The Role of Massive Stars in Young Starburst Galaxies

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Abstract

Starburst galaxies are defined as those galaxies undergoing violent star formation over relatively short periods of time (10 to 100 Myr). These objects may form stellar populations of greater than $10^6 \, M_\odot$, containing massive stars with masses $> 100 \, M_\odot$. Although most starburst galaxies are observed at relatively low redshift, recent evidence suggests that these types of galaxies were far more important in the high redshift past. It is believed that the chemical evolution of the Universe has been strongly influenced by this mode of star formation through the dense winds from massive stars and supernovae ejecta. Our understanding of starbursts is still relatively poor, since most are too distant to be resolved. We can gain some understanding of starbursts indirectly through the modelling of associated nebulae via the calculation of theoretical spectral energy distributions (SEDs) and photoionization modelling. This technique heavily relies upon the accuracy of the predicted far UV continuum of the massive star population. This thesis presents a new grid of SEDs for O stars, early B supergiants and Wolf-Rayet stars which have been incorporated into the evolutionary synthesis code Starburst99 (Leitherer et al. 1999). A total of 285 expanding, non-LTE, line-blanketed model atmospheres have been calculated to replace old, inaccurate LTE models for O stars, and pure helium, unblanketed models for W-R stars. These new grids cover five metallicities and the wind parameters are scaled with metallicity. We find that the new models yield significantly less ionizing flux below the He$^0$ ionizing edge at early phases and as a consequence, nebular He$^\text{II}$ $\lambda$4686 will not be observable in young starbursts. We use the photoionization code CLOUDY to test the accuracy of the predicted ionizing fluxes from our new models. We find that they are in much better agreement with observed optical and IR nebular line diagnostics than any previous models. The new W-R atmospheres are used in conjunction with 40 new O supergiant CMFGEN atmospheres to generate optical synthetic spectra of a starburst in its W-R phase. We demonstrate the use of this new spectral synthesis tool by modelling the observed spectra of five WR galaxies. We show for the first time that it is possible to derive consistent ages directly from the W-R stellar features and indirectly via the ionizing fluxes from the nebular line ratios.
To the banes and joys of my life
Tottenham Hotspur and Lowri
(Not necessarily in that order)
Acknowledgements

The work presented here would not have been possible without the help of many people. I would like to thank my supervisors Linda Smith and Paul Crowther for teaching me the skills needed to study starbursts and for pushing me through my Ph.D. My work colleagues in A5 must all be thanked for the comradeship and contributions to my work. These include my present office mates Jay Abbott and the Two Sams. I would also like to thank Allan Willis for the financial support during the final months of my writing up period.

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Finally I would like to thank my parents for their support over the seven years of my higher education, and my girlfriend Lowri who has put up with my fleeting mind. I would also like to thank Glenn Hoddle for steering Tottenham to a status of better than average Premiership football team.
Great spirits have always found violent opposition from mediocrities. The latter cannot understand it when a man does not thoughtlessly submit to hereditary prejudices but honestly and courageously uses his intelligence.

- Albert Einstein
- (1879 – 1955).
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Chapter 1

Introduction

1.1 Starbursts and Star Formation in the Universe

The chemical composition of our Universe is continually evolving due to the effects of star formation. Star formation exists in two different modes: the quiescent creation of stars in a host galaxy over long periods of time, and the violent gravitational collapse of starbursts where millions of stars form over very short time scales. A large number of galaxies at high redshift have been found to be undergoing intense star formation (Steidel et al. 1996; Lowenthal 1997) which suggests that starbursts were dominant in the early phases of the Universe. These starbursts still occur in the vicinity of our local Universe so can easily be studied in a variety of environments (e.g. Figure 1.1).

1.1.1 The Starburst Phenomenon

It has now been nearly thirty years since the discovery that some nearby galaxies contain regions of intense star formation. Early photometric and spectroscopic studies showed that some local galaxies have blue luminous regions which could only be attributed to hot young stars (Weedman 1973). The merger galaxy, NGC 7714 was the first object to be classified with the term “starburst” to describe its intense on-going star-forming activity (Weedman et al. 1980). A starburst may be formed when giant molecular hydrogen gas clouds of more than $10^6 M_\odot$ undergo gravitational collapse, triggered by some external mechanism (e.g. galaxy mergers), forming stars at rates of up to $50 M_\odot$ yr$^{-1}$ or greater (Izotova et al. 2000). Evidence now suggests that starbursts are an important part of our local Universe, where they contribute a very significant fraction of the total star formation. In fact about 25
Figure 1.1: HST image of the starbursting galaxy NGC 3310. This galaxy is undergoing intense star formation as shown by its blue colour and bright knots of young star clusters. This picture was taken by the WFPC2 instrument on the HST.
percent (Heckman 1998) of the massive star formation within a radius of 10 Mpc occurs in just four starburst galaxies. Massive stars, although far less common than their low mass counterparts, dominate the appearance of the starburst. The luminosity of a massive star (M > 5\(M_\odot\)) may be up to 10^6 times greater than that of a 1 \(M_\odot\) star. In fact it is these rare stars, always associated with the starburst phenomenon, which are responsible for the chemical enrichment of the surrounding gas and its host galaxy through powerful stellar winds and core collapse.

1.1.2 Quantifying Star Formation Rates

It is important to quantify the rate at which stars can be formed in a starburst. Determination of star formation rates (SFRs) in starburst galaxies can be measured via several different methods. Most of these use the fact that the most massive stars only exist in detectable concentrations in regions of recent star formation due to the bursts comparatively short lifetime, which is of the order \(\sim 10-100 \text{ Myr}\) (Coziol 1996).

At optical wavelengths, there are many narrow nebular lines that are only excited by the strong ionizing ultraviolet (UV) field radiated by O stars and their evolved descendants, the Wolf-Rayet (W-R) stars. Nebulae are always associated with populations of massive stars due to their high mass-loss rates and high UV output, coupled with the fact that these stars are so young that they have not had time to depart from their birth place. The most popular method for quantifying the SFR of a region is via the luminosity of the H\(\alpha\) \(\lambda 6563 \text{ Å}\) line (Thronson & Telesco 1986). This can be shown to be directly proportional to the SFR of the starburst region (Kennicutt 1998a).

\[
SFR_{H\alpha} = 7.9 \times 10^{-42} L(H\alpha) M_\odot \text{yr}^{-1}
\]  

(1.1)

This star formation rate determinant must be applied carefully, however, as the heavy dust obscuration found in some objects can lead to large uncertainties. Furthermore, some galaxies may not be detected at all at optical wavelengths because of high dust densities.

Another method for SFR quantification is to use longer wavelengths such as the far infra-red (FIR) luminosity of the parent galaxy (Hunter & Gallagher 1986). It is well known that dust converts UV and optical light to FIR wavelengths. Using the assumption that dust is converting the UV from massive stars to longer wavelengths it is possible to infer their number from the FIR luminosity of the parent galaxy. The SFR in the FIR
can be determined using the following equation (Izotova et al. 2000):

\[ SFR_{FIR} = 6.5 \times 10^{-10} \frac{L(FIR)}{L_B} M_\odot \text{yr}^{-1} \] \hspace{1cm} (1.2)

This is obtained using the assumption that all emission from OB stars is absorbed by dust and re-emitted into the FIR, and validated via the correlation found between \( L_B \) and \( L_{FIR} \) in the Second Byurakan survey of blue compact galaxies (Izotova et al. 2000). However, different methods for determining the SFR may not agree. It is also possible to infer the star formation rate from the 21cm luminosity using the following relation (Izotova et al. 2000):

\[ SFR_{21cm} = 2.5 \times 10^6 \frac{L(21\text{cm})}{L_B} M_\odot \text{yr}^{-1} \] \hspace{1cm} (1.3)

This equality uses the black-body radiation emitted from gas surrounding a starburst as a signature of star formation, using the assumption that ongoing star formation is heating the surrounding gas. A comparison of the FIR and 21cm star formation rates is presented in Table 1.1. This shows that the SFR rate derived from the 21cm luminosity can be up to ~

Table 1.1: Star formation rates in a sample of galaxies from (Izotova et al. 2000) derived from the FIR luminosity and 21cm luminosity.

<table>
<thead>
<tr>
<th>Name</th>
<th>( \log L_B )</th>
<th>( \log L_{FIR} )</th>
<th>( \log L_{21\text{cm}} )</th>
<th>SFR(<em>{FIR} ) ((M</em>\odot \text{yr}^{-1}))</th>
<th>SFR(<em>{21\text{cm}} ) ((M</em>\odot \text{yr}^{-1}))</th>
<th>SFR(<em>{21\text{cm}} ) (\text{SFR}</em>{FIR})</th>
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<td>1533+574A</td>
<td>42.05</td>
<td>42.71</td>
<td>28.01</td>
<td>0.85</td>
<td>6.59</td>
<td>7.76</td>
</tr>
<tr>
<td>1538+574</td>
<td>43.76</td>
<td>44.69</td>
<td>29.62</td>
<td>81.68</td>
<td>266.81</td>
<td>3.27</td>
</tr>
<tr>
<td>1556+583</td>
<td>44.00</td>
<td>44.08</td>
<td>29.03</td>
<td>19.82</td>
<td>62.52</td>
<td>3.49</td>
</tr>
<tr>
<td>1559+585</td>
<td>42.24</td>
<td>43.25</td>
<td>28.07</td>
<td>2.98</td>
<td>7.57</td>
<td>2.54</td>
</tr>
<tr>
<td>1629+205</td>
<td>43.45</td>
<td>43.76</td>
<td>28.69</td>
<td>9.59</td>
<td>31.32</td>
<td>3.27</td>
</tr>
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</table>

8 times larger than the FIR value, although on average they tend to be ~ 3.5 times larger. (Izotova et al. 2000) discuss this difference and suggest that the FIR star formation rate assumes 100 percent re-emission of UV to FIR, which is unlikely. The 21cm luminosity assumes that the radiation originates from the thermal nebular gas around a star forming region. Supernovae, however, can add significantly to the emission. These rates should be regarded as upper and lower limits to the star forming rate of a region.
Figure 1.2: The evolution of luminosity density as a function of redshift from Madau et al. (1998). Theoretical models plotted represent wavelengths of 0.1 (dotted line), 0.28 (solid line), 0.44 (short-dashed line), 1.0 (long-dashed line) and 2.2 (dot-dashed line) μm. All models assume a Salpeter IMF exponent, SMC-type dust and a universal \( E(B - V) = 0.1 \). The references for the data points can be found in Fig. 3. of Madau et al. (1998).
Figure 1.3: Redshift versus SFR density (SFRD) from Chapman et al. (2001). Open triangles are optical observations with and without dust correction, filled squares and hexagons and open squares are from sub mm sources, the open circles are radio sources with confirmed optical counterparts and the open stars are the sum of the optically selected (uncorrected for dust) and submillimeter data points. All references are cited in Chapman et al. (2001).
1.1.3 The Star Formation History of the Universe

Recent studies of high redshift galaxies have shown that star formation in the past far exceeds that which can be measured in the local Universe. Studies of the Hubble Deep Field (HDF) taken at optical wavelengths, led to the creation of the Madau plot shown in Figure 1.2, primarily attributed to Madau et al. (1996, 1998) who showed that the star formation rate peaks at around a redshift of $z \sim 1.5$, (or 40 percent of the current age of the Universe). Madau used the fact that UV light from massive stars ionises surrounding hydrogen gas to create nebular emission lines such as Ly$\alpha$ or H$\alpha$ which are only found in areas of recent star formation activity. However, others have recently realised that at high redshifts, dust obscuration can lead to significant reddening, and so may bias results to lower redshift. Other methods for determining SFRs have been developed to avoid this possible selection effect.

For example, recent work has focused on using infrared and submillimeter observations to search for re-emission of the absorbed UV light by the surrounding dust clouds associated with star formation. The absorbed light from the massive star population is re-radiated in the far-infrared at rest frame wavelengths of 60 - 100$\mu$m. This means that this emission at high redshifts is converted to the submillimeter domain of 175 - 1850 $\mu$m. Recently submillimetre (Blain & Natarajan 2000) measurements from the Submillimetre Common User Bolometer Array (SCUBA) and far-infrared measurements of the background extragalactic counts have been independently used to determine the star formation activity as a function of red shift. These alternative studies suggest that dust may obscure starbursting regions at higher redshift and star formation can be detectable via longer wavelength studies such as submillimetre which look at the dust re-emission of hot star fluxes. To this end a submillimetre Madau plot has been constructed. The SCUBA measurements suggest that the SFR is far higher and reaches back to a much earlier epoch of the Universe. Fig. 1.3 which shows the Madau plot (Chapman et al. 2001), including selected submillimetre and radio data, suggests a flattening of the star formation rate perhaps peaking beyond about $z \sim 3$, or 10 percent of the age of the Universe. This implies that star formation is more important to the early evolution of the Universe than previously thought.
Figure 1.4: HST image of the merging galaxies NGC 4038 and NGC 4039, named the Antennae. This picture was taken by the WFPC2 instrument on the HST.
1.1.4 Star Formation Triggering Mechanisms

The triggering mechanisms of starbursts are still a subject of hot debate, although it is thought that many are triggered by processes such as galactic interactions and mergers (Schweizer 1987; Jog & Das 1992), instabilities in the bar structures in disk galaxies (Shlosman et al. 1990) and shockwaves from supernovae and stellar winds (Heckman et al. 1990). However, these mechanisms have been studied in mature disk galaxies in the local Universe and may not be indicative of the higher rates of star formation in the past. Studies of dwarf irregular Blue Compact Galaxies (BCG), situated at redshift $z \approx 0.5$ and beyond, have led to more general theories on the triggering of starburst regions. BCGs are thought to be the local analogues of the star-forming galaxies common at high redshift in the Hubble Deep Field (Bergvall & Östlin 2002; Ellis 1997).

These galaxies are all metal poor and have high levels of star formation, and their gas consumption time scales are significantly shorter than the Hubble time. They therefore could not possibly form stars at high rates for long periods (Fanelli et al. 1988). At first it was postulated that these galaxies were in the process of forming their first stars, however this was later rejected when evidence for an old stellar population was detected in their haloes via photometry. A more sensible conclusion was that these galaxies underwent periods of intense starbursts followed by quieter periods of quiescent star formation (Searle et al. 1973).

One explanation for this type of behavior could be the ejection of hot gas into the intergalactic medium by supernovae (SNe) and dense stellar winds which prevent further star formation. This gas cools and, depending on primarily the mass of the galaxy, accrete back onto its host to trigger a new round of starbursts (Babul & Rees 1992). The number of active periods would still have to be surprisingly few to account for the low metallicity of these objects, as starburst super-winds would enrich the interstellar medium (ISM) quickly, therefore mechanisms must exist to inhibit the collapse of the ejected gas.

An alternative view is one with many similarities to the present day Universe. Tidal interactions due to the presence of a companion or a merger could be responsible. Indeed in our local Universe many examples of merging pairs of galaxies undergoing starburst phases can be found (c.f. Fig. 1.4, the Antennae galaxies). Although the HDF suggests that mergers were common in the past, other evidence by Compos-Aguilar et al. (1993) suggests that few BCGs have any companions within several hundred kpc. Supporting this
finding is work by Telles & Terlevich (1995) and similar conclusions were drawn by surveys looking at galaxies with narrow line emission and UV excess. However, work by Taylor et al. (1995, 1996) which surveyed H I companions to H II galaxies, solves this inconsistency and suggests that in fact 60 percent of BCGs have close companions in the form of faint inactive gas rich galaxies which are undetectable at optical wavelengths. This evidence tends to suggest that tidal interactions may be the main trigger for starburst galaxies.

1.2 The IMF and Stellar Populations

When a giant H I cloud collapses, a range of stars of different masses is always formed. In general, stars of masses ranging from 0.1\(M_\odot\) to as much as 120\(M_\odot\) are formed in these regions. It is now over 45 years since Salpeter published work describing the mass number density of the local solar neighbourhood population of stars (Salpeter 1955). This parameterization now termed the Initial Mass Function (IMF) is given by:

\[
N_{tot} = k \int_{M_{low}}^{M_{up}} m^{-\alpha} dm
\]

where \(N_{tot}\) is the total number of stars in a system, \(M_{up}\) is the mass of the most massive component, \(M_{low}\) is the mass of least massive component, \(\alpha\) is the power law exponent and \(k\) is a normalisation constant. Today the IMF is a fundamental parameter used in many areas of astrophysics, from star formation in starbursts to the chemical and dynamical evolution of the Universe.

Although the concept of the IMF has existed for nearly half a century, the values of its power law slope and the upper and lower mass limits of the stellar population are still extremely uncertain. It is still not known whether an IMF can be applied universally for all modes of star formation, or if stars form with different IMFs depending on the stellar environment.

1.2.1 The Importance of the IMF

Not only is an accurate determination of the IMF of a system important in the modelling of starbursts, but the IMF’s importance is much more far reaching. Star formation rates (see §1.1.2), are primarily determined via the radiative properties of massive stars above 5-10\(M_\odot\), which output most of their energy in the far UV region of the spectrum. Since
these stars make up only about 5-20 percent (Kennicutt 1998b) of the mass of the active star bursting region, estimates of the total rate depend very heavily on the IMF, where a modest change in the exponent or mass limits can lead to a large difference in the SFR. The most dramatic change occurs when the Hα line is used as the SFR diagnostic. Only extremely massive stars (M > 10 M☉) are capable of exciting the Hα emission in a surrounding H II region, making the SFR calculation more reliant on the form of the IMF. This could have wide reaching implications for observed SFRs if the IMF changes from galaxy to galaxy, as the varying IMF will be picked up as changes in the SFR by up to an order of magnitude. For example, changing the power law exponent from α = 2.35 to α = 3.5 can lower the star formation rate by a factor of 30 (calculated using a FORTRAN code based on Equ. 1.4, with M_up =100M☉ and M_low =1M☉).

Observations of extragalactic environments can often lead to other inferred properties of the host galaxy. As these galaxies are in general too far away to resolve their faint lower mass stars, the IMF has a wide reaching effect on our understanding of them. Predictions of the chemical yields of a star bursting region are heavily influenced by the IMF, as is the mass to light ratio of a galaxy and therefore the dark matter content within the galaxy, since lower mass stars are inferred by an IMF. Another extremely important area which is reliant on a good determination of the IMF is the formation and early evolution of galaxies. The IMF has important effects on the amount of feedback due to energy release (non-radiative and radiative), which inhibits gas from re-condensing and forming stars in a runaway process.

1.2.2 Determination of the IMF

Observationally it is possible to determine the IMF for different environments. The first attempt at this determination was for individual resolved stars in the solar neighbourhood (Salpeter 1955). These stars were simply classified photometrically and binned to create an empirical power law in the form of Eqn 1.4 in the mass range of 0.3 to 10M☉ with a slope of α = 2.35. The value of α from this work is still adopted for the power law exponent of the IMF today, although the origin of the IMF slope is still very poorly understood. More recently, Scalo (1986) discussed the form of the IMF over an extensive mass range. This work again considered field star formation, and cluster photometry, coupled with new luminosity functions (LF) and mass-luminosity relationships. The cluster work yielded a variation in the power law slope at about 10M☉, with stars lower than this mass adhering
to power law of 2.8, and more massive stars obeying an $\alpha = 2.2$ law, more consistent with a Salpeter IMF. It was also concluded that local field star surveys are too inaccurate to reveal a variation in the power law slope. This is due to problems with using stellar evolutionary tracks, as unlike clusters, field stars are not created simultaneously, but rather continuously; therefore the age and original mass of individual stars can be indeterminate.

Recent work has focussed on smaller mass ranges (e.g. Kroupa 2001), as evidence suggests that the IMF slope is not universal over the entire range of possible stellar masses. These can be grouped into three main areas.

**The Low Mass IMF**

Several works have tried to evaluate the IMF in the mass range $0.1M_\odot < m < 1.2M_\odot$. Reid & Gizis (1997) conducted an 8 parsec study of the solar area, in the range $0.1M_\odot < m < 1.0M_\odot$ and found that generally the IMF was flat in this region ($\alpha = 0$). However, Reid & Gizis (1997) found some evidence for a decline in numbers of stars lower than $0.1M_\odot$ ($\alpha < 0$). Scalo (1998) pointed out that most field stars in this mass range are too young to be on the Main Sequence (MS). These Pre–Main sequence stars (PMS) are more luminous than their MS counterparts, so corrupting the mass–luminosity relation. Scalo (1998) concludes that no work at present has derived a true IMF below the $0.1M_\odot$ limit. Other recent work by Mèra *et al.* (1996), Kroupa (1995a; 1995b) and others put the power law exponent for the low mass range at 0 to -0.5. Because of the large number of nearby low mass stars with measured parallax, this mass range is thought to be the most well known in the IMF in spite of the difficulties of the poorly defined mass luminosity relationship, over-luminous PMS stars and contamination due to binary systems.

**The Intermediate Mass IMF**

The intermediate mass range deals with stars from 1 to $3M_\odot$. There are many uncertainties that need to be dealt with in order to obtain a realistic value of $\alpha$. For masses greater than $0.85M_\odot$, the lifetimes of the stars become short enough that the earliest field stars will have evolved off the MS, so evolutionary effects must be considered. Another source of contamination is that star formation in the Milky Way is though to have occurred in bursts with peaks in the galactic SFR at 2-5 and 7-9 Gyr ago (Rocha-Pinto *et al.* 2000), creating a synthetic peaks in the field IMF for stars with intermediate and low masses (existing higher mass stars would have to have formed more recently at a lower rate).
There are several works that have considered the intermediate range, avoiding most of these problems by considering only clusters and associations (e.g. Grebel & Chu 2000). Clusters with ages greater than 10 Myr have no PMS stars in this range, and are relatively uncontaminated by binaries (which alter the luminosity of a system and hence its mass) and most importantly there is no uncertainty due to the SFR of the galaxy at earlier epochs. The problems that plague clusters, such as field star contamination and radial segregation are considered small in comparison. Many studies of the IMF of clusters have been undertaken resulting in a range of IMFs resulting in an average of $\alpha = 2.3$ (Kroupa 2001).

**The High Mass IMF**

The massive star IMF is extremely difficult to define for several reasons. The nature of the IMF means that the number of these stars is extremely small compared to the other mass ranges, so errors will be high. Massive stars also evolve very quickly, for instance, a $120M_\odot$ star will only spend 2.6 million years on the MS if non-rotating (Meynet & Maeder 2000), so only very young regions will contain these most massive stars. During this MS lifetime, the luminosity of the star may also change by a factor of 10 due to its evolution from dwarf to supergiant. An older association may have a mixture of massive evolved stars combined with those still on the MS, due to variations in the lifetimes of these stars as a function of mass. This makes photometry very difficult as the most massive stars may be cooler post-main sequence supergiants rather than very blue MS stars. More importantly, O and early B stars are photometrically degenerate, so stars of $100M_\odot$ and $10M_\odot$ are indistinguishable. The only reliable way to determine the IMF at these masses is to use spectroscopy, which is far more work intensive than the use of photometry. Spectroscopic measurements of stars in OB associations in the Milky Way, and the Magellanic Clouds have been the principle way to determine the massive star IMF. Oey & Massey (1995) and Oey (1996) are two examples of some of the surveys undertaken to determine the stellar mass content of some of these associations. These works considered 10 LMC, 1 SMC and 13 Milky Way OB associations with ages between 1 to 11 Myr with the most massive star measured being in excess of $120M_\odot$. This study like more recent ones (Elmegreen 1997) concluded that the IMF for very massive stars has a power law similar to that of the original Salpeter IMF of $\alpha = 2.35$.

The upper mass limit of the IMF is uncertain, although very few stars are found with
masses in excess of $120M_\odot$. There are several candidates for the most massive star known including the doomed star $\eta$ Carina, whose mass is still hotly debated (100 - 250$M_\odot$) and the Pistol star, estimated to have had an initial mass of $200M_\odot$ (Figer et al. 1998).

**Current Impressions of the IMF**

Measurements of the IMF in all mass ranges are thought to be uncertain enough not to be able to answer the most important question in this field of research; is the IMF universal? Theoreticians would certainly like this to be the case, although for the moment the question remains unanswerable. Certainly there is strong evidence for several breaks in the IMF power law slope, segregating the IMF into 3 (or possibly more) parts. However several points can be made about the IMF.

1. Low mass cutoff is probably < $0.1M_\odot$, corresponding to the H-burning limit of $0.08M_\odot$.
2. $\alpha = 1.2 \pm 0.3$ for $0.1 < M < 1.0M_\odot$
3. $\alpha = 2.3 \pm 0.5$ for $1 < M < 10M_\odot$
4. $\alpha = 2.3 \pm 0.5$ for $10 < M < 100M_\odot$
5. The upper mass cutoff is probably greater than $120M_\odot$, although the observed value does change from region to region.
6. The high mass stars are long gone before the low mass stars even reach the MS.

Kroupa (2001) includes one important caveat in his discussion of the IMF, giving a form similar to the one mentioned above. Binary systems are not usually included in the IMF determination, which will bias the power law exponent to a shallower (lower) value. The actual slope may be steeper by 0.5 and new observations may be able to resolve these systems.

### 1.3 The Stellar Population of Starbursts

Starbursts are always dominated by the properties of the most massive components of the stellar population. Although these stars, as discussed above are far fewer than their low mass counterparts, they can be up to $10^6$ times more luminous. These massive stars
evolve so quickly that within 10 million years from their initial collapse they will die, long before the lower ($1M_\odot$) mass stars even reach the MS.

### 1.3.1 Massive Star Evolution

During the time it takes a massive star to burn its fuel, the star’s appearance and processes within it can change dramatically. The rate of change and the stages a star may go through depend greatly on the initial mass of a star. In a defined chemical environment such as that of our own solar neighbourhood, we can use evolutionary tracks such as those of Meynet et al. (1994) to calculate the evolutionary sequence of a massive star. Below is the inferred evolutionary sequence of a solar metallicity massive star as a function of mass from the Meynet et al. (1994) tracks.

- **120 $M_\odot$**: \(O V \rightarrow BSG \rightarrow WNL \rightarrow WNE \rightarrow WCE \rightarrow WO\)
- **85 $M_\odot$**: \(O V \rightarrow BSG \rightarrow WNL \rightarrow WNE \rightarrow WCL \rightarrow WCE \rightarrow WO\)
- **60 $M_\odot$**: \(O V \rightarrow BSG \rightarrow LBV \rightarrow WNL \rightarrow WNE \rightarrow WCL \rightarrow WCE \rightarrow WO\)
- **40 $M_\odot$**: \(O V \rightarrow BSG \rightarrow RSG \rightarrow WNL \rightarrow WCE \rightarrow WO\)
- **25 $M_\odot$**: \(O V \rightarrow BSG \rightarrow RSG\)

Where $O V$ signifies the young MS O dwarf stars, BSG is Blue Supergiant, RSG is Red Supergiant, LBV is Luminous Blue Variable and WNL, WNE, WCL, WCE and WO signifies the various Wolf-Rayet star stages. We can see that all massive stars go through a BSG phase as they become more luminous. At solar metallicity, only stars as massive as 40 $M_\odot$ and greater become W-R stars, although through different routes. A 40 $M_\odot$ star will cool down to become a RSG before it becomes a W-R star whereas a 120 $M_\odot$ star will move straight into the W-R phase. Chapters 2 and 4 comment on the O supergiant (BSG) and W-R phases of massive star evolution.

### 1.3.2 The HRD

Perhaps the most famous diagram in all stellar physics is the Hertzsprung–Russell Diagram, or HRD. In 1912, Einar Hertzsprung and Henry Norris Russell, working indepen-
dently discovered that in most stars surface temperature and luminosity could be related. This was the discovery of the MS, and now refers to a star which is core hydrogen burning. Figure 1.5 shows the MS on the upper part of the HRD, concerning the hot luminous

![Figure 1.5: The upper Hertzsprung-Russell diagram, showing MS, core hydrogen burning stars, the luminous Blue Supergiants (BSG) and Luminous Blue Variables (LBV), cool luminous Red Supergiants (RSG) and the evolved Wolf-Rayet stars. The plot is truncated at the beginning of core helium burning.](image)

massive star population. This figure plots the synthetic population produced by the evolutionary synthesis code Starburst99 (Leitherer et al. 1999) from 1 to 10 Myr. This figure shows that after 1Myr, all massive stars lie on the MS, where they burn hydrogen in a relatively steady state. Massive stars reach the zero age main sequence (ZAMS) \( \sim 1 \) Myr after their birth. Lower mass stars may not reach the MS, for instance, a \( 3M_\odot \) star will only reach the MS 20Myr after its birth (Palla & Stahler 1999). This is shown in the HRD as an increase in luminosity. Further on in its evolution, a massive star will cool as
it becomes larger moving to the Blue Supergiant region, or in lower mass cases (< 40M\(_\odot\)) into the RSG region, where a massive star could be classified as a G or even F or M type star. At the other end of the temperature range, extremely massive stars enter the W-R phase without becoming red supergiants.

The HR diagram is a good test of theoretical results and evolutionary tracks are often plotted over the temperature-luminosity determinations of cluster members to test their validity. One useful test of modern evolutionary codes is the prediction of numbers of red and blue supergiants, or the numbers of W-R stars in young clusters. In later chapters we will present the Meynet et al. (1994) evolutionary tracks and discuss their predictions when applied to synthesis models.
Chapter 2

Stellar Atmospheres

For over one hundred years the atmospheres of stars have been modelled in an attempt to understand them. In 1900 Max Planck announced his formula of the description of the spectrum emitted from a black body by combining the formulae of Wien & Rayleigh. Although a distinct temperature – colour relationship could be described, Planck realised that classical physics could no longer fully describe the spectra of stars. In 1913 Niels Bohr correctly predicted the spectral positions of the hydrogen lines in stellar spectra using his model of the atom. The theory of quantum mechanics was born. Today, modelling stellar atmospheres is still extremely important in understanding the physics of these extreme environments. From this understanding, not only is it possible to determine the chemical composition and temperature directly via measurements of emission and absorption lines, but using modelling techniques, it is also possible to discover indirectly the mass, age and stellar wind properties of a star. There are many useful observable properties that can be modelled in a star. Here, we concentrate on the modelling of stellar atmospheres which predict the Spectral Energy Distribution (SED) emitted from a star. This technique is crucial as it provides the fingerprint of the star, since its spectrum originates in its atmosphere, where light is absorbed and re-emitted by the atomic species that are present within it. Classically a stellar atmosphere comprises a thin layer of hot gas residing between the surface of the star and the interstellar medium. More recently it has been shown that the presence of a corona and a stellar wind also have a significant effect on the appearance of the star, and are now considered as part of the atmosphere. As discussed below, the chemistry, temperature, surface gravity and pressure in the atmosphere all affect the energy and radiation balances which produce the stellar SED. All models discussed
here use complex physical methods to describe processes in the stellar atmosphere. One of the most important processes is that of line blanketing. This important process is the absorption of radiation due to the spectral lines of an atom, which is re-emitted at longer wavelengths. This process is discussed at length in this chapter.

In the case of a massive star, there are sufficient high energy photons to excite ions in an associated gas cloud into higher level energy states and create the characteristic emission lines of a nebula. These nebular lines can be detected by modern telescopes with reasonable signal to noise even outside our Local Group and can be used to indirectly probe the far-UV stellar continuum. This relies upon the accuracy of evolutionary predictions and more importantly the accuracy of the shape of the continuum predicted by an atmospheric model, which deviates from the Planck black-body function due to the action of atomic physics. In §2.1, the structure of a basic Local Thermodynamic Equilibrium (LTE) model is discussed. From this model the more complicated assumptions of a non-LTE model are given, concerning the more realistic approach used to describe the quantum states of the atomic species in the stellar atmosphere. The two models are compared and the failings of the LTE case are discussed. §2.2 introduces the various competing non-LTE models that are currently available for computing a realistic grid of hot stellar atmospheres, briefly describing the workings and assumptions of the most popular models (WM-basic, TLUSTY, CMFGEN and CoStar). §2.3 describes the choice of parameters for the new O and early B star atmosphere grid. The properties calculated in the grid are presented for the 165 models including the ionizing properties of the atmospheres beyond the H\(^0\), He\(^0\) and He\(^+\) continuum limits. These models are also compared with the existing CoStar atmosphere grid of Schaerer & Vacca (1998) calculated with the use of isa-wind (de Koter et al. 1997) and show that the new continua are softer in the ionizing continuum due to a more comprehensive calculation of line blanketing. It is shown that the WM-basic grid agrees more closely with observations. §2.4 presents the new Wolf-Rayet (W-R) grid with its choice of parameters, justifying the reasons for using fixed luminosity and the level of mass-loss in these wind dominated objects. These new models are compared to the Schmutz et al. (1992), pure helium, unblanketed grid, showing extreme differences in their ionizing fluxes especially at lower temperatures.
CHAPTER 2. STELLAR ATMOSPHERES

2.1 Model Atmospheres

In order to describe a fully working complex non-LTE model, it is simpler to describe a basic LTE case and replace the underlying assumptions with that of the more complex non-LTE physics. The basic physics of radiatively driven wind theory is then discussed, introducing new thoughts on the metallicity scaling of outflows from massive stars.

2.1.1 A Simple LTE Atmosphere

A modern model stellar atmosphere can be thought of as the coupling of the macroscopic conditions of a gas, such as temperature, density and velocity field with the microscopic quantum processes of each individual atomic component. In order to describe this plasma, one must not only describe its large scale structure through hydrostatic/hydrodynamic equations and radiative equilibrium using a suitable radiative transfer equation, but one must also define all the quantum states of all of the particles that make up the stellar atmosphere. Initially, a model geometry must be assumed. In Main Sequence (MS) stars, the emergent flux is mainly from the photosphere, which can be shown to be small fraction of the star's radius (1/1000 R*, Philips 1999). This makes a plane-parallel geometry favourable as it can easily be approximated to a one-dimensional hydrostatic case which can be modelled as a series of slabs of atmosphere each at sequential distances from the stellar surface as described in Figure 2.1. However, some atmospheres are known to be extended, due to strong stellar winds, present in evolved stars such as supergiants or W-R stars. In these cases, more complex spherically expanding hydrodynamical models must be used to take account of the large velocity gradient experienced between the inner atmosphere and the outer edge of the wind. Although this model geometry is more complex, it can model wind effects very well, and gives rise to structures known to be observed in emission line profiles. Once a geometry is assumed, equations of state can then be found to describe the plasma at different points from the stellar surface. These equations can be solved to find the pressure at any point in the atmosphere.

The pressure, $P$ is represented by the sum of $P_r$ and $P_p$ the radiation and particle pressures. In cooler stars the radiation pressure can be omitted due to its insignificant effect, but in hot stars this effect can be as large as 40 percent of the total pressure (Philips 1999). Hydrostatic equilibrium generally holds in the case of a stellar photosphere, but not for the wind, for which a net outward movement of matter must be considered in
Figure 2.1: A simple diagram to explain the workings of a plane-parallel model atmosphere.
a hydrodynamic case. This is most important in stars with very high mass-loss such as Luminous Blue Variables (LBVs), O supergiants (O If) and their evolved descendants, the W-R stars where winds can be opaque and completely hide the photosphere of the star.

In tandem with the pressure, the temperature structure of the atmosphere must be defined. The simplest case, and therefore the best starting point, is to use a grey case temperature solution, which reduces to an analytical expression. The opacity of a system is extremely important to the model. It represents the atmosphere’s ability to absorb radiation at a point in the atmosphere. A simple description uses electron scattering and Kramers’ law i.e. \( \kappa \propto \rho T^{-\frac{3}{2}} \) for free-free and bound-free interactions (these are assumed dominant), where \( \kappa \) is the opacity, \( \rho \) is the density of the atmosphere and \( T \) is its temperature. This simple model is not good enough for a realistic representation of an atmosphere however, where the opacity is strongly coupled with the quantum state of the gas. Once the macroscopic state of the atmosphere is defined, the radiation transport within it can be described. Any such transport must satisfy the condition of the radiative equilibrium:

\[
\int_0^\infty (\chi_\nu J_\nu - \eta_\nu) d\nu = 0 \tag{2.1}
\]

where \( \eta_\nu \) is the emission and \( \chi_\nu \), the extinction, is defined as the sum of the opacity (or absorption) and scattering properties of the system. \( J_\nu \) is the mean intensity of the radiation. Equations 2.1 stipulates that the product of the extinction and incoming flux must equal the emissivity of the system integrated over all frequencies, i.e. the total energy of the system must be conserved. Although this may not be the case globally (especially obvious since the star emits light), it may be true locally, as approximated by LTE. This also means that the model neglects convection, present in cool stars, magnetic heating present in the solar corona, and any nuclear processes, all of which should be negligible in the atmosphere of a hot star. Equation 2.1 is then coupled with the general radiative transport equation below

\[
\left( \frac{1}{c} \frac{\partial}{\partial t} + n. \nabla \right) I(\nu, r, n, t) = \eta(\nu, r, n, t) - \chi(\nu, r, n, t) I(\nu, r, n, t) \tag{2.2}
\]

Equations 2.1, 2.2, 2.3 and 2.4 describing the radiation transport are taken from Hubeny (1999), although a similar discussion can be found in Mihalas (1978). \( I_\nu \) is the intensity over a unit time interval, per unit solid angle, per frequency range. \( \eta_\nu \) represents the radiation emitted thermally from the plasma taking into consideration the effect of thermal
excitation and spontaneous emission, and is defined by

$$
\eta_\nu = \frac{2\hbar \nu^2}{c^2} \left( \sum_i \sum_{j>i} n_j (g_i/g_j) \sigma_{ij}(\nu) \right) \\
+ \sum_i (n_i^* \sigma_{i\nu}(\nu) e^{-h\nu/kT}) \\
+ \sum_\kappa n_\kappa n_\kappa \sigma_{\kappa\kappa}(\nu,T) e^{-h\nu/kT}
$$

(2.3)

In thermal equilibrium, $\eta_\nu$ reduces to the Planck black body function. $\chi_\nu$ is the sum of absorption and emission coefficients and represents a realistic form of the opacity of a local region of atmosphere. The product of $I_\nu$ and $\chi_\nu$ therefore represents the absorption corrected intensity. $\chi_\nu$ is a complex function which represents absorption and emission from four types of quantum mechanical process, namely bound-bound, bound-free, free-free transitions and electron scattering. Bound-bound transitions arise from the excitation and de-excitation of electrons to different energy levels in an ion and are responsible for the creation of spectral lines. This could be in the form of either absorption of emission depending on the conditions driving the population levels of each ion. Bound-free transitions are due to ionisation or recombination, and create the underlying flux level of the continuum. Free-free transitions, accounting for Bremsstrahlung radiation and electron scattering are also usually included. $\chi_\nu$ is normally written as in equation 2.4

$$
\chi_\nu = \sum_i \sum_{j>i} [n_i - g_i/g_j n_j] \sigma_{ij}(\nu) \\
+ \sum_i (n_i - n_i^* e^{-h\nu/kT}) \sigma_{i\nu}(\nu) \\
+ \sum_\kappa n_\kappa n_\kappa \sigma_{\kappa\kappa}(\nu,T) (1 - e^{-h\nu/kT}) + n_e \sigma_e
$$

(2.4)

$n_i$ represents the population density of state $i$, with $g_i$ its corresponding statistical weight. The $n^*$ represents the population for statistical equilibrium (LTE). $\sigma$ represents the cross section of the atom at frequency $\nu$ with $\kappa$ denoting the continuum. The thermal emission of radiation $\eta_\nu$ has components defined in a similar way. Both $\eta_\nu$ and $\chi_\nu$ must take into consideration the range of velocities of particles in the local rest frame. This velocity distribution will give rise to Doppler shift effects and broaden the absorption/emission properties of the gas. In LTE a Maxwellian distribution is assumed. This distribution gives rise to the underlying shape of the line profiles (although in more complicated models
other effects can have a large effect.). In addition, to these macroscopic equations (albeit with microscopic components), atmospheric models need a microscopic set of equations to determine the statistical distribution of the atomic level populations.

These atmospheres require a description of the state of not only all atoms, but all atomic species at all points in the system. The calculation of the state of every ionic species at every point in the atmosphere cannot be performed exactly, as every state is coupled with every other. This method would be extremely computationally intensive, with maybe one million energy levels being required for a reliable solution. One way of reducing this computing time is to assume a LTE solution. Although just as many states have to be dealt with, the statistical distribution of level populations can then be calculated from the Saha-Boltzmann equation due to the equilibrium *locally*. The Saha-Boltzmann equation describes the number of ions in an associated quantum state. This equation is simply the product of the Saha equation and the Boltzmann equation. The Saha equation describes the relative fraction of an ionic species in an excitation state \( r + 1 \) relative to the number of ions in the lower state \( r \). The Boltzmann excitation equation represents the number of ions, \( N_r \), that are in energy state \( s \). Once the absolute distribution of the energy levels for each ion is known, it is then possible to calculate \( \chi_\nu \) and \( \eta_\nu \). Now we have all the physics to represent a component of a stellar atmosphere in LTE, we can calculate the way a simple slab reacts to input radiation from the physics above. Only one assumption remains, the radiation at the inner boundary must be specified. Normally it is assumed that a simple grey body is good enough to represent the output from the stellar core.

A Planck function may be used in a simple model, but for a non-LTE model an LTE model would be used as an input to help model convergence. The physics used in the model described above are relatively simple and therefore can be calculated linearly. In any real atmosphere model (LTE or non-LTE) this is not the case, since temperature depends on opacity and opacity depends on temperature through collisional rates. This means that the macroscopic and microscopic physics are dependent and should be worked out simultaneously. As this is impossible, an iterative solution is required.

2.1.2 Radiatively Driven Winds

For three quarters of a century it has been realised that radiation could be the mechanism for driving dense winds from massive stars, due to the interaction between radiation and matter. In 1926, Milne (1926) proposed a mechanism in which radiation pressure applied
force to ions in the atmosphere of stars, pushing them outwards to create a wind. This paper also mentions the possible existence of a terminal velocity, a valuable quantity which we now term \( V_\infty \). It was then over 40 years later that observations of three stars in Orion revived interest in mass-loss mechanism of massive stars. In the autumn of 1965, the Aerobee rocket obtained far-UV spectra for \( \delta, \epsilon \) and \( \zeta \) Orionis, indicating wind velocities of 2000 km s\(^{-1}\) and mass-loss of the order of \( 10^{-4} M_\odot \) yr\(^{-1}\) (Morton 1967). In 1970 Lucy & Solomon (1970) found that their model of massive star atmospheres showed the ions in the atmosphere could have strong negative effective gravities because of the radiation pressure from the star. Further development was reached with the famous paper of Castor et al. (1975), which developed a formal theory of the radiation pressure on absorption lines as the mechanism behind the driving of wind outflows of massive stars. Today far more is known about this phenomenon, due to additions to these pioneering works.

Unlike a low mass star, which may have a transparent wind driven by a magnetic field, a massive star has an almost opaque wind (or completely opaque in the case of a W-R star). Since O stars emit strongly in the UV, where many atomic species can have thousands of lines, the wind is line driven, rather than through continuum scattering processes. In fact the opacity due to the C\( \text{IV} \) 1550 Å line can reach \( 10^6 \) times that of the electron scattering value (Lamers & Cassinelli 1999). A simply way to describe the effect of absorption on the momentum of the wind is to look at the effect of one line in the wind and generalise as in Lamers & Cassinelli (1999).

**A Single Line Model**

if we consider an optically thick wind, then all photons with frequency \( \nu_0 \) (the frequency which corresponds to the transition energy of the line) will be absorbed at the photosphere where material is not moving. As all the momentum of the photon is transferred to the atmosphere, we assume that the wind has a velocity field associated with it from \( v = 0 \) at the photosphere, to \( v = v_\infty \) as \( r \to \infty \) (the terminal velocity). This means that, due to the Doppler effect, the frequency that is absorbed at the terminal velocity is \( \nu = \nu_0 v_\infty / c \).

The total range of frequencies absorbed by the wind is then \( \nu = \nu_0 (1 + v_\infty / c) \).

The rate at which momentum is carried away by the wind is:

\[
M v_\infty = \frac{1}{c} \int_{\nu_0}^{\nu_0 (1 + v_\infty / c)} 4\pi R^2 \mathcal{F}_\nu \, d\nu \simeq \frac{4\pi R^2}{c} \mathcal{F}_{\nu_0} \nu_0 - \frac{v_\infty}{c} \quad \text{(2.5)}
\]

Since \( v_\infty \) is on both sides of the equation, the momentum transfer is independent of the
terminal velocity. Also, if we use a Planck function to evaluate this quantity, then we find:

$$\dot{M} \simeq \frac{L}{c^2}$$

However, if the wind has many optically thick lines this can be generalised to

$$\dot{M} \simeq N_{\text{opt}}\frac{L}{c^2}$$

Here we can see that the momentum transfer is proportional to the number of lines that can interact with the wind. In the most extreme case, where many lines are present, the momentum transfer is optimised where lines absorb all the light from the star and are scattered in the wind. Generally, the wind is not one hundred percent efficient at absorbing photon momentum into the wind. The wind efficiency can be quantified as:

$$\eta \equiv \frac{\dot{M}v_\infty}{\frac{L}{c}}$$

This value lies between 0 and 1, with 1 being 100 percent efficient. This is generally not the case since the wind may not be optically thick at every point, as the opacity of the wind changes with radius, and lines will not cover the entire range of the spectrum, and are usually most efficient in the UV.

A Realistic Model of Line Driven Winds

A much more complete description of the wind is derived from the work of Castor et al. (1975) (CAK) and later additions by Kudritzki et al. (1989). Here the total radiative acceleration is assumed to be the product of electron scattering and a force multiplier, $M(t)$ defined as

$$M(t) = k t^{-\alpha} \left(10^{-11} n_e/W\right)^{\delta}$$

where $k, \alpha$ and $\delta$ are called the force multiplier parameters. $k$ is a measure of the number of lines acting on the wind, $\alpha$ is a measure of the number of optically thick to thin lines and $W$ is a geometrical dilution factor needed to correct for the finite size of the stellar disk. Finally $\delta$ quantifies the degree of ionization in the wind. Using this approximation it is possible to derive an equation of motion for the wind. From this, the formal solution
for the mass–loss rate is:

\[ \dot{M} = \left( \frac{\sigma_e v_{th}}{4\pi} \right) \left( \frac{\sigma_e}{4\pi} \right)^{\frac{1}{\alpha}} \left( \frac{1 - \alpha}{\alpha} \right)^{\frac{1-a}{\alpha}} (\alpha K)^{\frac{1}{\alpha}} \left( \frac{L_*}{c} \right)^{\frac{1}{\alpha}} \left( GM_*(1 - \Gamma_e) \right)^{\frac{a-1}{\alpha}} \]

(2.10)

where \( K \) is equal to \( (n_e/W)^{\delta} \) and \( M_*(1 - \Gamma_e) \) is the effective mass. \( \Gamma_e \) is the radiative force due to electron scattering. The solution is presented in Vink (2000), adapted from the analytic solution given in Kudritzki et al. (1989). This solution, however, is only valid when the Castor et al. (1975) force multiplier parameters are constant. This is certainly not the case for a modern wind model, where all three values change in the wind due to the varying ionization balance at different points in the wind. The dependence on the solution of the number and nature of the lines is extremely high, so as a consequence the metallicity of a star is extremely important to the rate of mass–loss. One might well expect the mass–loss of a star to be dependent on metallicity, since certainly the number of heavy element lines will diminish with metallicity, affecting both the balance of optically thick to thin lines (\( \alpha \)) and the number of lines (since Fe has the most number of lines in the UV spectrum and will decrease with metallicity).

Similarly, there is a solution for the calculation of \( v_{\infty} \). This turns out to be \( v_{\infty} = \sqrt{\frac{\alpha}{1-\alpha}} v_{esc} \), where \( v_{esc} = \sqrt{2(1 - \Gamma_e)GM_*/R_*} \), and is the velocity a particle needs in order to escape the gravitational well of the star. This is dependent only on \( \alpha \).

**Metallicity Dependence**

As discussed above, the chemistry of the wind will have an effect on the radiative forces acting on it and hence the mass–loss rate. The main body of observational evidence for a metallicity dependence on the properties of stellar winds is primarily from the measurement of terminal velocities. Garmany & Conti (1985) and Prinja (1987) both note that \( v_{\infty} \)'s for stars in the Magellanic Clouds are lower than that of Galactic stars.

Theoretically, \( \dot{M} \) is expected to be proportional to the metallicity of the environment to a power exponent (equation 2.11)

\[ \dot{M} \propto Z^m \]

(2.11)
from the studies of Castor et al. (1975), Abbott (1982), and Kudritzki et al. (1987), where $m$ ranges from 0.5 (Kudritzki et al. 1987) to 0.94 (Abbott 1982). These results are generally accepted for main sequence O and B type stars, with the lower 0.5 value being used in the evolutionary tracks of Meynet et al. (1994). However, more recent work (Vink et al. 2001; Leitherer et al. 1992) points towards a mid range value of about 0.8 (0.85 for Vink et al. (2001) and 0.8 for Leitherer et al. (1992)). Furthermore, Leitherer et al. (1992) postulate that $v_\infty$ is also dependent on some power-law scaling, and suggest a exponent value of 0.13.

Although there is some agreement about scaling winds for O and B type stars, there is no real evidence for W-R stars. We may well expect a scaling from the considerations of the theoretical mass-loss solution, but there certainly is no solid observational evidence. Theoretically, uncertainty in W-R wind scaling arises due to the enrichment that these stars have acquired during their lifetimes. The key to determining if W-R winds scale with $Z$ lies in defining the elements/ions responsible for driving the winds. It is unclear whether enrichment by CNO elements could drive a low metallicity outflow in place of Fe, Ni and Zn. If this is true then we would see no metallicity scaling. However, iron group species have many more lines than CNO, with which to drive an outflow (Pauldrach 1998). Recent theoretical work (Hillier, private communication) shows that these heavy metal lines (including other lines such as Ni) are most likely responsible, although to get the correct force to drive the wind, efficiencies of 2–20 are currently needed in the momentum transfer (Nugis & Lamers 2000).

2.1.3 LTE and Non-LTE Atmospheres

Comparing Individual Models

Although a well-built LTE model can satisfactorily describe a cool star's atmosphere, the LTE assumption breaks down as the temperature of the atmosphere rises. LTE can comfortably describe main sequence stars of spectral types from late M to early A, as the mean free path of a photon is sufficiently greater than that of matter. This couples large parts of the atmosphere together, driving the temperature towards an equilibrium and therefore lowering the temperature gradient, so a non-LTE atmosphere is indistinguishable from the simpler LTE model. However in low gravity or high temperature conditions, the LTE assumption breaks down. This means that cool supergiants, LBVs, W-R stars and
O and early B stars are poorly modelled by the LTE assumption. There are several main contributing factors to the break down of LTE, affecting both the fundamental temperature and density structure of the atmosphere and the energy level states:

1. **The radiation field intensity.** Under conditions of extreme radiation field intensity, photon energy is absorbed at a much higher rate than that of spontaneous and collisional emission. This has the effect of raising the number of ions in higher energy states, pushing $\frac{N_{i+1}}{N_i}$ up to be far higher than in an LTE case, thus creating a large departure from the Saha-Boltzmann equation.

2. **Low density environments.** Massive supergiants often have a low density zone at the edge of their atmospheres, especially in their outer wind. This low relative density can deprive the atmosphere of the collisional de-excitation needed to lower atoms into the LTE energy states. This phenomena does not necessarily require the atmosphere to be hot, and can be just as important in low temperature stars.

3. **Large parameter gradients.** Hot stars tend to have extremely steep temperature and pressure gradients, and new models even suggest discontinuities such as "clumped" winds. If the gradient is too steep, the LTE approximation may break down, as the energy balance at one edge of an atmosphere increment is in a different state to the opposite side.

Once it is realised that there may be large differences between LTE and non-LTE models in certain conditions, it is important to quantify the level of difference to justify the use of a far more computationally intensive code.

**The non-LTE Rate Equations**

The equations of statistical equilibrium or rate equations for non-LTE can be represented by equation 2.12, taken from Mihalas (1978). All notation is defined in that reference.

$$
- \sum_{i<j} n_i (R_{ij} + C_{ij}) + n_j \left[ \sum_{i<j} (n_i/n_j)^* (R_{ji} + C_{ji}) + \sum_{k>j}^l (R_{jk} + C_{jk}) \right] - \sum_{k>j}^l n_k (n_j/n_k)^* (R_{kj} + C_{kj}) = 0 \quad (2.12)
$$

This equation represents all radiative and collisional rates for transitions whether they are bound-bound or bound-free. The * denotes a thermal equilibrium value. $n_i R_{ij}$ represents
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all the upward radiative rates where \((i \rightarrow j)\), and \(R_{ij}\) is given by

\[
R_{ij} = 4\pi \int_{\nu_0}^{\infty} \alpha_{ij}(\nu)(h\nu)^{-1}J_\nu d\nu
\]  
(2.13)

\(\alpha_{ij}(\nu)\) is the photoionization cross section at frequency \(\nu\), and the integral is performed from the threshold frequency to infinity \(J_\nu\) is the flux. The downward radiative rates \((j \rightarrow i)\) are represented by \(n_j (n_i/n_j)^*R_{ji}\) where \(R_{ji}\) is given by

\[
R_{ij} = 4\pi \int_{\nu_0}^{\infty} \alpha_{ji}(\nu)(h\nu)^{-1}[(2h\nu^3/c^2) + J_\nu e^{h\nu/k_BT}]d\nu
\]  
(2.14)

In equilibrium the two rates are exactly the same, i.e. \(R_{ij}^* = R_{ji}^*\). For the collisional states, \(n_iC_{ij}\), the total number of upward collisions is denoted

\[
n_iC_{ij} = n_i n_e \int_{\nu_0}^{\infty} \sigma_{ij}(v)f(v)v dv
\]  
(2.15)

where this time \(v\) devotes not frequency but electron velocity. \(\sigma_{ij}(v)\) is the electron cross section and \(f(v)\) is the Maxwellian velocity distribution. \(\nu_0\) denotes the velocity corresponding to the threshold energy of the process. The number of downward transitions is

\[
n_jC_{ji} = n_j (n_i/n_j)^*C_{ij}
\]  
(2.16)

Equation 2.12 is the equivalent of the Saha-Boltzman equation in full non-LTE. From this rate equation, not only can the density of the energy states be calculated, but more importantly the ionization structure of the atmosphere.

**Bulk Differences Between an LTE and a non-LTE Model Grid**

We compare the non-LTE WM-basic model which includes a spherical geometry to that of the more basic plane-parallel static LTE models from the grid of Lejeune *et al.* (1997). It can be the geometry that can often account for differences between the models. Figure 2.2 plots the LTE and non-LTE models of three hot stars, an extreme temperature main sequence dwarf (spectral type 03V), a hot giant of 37.5 kK (spectral type O7.5III), a cooler early B supergiant (spectral type B0.5I). The LTE models have extremely high \(\log g\) values to represent supergiants, since LTE modelling works best in a high gravity regime, so a comparison with the low gravity WM-basic models was not possible. Instead, we pick an evolutionary point from the evolutionary tracks of Meynet *et al.* (1994) and fit an LTE and non-LTE model to it as closely as possible. This sort of comparison is also used in Schaeerer & de Koter (1997) between CoStar and Kurucz models.
Figure 2.2: LTE models of three stars are compared (unbroken lines) with their non-LTE counterparts (broken lines). The plot shows that the hotter models disagree by the largest amount, especially at wavelengths short-ward of 228 Å. The LTE models are taken from the grid of Lejeune et al. (1997) and the non-LTE models are from the grid calculated for this thesis. The ionization edges for H₀(912 Å), He₀(504 Å) and He⁺(228 Å) are also shown.
In the top plot (the hottest star), showing the LTE and non-LTE 50 kK O dwarf models the difference in predicted flux can clearly be seen in the continuum short-ward of 228 Å. The difference in the He$^+$ ionizing continuum for stars this hot can be as much as a factor of $10^3$ more in non-LTE models and is largest in the most extreme models. The giant model (middle) exhibits a different behaviour, with a greater amount of excess non-LTE flux in the $\lambda < 504$ Å He$^0$ ionizing continuum. The non-LTE model has less flux in the He$^+$ ionizing continuum in this case. As explained below, this could be due to the non-LTE line blanketing in the WM-basic models, especially the modelling of the line depths which is greater in a non-LTE regime. The bottom comparison shows that there is still significant non-LTE ionizing continuum excess even at 25kK, although this manifests itself in the He$^0$ continuum. This may be due to the low densities in the outer parts of the supergiant's atmosphere. Figure 2.3 shows the ratio of the Q values predicted from the solar metallicity non-LTE spherically expanding O star models presented in this thesis to that predicted by the LTE plane parallel models of nearest matching parameters compiled by Lejeune et al. (1997). The Q values are defined in equation 2.17, described in Ferland (2002)

$$Q_{0,1,2} = Q(HI, HeI, HeII) = 4\pi R_{\text{star}}^2 \int_{\nu_1}^{\infty} \frac{F_\nu}{h\nu} d\nu$$

where $\nu_1$ is $1.32 \times 10^{16}$Hz for $Q_2$ (228 Å), $5.95 \times 10^{15}$Hz for $Q_1$ (504 Å) and $3.3 \times 10^{15}$Hz for $Q_0$ (912 Å). The concept of the Q value is particularly useful for photoionization applications, since it measures the total ionizing flux in a certain continuum region emitted by an object. Figure 2.3 shows several important differences between the LTE and non-LTE sets of models.

The non-LTE supergiants have a $Q_1$ luminosity excess approaching a factor of 10 compared to their LTE counterparts, especially in the 30-40kK temperature range, although all spectral classes seem to have some small gains in flux over the LTE models above 40kK. This effect is likely due to the non-LTE models differences. H and He can become fully ionized, and electron scattering becomes an important source of opacity in the He$^0$ continuum. Wind effects therefore have little effect on the He$^0$ continuum and a large amount of flux is emitted. Since O supergiants have the highest density winds, line blanketing will be most effective in these spectral classes. We can also see that some O dwarf models have a small deficit of flux compared to the LTE grid in the 30-35kK range. Schaerer & de Koter (1997) also see this phenomenon when comparing their models to the Kurucz grid.
Figure 2.3: The difference between LTE and non-LTE He II ($Q_2$), He I ($Q_1$) and H I ($Q_0$) ionizing continua for solar metallicity dwarf (unbroken), giant (dashed) and supergiant (dotted) models. The LTE models are taken from the grid of Lejeune et al. (1997) and the non-LTE models are from the grid calculated for this thesis.
They believe that this effect is due to the difference between plane-parallel and spherically expanding models, and the addition of a wind outflow which causes high lying energy levels to become optically thin. Subsequent electron cascading can cause the ground state population to increase, lowering the He\(^0\) flux. This is primarily seen in the lower temperature dwarf models in the WM-basic grid. As temperature increases, we revert back to the case of an excess of \(Q_1\) compared to the LTE models.

The biggest non-LTE/LTE differences are in the He\(^+\) continuum, corresponding to \(Q_2\). Here we see the surprising result that the non-LTE supergiants are far less luminous than their LTE counterparts, by a factor of 10\(^3\) for all temperatures greater that 30kK. Giant and dwarf non-LTE models of between 30 and 35kK also contribute up to one hundred times less He \(\text{II}\) continuum flux than in LTE. It is suspected that this is coupled to the excess in \(Q_1\) above, where high density winds seem to be blocking the far UV continuum flux and redistributing it to longer wavelengths.

The \(Q_2\) plot shows no data for temperatures of less than 30kK for any spectral class. This is because the LTE models have no flux in this region of the spectrum at low temperatures and the ratio is infinitely large.

High temperature dwarf and giant models (38-50kK), also show a large excess of far UV flux in the He \(\text{II}\) continuum of up to a factor of 10\(^4\). This is because the continuum is formed in a region corresponding to that of strong depopulation. Schaerer & de Koter (1997) believe this effect results in the strong continuum excess observed above.

The \(Q_0\) ratio is much closer to 1 than any for the other Q values, although we see a slight excess throughout most of the non-LTE models. In this wavelength range, an agreement should be found as the \(\text{H}\(^0\) ionizing continuum is formed under similar conditions in both models, where geometric and LTE/non-LTE assumptions are roughly equivalent. This means that the effective difference of LTE to non-LTE is small and the continuum levels should roughly agree. The dwarf non-LTE models show a small deficit compared to their counterparts, probably due to larger line depths in the lower temperature non-LTE model. At higher temperatures, these larger line depths redistribute He\(^+\) continuum flux to the \(\text{H}\(^0\) continuum increasing its value. The excess of flux experienced by the giant and supergiant models may be due to a slight mismatch between model parameters, since no low gravity LTE models exist, and higher gravity models had to be used with small radii, hence causing the LTE models to exhibit a lower level of flux. The differences in the \(Q_2\) and \(Q_1\) fluxes are large enough to have a significant effect on the predicted nebular line
2.2 Comparisons of Non-LTE Atmospheres

Several freely available, contemporary non-LTE atmospheres exist. Some of these codes are designed to calculate the atmospheres of hot, evolved stars with winds, while others are intended for use in the cool, low gravity environments of low mass stars, where deviation from LTE is just as important. The codes available for the calculation of hot, massive stars are TLUSTY by Hubeny et al. (1994), ISA-WIND by de Koter et al. (1997), CoStar by Schaerer & de Koter (1997), WM-Basic by Pauldrach (1998) and CMFGEN by Hillier & Miller (1998). Other codes do exist privately, but were not at the time available for public use. These codes employ various different means of calculating the state variables of the atmosphere, including the use of spherically expanding geometry which includes a representation of the stellar wind in the model atmosphere. This geometry is suitable for not only O stars, but for more evolved stars such as O supergiants and even the complicated W-R stars. Another more simple way is to neglect the effect of winds and use a plane-parallel geometry, or to use a Monte-Carlo simulation to determine the state of the atmosphere, or even a combination of the two. These methods are described below.

2.2.1 Plane Parallel Models

Although young O stars have powerful UV radiation fields, their stellar winds are relatively weak compared to their evolved descendants. This means that in some cases it is possible to use the less complicated assumption of hydrostatic equilibrium used in the plane parallel geometry described in §2.1, or the complex fluid dynamical treatment representing an atmosphere with a dense wind, to get the same result. Other massive stars such as O supergiants and W-R stars have strong winds and must be modelled using an out-flowing geometry. TLUSTY is an example of a truly non-LTE code that relies on a plane parallel type of assumption. This assumption allows for a much simpler method of computation. It is possible to calculate an atmosphere far more quickly than the outflow models, due to the simplified macroscopic state physics. This means that not only are the populations of the atomic species calculated in full non-LTE, but line blanketing and the effect of the lines on the continuum are also fully non-LTE. Early non-LTE models used a non-LTE case to calculate the number of each atomic species, while performing line blanketing in LTE e.g.
Figure 2.4: Plotted are the non-LTE models of an O star with a temperature of 25kK, with an effective gravity of \( \log g = 4.0 \). The WM-basic and CMFGEN models are spherically expanding with full treatment of line blanketing, CoStar uses the \texttt{isa-wind} spherically expanding model structure, with a Monte-Carlo line blanketing approximation. The \texttt{TLUSTY} model here treats just hydrogen and helium for line blanketing in a plane-parallel geometry.
Eastman & Kirshner (1989), since the effect of a non-LTE rather than an LTE regime in the calculating of ionic populations is larger than that of line blanketing (Schaerer & Schmutz 1994).

An output continuum of TLUSTY model is compared to three spherically expanding outflow models in Fig. 2.4 for a 25kK early B star for which the stellar wind is sufficiently weak to allow for a plane parallel approximation. This figure illustrates that although the model produces the ionizing continuum in good agreement with the other models, no emission lines are produced. This is because the physical conditions for emission to occur are found in the outward moving wind of a star, so therefore the emission is absent. However, this model does have the advantage that the absorption lines produced in the photosphere (the TLUSTY model presented in Fig. 2.4 has only helium and hydrogen atomic data included in it, although a wider range of data is available for new calculations) give excellent agreement with observations of stars with weak winds, due to the high resolution description of the photosphere.

2.2.2 Spherically Expanding Wind Models

There are several non-LTE codes which use a spherically expanding geometry. Here we review three of them; CMFGEN, WM-basic and CoStar. These models use the assumption that the atmosphere can be approximated by an expanding shell of wind driven from the stellar interior by radiation pressure (all magnetic fields are neglected). These codes fall into two sub-categories, codes that use co-moving frame geometry (historically Mihalas et al. (1975)), currently CMFGEN and CoStar, or use a static wind geometry and the Sobolev approximation (Castor et al. 1975). These codes have historically used few atomic species in their calculations such as H, He and occasionally C, N and O but neglecting iron group elements, since these have thousands of lines and require large calculations (Hillier (1990); Hamann (1985); Hamann & Schmutz (1987); Gabler et al. (1989); de Koter et al. (1993)).

Recently, computing power has increased sufficiently to allow codes to use many atomic species under what is known as the super-level approximation. This method, used by both WM-basic and CMFGEN, groups together the influence of lines of similar energy level. This method is described in the CMFGEN section. Another factor in the improvement of atmosphere codes in general is that since the conclusion of the Opacity Project (OPACITY, Seaton (1987)), a large atomic dataset has become available. This allows for the inclusion
of reliable model atoms. Models such as cmfgen have benefitted hugely from this and now include atomic species such as Fe, Si, S, Ne, Ar, P and Ca in their calculations.

The WM-basic Code

Pauldrach et al. (2001) developed the Wind Momentum - (visual) basic code to be used on average speed Intel type processors with a reasonably fast calculation time (~ 3 hours). This meant that some limiting approximations had to be made in their model. The WM-basic code uses a homogeneous, stationary and spherically symmetric radiation driven wind model. In this model the expansion of the atmosphere is not assumed, but rather driven by the scattering and absorption of Doppler shifted spectral lines. This mechanism is famously discussed in Castor et al. (1975), and reviewed in §2.1.2. The main iterative loop of the code is shown in Fig. 2.5. The calculation starts with a spherical grey LTE model which gives a first guess at the temperature and continuum radiation pressure of the atmosphere. Then the calculation couples two routines to generate a temperature and density structure which is repeated as an iterative cycle until the solution converges. The first routine is the hydrodynamic iteration cycle which computes the velocity field and the temperature and density structure of the wind. This information is then fed into the second non-LTE spherical routine which is the most complex part of the model where the emissivity $\eta$ and opacities $\kappa$ of the continuum are calculated. As the opacity calculation is extremely large, involving all of the energy levels of all atomic species present in the wind, the authors decided to approximate this calculation by using an “opacity sampling” technique. Instead of every transition being accounted for, the opacity of the gas is calculated at discrete intervals spaced logarithmically along the continuum starting from approximately 50 Å. This method can approximate to the real opacity to within a few percent by using only a few thousand grid points (Pauldrach et al. 2001). The last two iterations do not use this opacity sampling technique, instead a more rigorous calculation is used to obtain a solution that will converge. In all 149, ionization stages of the 26 elements H to Zn are considered, apart from Li, Be, B, and Sc. Overall, this code works extremely well at reproducing the observed spectra of most O stars, as shown by the example in Fig. 2.6. This plot is reproduced from Pauldrach et al. (2001) and shows their best fit to the O3V((f)) star HD 93250. The C III, IV and V P Cygni profiles are modelled extremely well in this case.
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Figure 2.5: A flowchart of the main processes in the WM-basic program. This figure is taken from Pauldrach et al. (2001).
Figure 2.6: Plot showing a best fit WM-basic spectra to that of an observation of the O3V((f)) star HD 93250. This figure is taken from Pauldrach et al. (2001).
The CMFGEN code

The Co-Moving Frame GEN atmosphere code, models the stellar atmosphere as an outflowing wind in the co-moving frame. The code defines the geometry of the model through input parameters, so assumes a wind already in motion, with no attempt to calculate the radiation pressure in the atmosphere. To this end the user defines the velocity field through:

\[
v(r) = \frac{V_0 + (V_\infty - V_\text{ext} - V_0)(1 - R_*/r)^{\beta_1} + V_\text{ext}(1 - R_*/r)^{\beta_2}}{1 + (V_0/V_{\text{core}})e^{\exp[(R_*/r)/\kappa_{\text{eff}}]}}
\]  

(2.18)

This equation allows for two power laws to define the velocity field, \( \beta_2 \) in the outer part of the wind and \( \beta_1 \) in the inner part, which must be specified along with \( V_\infty \) (the velocity of the wind at infinite distance from the star) and \( V_{\text{core}} \) (the velocity at the base of the wind). This method of modelling the atmosphere allows for a more complex wind structure to be added, so that wind “clumping” can be considered. Clumping is the term given to an extremely inhomogeneous wind, in which high density bubbles of gas exist in a lower density medium. Observationally, clumping in stellar winds has been known about for some time, and manifests itself in the variability of line profiles with periods in the order of days (Prinja et al. 1992). Recently evidence also exists, via the shapes of emission line profiles, that W-R stars exhibit the clumped wind characteristics. Evidence for this comes from observations of line profile variability (e.g. Moffat et al. 1988), the slope of the infrared and radio spectrum (Nugis 1994; Nugis et al. 1998), and the observed weakness of the electron scattering wings compared to theoretical line profiles (Hillier 1991). The radiative transfer of the stellar atmosphere is treated exactly in CMFGEN, meaning that no opacity redistribution or sampling techniques are used and the opacity is calculated implicitly for all atomic species. As this technique is extremely computer intensive however, assumptions are made about the model atoms used in the calculation. Instead of all the energy levels being used in the atomic model, energy levels of a similar value are “packed” together to form what is known as a super-level approximation of the ion. A super-level is used to represent several spectral lines in the same ion which have very similar energy levels. The super-level is made to exhibit the bulk transition properties of the group. The super-level approximation is a good way of lowering the number of levels in the atom, providing the new levels are appropriately normalised. This assumption is not only made in CMFGEN, but also in WM-basic, as it lowers the calculation size considerably, with very little loss of accuracy (Hillier & Miller 1998).
The CoStar Code

The Combined Stellar structure and atmosphere or CoStar model is a combined evolutionary and atmosphere code. The atmosphere section is based on that of the isA-wind code of de Koter et al. (1997) with non-LTE treatment of all major species from H to Zn. This atmosphere section is solved in three steps:

1. **Step 1: The Atmosphere Solution.**
   Atoms of hydrogen and helium are used to simultaneously solve for the statistical equilibrium, with other elements used in approximate treatments. The radiative transfer is solved in the co-moving frame using a spherically expanding atmosphere. The solution is based on that of Hamann & Schmutz (1987) and Wessolowski et al. (1988). The temperature structure is solved using H and He and improved using two techniques:
   - A modified version of the Karp (1972) temperature correction procedure for optical depths of order 1 or larger. Below this value a Grey atmosphere is used.
   - The energy equation is satisfied, including the cooling of the atmosphere via C, N, O, Ne, Ar, Si and Fe in a nebular approximation. This technique may be inaccurate for many stars since the nebular approximation only applies at low density.

   Both these techniques are approximate and tend to underestimate the true temperature, therefore the highest temperature solution is used.

2. **Step 2: Opacity Sampling.**
   In order to obtain a realistic solution, line blanketing is included in the model’s framework. This is accomplished by the intensity-weighted mean opacities of bound-free and free-free transitions. These weights are evaluated via a Monte-Carlo technique, and generate the ionization structure of the atmosphere required for the formal solution. The full line blanketing is performed via a main Monte-Carlo routine described by Schmutz (1991) and also in Abbott & Lucy (1985) and Lucy & Abbott (1993). The formal solution of the equation of transfer is performed in the differentially expanding outflow. This allows for scattering by electrons, the effect of thousands of overlapping lines and continuum scattering and absorption processes.

The final step is to solve for $M$ and $v_\infty$ in a similar way to WM-basic. The code works in the iterative way described in Schaerer et al. (1996). An LTE model is fed in from step 1 and combined with solutions from step 3 to input into step 2. The output from step 2 is fed back into steps 1 and 3. The process is repeated until convergence.

This Monte-Carlo approach has the advantage that millions of lines can influence the emergent spectrum. Disadvantages however, include the lack of line broadening and line blanketing in the low velocity part of the wind providing an over estimate of ionizing flux.

2.3 O and B Star Grid

For this work, it was concluded that a new grid of models must be computed that covers the entire upper H-R diagram for massive stars with stellar winds to replace older, inaccurate models. This section deals with the calculation of main-sequence O and early B dwarf stars and their slightly more evolved descendants, the giants and supergiants. Observationally, the main distinction between the dwarfs and more evolved species is mostly one of a luminosity difference, where a supergiant can be up to two magnitudes brighter than its dwarf spectral counterpart (see Fig. 2.8). Other more subtle differences point towards an evolutionary connection, such as lower surface gravity, stronger winds and more evolved surface chemistry.

From the non-LTE models discussed above, it is clear that a plane-parallel model must be discarded since it ignores outflows caused by a stellar wind. Since the CoStar models may have underestimated line blanketing, and a low resolution, publicly available grid already exists, it was decided that a larger grid, as required by the scope of this thesis should include a full treatment of line blocking and blanketing. For this reason CoStar is excluded. It was decided that WM-basic should be used, due to the availability of a 350 MHz Pentium II PC and a short calculation time of approximately 3 hours. The differences in the predicted ionizing fluxes of the CoStar and WM-basic codes are discussed later in this section.

Models were calculated for five metallicities: $0.05, 0.2, 0.4, 1, \text{ and } 2 Z_\odot$, as defined by the evolutionary tracks of Meynet et al. (1994), since the main application for the new models is the population synthesis code Starburst99 (Leitherer et al. 1999), which uses these tracks. Each metallicity has 33 models calculated spanning the surface gravity –
temperature parameter space required to represent all O and early B stars, from young main sequence dwarf stars, with typically a $\log g = 4.0$, to evolved supergiants with surface gravity $\log g < 3.1$. Three types of spectral class were considered: dwarf, giant and supergiant, each class having 11 temperatures to represent a large range of spectral types. The giant models were not calculated with WM-basic, but instead were interpolated from the dwarf and supergiant grid. The interpolation was carried out via a simple linear routine, using the nearest temperature dwarf and supergiant models as reference in a wavelength point by wavelength point approach. This simple method of interpolation yields a continuum distribution consistent with a calculated model of the same parameters.

A 26kK giant model is compared in Fig. 2.7 with its interpolated counterparts, showing
remarkable good agreement. Tests show that the interpolated model continuum is extremely close to its calculated counterpart for temperatures of 25 ≤ Teff < 40kK, with some over prediction of the He+ continuum above those temperatures. However this deviation is minimal and the use of high temperature O III models is minimal in the SB99 (starburst99; (Leitherer et al. 1999)) code due to their relatively low number predicted by the evolutionary tracks. This method saves approximately 165 hours of computing time, with little loss of accuracy. Although the flux distribution is well produced by a simple interpolation method, there is no guarantee the the total flux of these interpolated models is conserved, as interpolation gives no allowance for flux conservation. To this end, careful consideration must be taken when applying these models to physical situations. To ensure flux conservation, one must use the following formula:

$$F_{\text{cond}} = \int_0^\infty \frac{F_{\text{int}}}{B_\lambda(T)} d\lambda$$

Where $F_{\text{cond}}$ is the conserved flux, $F_{\text{int}}$ is the interpolated flux and the Planck function $B_\lambda$ is used to normalise the flux since the overall energy is conserved in this form, and is easy to calculate. The main intended use for the O star model atmospheres is discussed in Chapter 4, where their implementation into Starburst99 is presented.

Tables 2.1, 2.2 and 2.3 present the entire OB model grid that was computed for this thesis.

2.3.1 Definition of the Fundamental Atmosphere Parameters

In order to create a sufficiently accurate model O and B star grid, we must define several fundamental parameters using the most accurate data available. These parameters are the effective temperature $T_{\text{eff}}$ of the stellar atmosphere; the surface gravity, usually expressed as $\log g$ in cgs units (cm s$^{-1}$) and photospheric radius $R_*$. Since the ability to represent a reliable stellar atmosphere rests on the accuracy of these parameters, great care was taken in their definition. The following approach was used to define these parameters;

First we needed to use a temperature luminosity calibration for O stars, to link the gravity of a star to a luminosity in order to produce three grids to represent the dwarfs, giants and supergiant spectral classes. The most obvious source for this was a study of Schmidt-Kaler (1982). However, this work is not reliable for extremely massive stars, since few were known to exist at the time of publication, in fact so few that no O3 type stars had been observed. It was decided to create a new $M_V$– spectral type scale via a survey of
Table 2.1: OB dwarf model atmosphere grid

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<th>$0.2 Z_\odot$</th>
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Notes: Mass-loss rates $\dot{M}$ and terminal velocities $v_\infty$ are in units of $M_\odot\, \text{yr}^{-1}$ and km s$^{-1}$ respectively. $Q_0, Q_1$ and $Q_2$ are on a log scale and represent the photon luminosity (s$^{-1}$) in the H I, He I and He II ionizing continua.
Table 2.2: OB giant model atmosphere grid

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Notes: Mass-loss rates $\dot{M}$ and terminal velocities $v_\infty$ are in units of $M_\odot$ yr$^{-1}$ and km s$^{-1}$ respectively. $Q_0$, $Q_1$ and $Q_2$ are on a log scale and represent the photon luminosity (s$^{-1}$) in the H I, He I and He II ionizing continua.
Table 2.3: OB supergiant model atmosphere grid

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<td>46.4</td>
<td>32.8</td>
<td>49.0</td>
<td>46.6</td>
<td>34.6</td>
<td>49.0</td>
<td>46.4</td>
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<td>27.8</td>
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<td>1500</td>
<td>-6.45</td>
<td>1210</td>
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</tr>
<tr>
<td>B0I</td>
<td>48.4</td>
<td>45.6</td>
<td>32.1</td>
<td>48.3</td>
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<td>33.0</td>
<td>48.4</td>
<td>45.5</td>
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<td>2.99</td>
<td>26.3</td>
<td>-6.08</td>
<td>1400</td>
<td>-6.64</td>
<td>1130</td>
<td>-6.40</td>
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<td>B0I</td>
<td>47.8</td>
<td>45.2</td>
<td>...</td>
<td>47.5</td>
<td>44.8</td>
<td>31.8</td>
<td>47.6</td>
<td>45.0</td>
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<tr>
<td>OB#33</td>
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<td>2.95</td>
<td>26.9</td>
<td>-6.11</td>
<td>1200</td>
<td>-6.67</td>
<td>970</td>
<td>-6.43</td>
</tr>
</tbody>
</table>
| B0.5I      | 47.6      | 44.9  | ...     | 47.1  | 44.2   | 31.5   | 47.4   | 44.6   | ...    | 47.7   | 45.1   | ...    | ...    | ...

Notes: Mass-loss rates \( \dot{M} \) and terminal velocities \( v_\infty \) are in units of \( M_\odot \text{yr}^{-1} \) and \( \text{km s}^{-1} \) respectively. \( Q_0, Q_1 \) and \( Q_2 \) are on a log scale and represent the photon luminosity (s\(^{-1}\)) in the H\(_1\), He\(_i\) and He\(_{II}\) ionizing continua.
the existing observations, using the work of Crowther & Dessart (1998) which included a sample of Galactic and LMC stars, for spectral types O3 to O5. The scale for later types (O5.5 to B1.5) was determined manually by using the references cited in that paper. A fit was made to individual absolute magnitudes as illustrated in Fig. 2.8.

This $M_V$–spectral type calibration was then transformed to luminosity–temperature by the simple use of the well known $M_V$ luminosity relation, using the $T_{\text{eff}}$-bolometric correction calibration of Vacca et al. (1996) to convert $M_V$ to luminosity, and a spectral type temperature calibration to calculate temperature from spectral type. The temperature calibration used was that of Crowther (1998) presented in Table 2.4.

<table>
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<tr>
<th>Spectral Type</th>
<th>Spectral Classification</th>
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<td>O 3.0</td>
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<tr>
<td>O 4.0</td>
<td>44000 47000 46500</td>
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<tr>
<td>O 5.0</td>
<td>43000 45000 43500</td>
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<td>O 6.0</td>
<td>42000 41000 40000</td>
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<td>O 7.0</td>
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<td>O 8.0</td>
<td>35000 36000 35750</td>
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<td>O 9.0</td>
<td>31800 33000 34000</td>
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<tr>
<td>O 9.5</td>
<td>30800 31000 32500</td>
</tr>
<tr>
<td>B 0.0</td>
<td>29800 30000 28500</td>
</tr>
<tr>
<td>B 0.5</td>
<td>28500 27500 27000</td>
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<tr>
<td>B 1.0</td>
<td>26100 26000 23500</td>
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<tr>
<td>B 1.5</td>
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<td>B 2.0</td>
<td>21300 22000 20000</td>
</tr>
<tr>
<td>B 2.5</td>
<td>19000 19500 19500</td>
</tr>
</tbody>
</table>

The Crowther (1998) calibration is updated relative to that of Vacca et al. (1996) which uses helium-hydrogen plane-parallel non-LTE models that neglect line blanketing and wind effects to determine the temperatures. Crowther (1998) improved on this determination by
including some allowance for stellar winds. Models have now become more accurate, the temperature scales have changed, even in fact since the creation of this OB star grid. New work (Martins et al. 2002; Crowther et al. 2003a) suggests lower temperatures due to use of more realistic non-LTE spherically expanding models with line-blanketed atmospheres. Unfortunately, these changes are relatively large, revising temperature scales downward by \( \sim 3-4kK \); (Martins et al. 2002) or \( \sim 8kK \) (Crowther et al. 2003a). Once the \( M_V \)-spectral type calibration was converted to that of luminosity-temperature, it was smoothed by way of averaging to avoid lower temperature stars being more luminous than their hotter counterparts in the same spectral class.

This new calibration is presented in Figure 2.9, and compared with that of Schmidt-Kaler (1982). The largest difference is that the new work gives lower luminosities for the hottest stars, and lowers the overall luminosities of the giant stars.

The effective temperature grid was calculated so that it would span the domain of O to
Figure 2.9: The adopted luminosity - temperature calibration of this thesis. The new calibration is the solid line, whereas the old standard, the Schmidt-Kaler work is the broken line.
early B stars from 25 000–51 000 K, in log space at typically 0.03 dex intervals to give 11 models covering the spectral classes from O3 to B1.5, where spectral types become more closely grouped, as the spectral lines change rapidly with temperature around O8 to B2. A temperature–luminosity–radius calibration was used to calculate the photospheric radii for each model.

\[
\log R_{\ast, cgs} = 10.8426 + 0.5 \log \frac{L_{\ast}}{L_{\odot}} - 2.0 \log T_{\text{eff}} + 7.52 \tag{2.20}
\]

Since a large discrepancy remains between masses derived from spectroscopic and evolutionary models (Herrero et al. 1992), we calculated the mean surface gravities in log space from Vacca et al. (1996) in all cases.

Since it is well known that the terminal velocity \( v_\infty \), scales with the energy output from a star (§2.1.2) we used the \( v_\infty \)-spectral type calibrations of Prinja et al. (1990) and Lamers et al. (1995), converting spectral type to temperature with our adopted scale. These values were averaged and then smoothed to give a linearly rising distribution with temperature. Although models relating wind density to the fundamental parameters of stars do exist, (e.g. see explanation in §2.1.2) these are still not reliable, so it was decided that an empirical relation would be the best option. The mass–loss rate \( \dot{M} \) was derived from the wind momentum–luminosity relationship of Kudritzki & Puls (2000):

\[
\log \left[ \dot{M} v_\infty \left( R_\ast / R_\odot \right)^{0.5} \right] = \log D_0 + x \log \frac{L}{L_\odot} \tag{2.21}
\]

where \((\log D_0 \text{ and } x) = (20.69, 1.51) \) (for O supergiants); 19.87, 1.57 (for O dwarfs); and 21.24, 1.34 (early B supergiants). Since we wished to create a grid of five metallicities, some \( \dot{M} \) and \( v_\infty \) scaling must be adopted. The reason behind a scaling law is discussed in §2.1.2. Several power exponents are predicted for this scaling, lying between 0.5 and 1.0. The prescription of Leitherer et al. (1992) was chosen as the scaling for both \( \dot{M} \) and \( v_\infty \) with metallicity by using power law exponents of 0.8 and 0.13 respectively

\[
\dot{M}_Z = Z^{0.8} \dot{M}_\odot; \quad v_{\infty, Z} = Z^{0.13} v_{\infty, \odot} \tag{2.22}
\]

This choice of the Leitherer et al. (1992) value gives consistency with Starburst99. The scaled values for \( \dot{M} \) and \( v_\infty \) are given for each metallicity in Tables 2.1, 2.2 and 2.3.

WM-basic was used to compute the grid for the reasons explained above. The latest version at the time of calculation was v 1.16. Since this code models both the wind structure and the emergent flux of the star, input parameters were not only those defined
above, but also the "force multiplier parameters" of Castor et al. (1975), explained in §2.1.2. As a mass-loss rate and terminal velocity is not obvious with the value of \( k, \alpha \) and \( \delta \), a manual iterative technique was used to achieve the wind density values previously defined for the models. An exact match for each value was not practical, so instead the values quoted in Tables 2.1, 2.2 and 2.3 are accurate to within 5 per cent for \( \dot{M} \) and to within ±100 km s\(^{-1}\) for the \( v_\infty \).

### 2.4 Model Atmosphere Comparisons

#### 2.4.1 O Star Models

The first line-blanketed non-LTE grid of O star model atmospheres was introduced by Schaerer & de Koter (1997, hereafter SK97). The CoStar models are used extensively in single star H\( \text{II} \) region analyses (e.g. Oey et al. (2000)) and spectral synthesis studies (e.g. Stasińska et al. (2001)). The CoStar grid consists of 27 models covering the main sequence evolution of O3 to early B spectral types at metallicities of solar and 0.2 \( Z_\odot \). The mass-loss rates are taken from Meynet et al. (1994) and are scaled with metallicity using an exponent of 0.5. To compare the WM-basic with the CoStar models, two very different models have been chosen for comparison, a hot \( (T_{\text{eff}} = 43 \text{ kK}; \text{OB#25}) \) early O supergiant model with a strong wind, and a cool \( (T_{\text{eff}} = 33 \text{ kK}; \text{OB#7}) \) late O dwarf model with a weak wind. Figs. 2.10 and 2.11 show the comparisons at solar and at 0.2 \( Z_\odot \) and also plot Kurucz (1992) LTE plane parallel models for reference. The details are displayed in Table 2.5 below.

If we first consider Fig. 2.10, we can see significant differences between the WM-basic and CoStar models, especially in the far UV emergent flux. In the top plot, the solar metallicity case, we see that the WM-basic model has no flux below the \( \text{He}^+ \) edge, in contrast to the CoStar model, which has an observationally significant value. This difference is probably due to the different line-blanking approximations, where WM-basic has a flux below that even that of the LTE model (as discussed in §2.1.3). At lower metallicity (bottom plot Fig. 2.10), we now see the \( \text{He}^+ \) ionizing flux of the WM-basic model is significantly increased. This is due to the lowering of the wind density due to our adopted mass-loss and \( v_\infty \) scaling. Since the wind is more transparent the emergent flux resembles that of the CoStar model, while the LTE case produces far fewer ionizing photons. The differences between CoStar and WM-basic may well be the line
Table 2.5: Reference table showing comparison model parameters presented in Figs. 2.10 and 2.11 for solar metallicity only. 0.2 \( Z_{\odot} \) parameters are scaled as described in the text.

<table>
<thead>
<tr>
<th>Code</th>
<th>Model #</th>
<th>( T_{\text{eff}} ) (kK)</th>
<th>( \log g )</th>
<th>( R_* ) (( R_{\odot} ))</th>
<th>( \log \dot{M} ) (( M_{\odot} ) yr(^{-1} ))</th>
<th>( v_{\infty} ) (km s(^{-1} ))</th>
</tr>
</thead>
<tbody>
<tr>
<td>WM-basic</td>
<td>OB#25</td>
<td>42.6</td>
<td>3.67</td>
<td>19.7</td>
<td>-4.92</td>
<td>2300.0</td>
</tr>
<tr>
<td>WM-basic(0.2Z_{\odot})</td>
<td>OB#25</td>
<td>42.6</td>
<td>3.67</td>
<td>19.7</td>
<td>-5.48</td>
<td>1860.0</td>
</tr>
<tr>
<td>CoStar</td>
<td>#E2</td>
<td>42.6</td>
<td>3.71</td>
<td>20.0</td>
<td>-4.88</td>
<td>2538.0</td>
</tr>
<tr>
<td>CoStar(0.2Z_{\odot})</td>
<td>#E3</td>
<td>42.6</td>
<td>3.71</td>
<td>20.0</td>
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<td>2059.0</td>
</tr>
<tr>
<td>Kurucz</td>
<td>-</td>
<td>42.5</td>
<td>5.0</td>
<td></td>
<td></td>
<td></td>
</tr>
</tbody>
</table>

---

If we move on to the cool dwarf models plotted in Fig. 2.11, both metallicity cases show a WM-basic model that is close to the LTE model of Kurucz (1979), yet predicting lower far UV fluxes. Sellmaier et al. (1996) suggested that this may be due to the non-LTE effect of line blanketing causing deeper line cores and reducing the flux in these regions where many lines form. Schaerer & de Koter (1997) also notice this excess compared to LTE models and find that the temperature of the wind is too high in the region that the He\(^0\) ionizing continuum forms. They caution the use of these models. A comparison of the WM-basic O star grids with application to observation is carried out in the next Chapter.
Figures 2.12 and 2.13 show the $Q_0$ and $Q_1$ continuum luminosity values versus absolute luminosity. On all of the plots, we can see that an O7.5I actually emits roughly the same amount of hydrogen ionizing flux as an O3V star, making them extremely important contributors to the overall ionizing flux. A B0V star emits an insignificant level of flux compared to the giants and supergiant stars, producing almost 1000 times less ionizing flux than the most extreme O stars. In the He$^0$ ionizing continuum, only the hottest stars contribute significantly. A B0 dwarf star contributes less than a ten thousandth of the flux to the He$^0$ continuum compared to an O3 supergiant.

2.5 Wolf-Rayet Star Grid

As W-R stars are extreme in terms of both their wind properties and temperatures, there have already been attempts to model their atmospheres with spherically expanding models. Schmutz et al. (1992) created a pure helium grid to try to reproduce their ionizing properties, however since basic parameters of W-R stars have been substantially improved in recent years and due to the availability of increasingly sophisticated model atmospheres, this grid is found to be more and more inaccurate (Crowther 1999). New codes are now able to reproduce the optical signatures of W-R stars, including their complex emission line spectra (see §2.2 or the conference reviews of Hillier (1999) and Crowther (1999)).

To complement our new O star grid, a grid of line-blanketed WN and WC model atmospheres was calculated with the CMFGEN code of Hillier & Miller (1998) using what is considered to be the most accurate parameters available for these stars.

Since line blanketing is extremely important in W-R stars (Crowther et al. 1999), we included as many dominant ionization states in the W-R atmospheres as possible within the time and computer memory constraints imposed on us. H, He, C, N, O, Ne, Si, S, Fe were included in WN stars and He, C, O, Si, Fe in WC stars. Computational advances and wider ranging atomic data have permitted many other elements and ionization stages to be included in CMFGEN since the completion of the grid (including Ni, Ca and Ar). However, tests now indicate that this will not add to greater accuracy in the ionizing continuum of a W-R spectrum, but may allow for better fits of emission lines.

Tables 2.6 and 2.7, present the parameters adopted in the WN and WC grids, parameterised by the stellar temperature $T_{\text{wind}}$, the stellar radius $R_*$ (both defined at a Rosseland optical depth of $\sim 20$), and the helium mass fraction $Y$. 

Figure 2.10: Comparison of the emergent fluxes for an early O supergiant from a WM-basic (solid), a CoStar (dashed), and a Kurucz model (dotted) at solar metallicity and 0.2 $Z_\odot$. The model details are presented in Table 2.5. The model parameters for 0.2 $Z_\odot$ are scaled as discussed in the text.
Figure 2.11: Comparison of the emergent fluxes for a late O dwarf from a WM-basic (solid), a CoStar (dashed), and a Kurucz model (dotted) at solar metallicity and $0.2 \, Z_{\odot}$. The model details are presented in Table 2.5. The model parameters for $0.2 \, Z_{\odot}$ are scaled as discussed in the text.
Figure 2.12: A $Q_0$ vs log ($L/L_\odot$) plot of the new 0.05, 0.04 and 2.0 $Z_\odot$ metallicity O star grids, for all three effective gravities; supergiants are plotted with pentagonal symbols (dotted, green line); giants with square symbols (dashed, red line) and dwarfs with triangular symbols (solid, black line). Also marked are the symbols which represent the O3, O7.5 and B0 spectral types.
Figure 2.13: A $Q_1$ vs log $(L/L_\odot)$ plot of the new 0.05, 0.04 and 2.0 $Z_\odot$ metallicity O star grids, for all three effective gravities; supergiants are plotted with pentagonal symbols (dotted, green line); giants with square symbols (dashed, red line) and dwarfs with triangular symbols (solid, black line). Also marked on are the symbols which represent the O3, O7.5 and B0 spectral types.
Table 2.6: WN model atmosphere grid

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<th>Model Ref.</th>
<th>$T_\ast$ (kK)</th>
<th>$R_\ast$ ($R_\odot$)</th>
<th>$Y$</th>
<th>$\frac{Z}{Z_\odot}$</th>
<th>$\frac{0.2 Z}{Z_\odot}$</th>
<th>$\frac{0.4 Z}{Z_\odot}$</th>
<th>$\frac{0.05 Z}{Z_\odot}$</th>
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<td>-4.93 1100</td>
<td>-5.49 890</td>
<td>-5.25 970</td>
<td>-5.97 740</td>
<td>-4.69 1200</td>
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<td>-4.59 1490</td>
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<td>-5.01 1730</td>
<td>-5.73 1320</td>
<td>-4.45 2140</td>
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<td>-5.73 2110</td>
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Notes: Mass–loss rates $\dot{M}$ and terminal velocities $v_\infty$ are in units of $M_\odot\,yr^{-1}$ and $\mathrm{km\,s^{-1}}$ respectively. $Q_0$, $Q_1$ and $Q_2$ are on a log scale and represent the photon luminosity ($s^{-1}$) in the $\mathrm{H\,\i}$, $\mathrm{\text{He\,\i}}$ and $\mathrm{\text{He\,\ii}}$ ionizing continua.
Table 2.7: WC model atmosphere grid

<table>
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<th>Model Ref.</th>
<th>$T_\ast$ (kK)</th>
<th>$R_\ast$ (R$_\odot$)</th>
<th>$Y$</th>
<th>$Z_\odot$</th>
<th>0.2 $Z_\odot$</th>
<th>0.4 $Z_\odot$</th>
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<td>$T_\ast$</td>
<td>$R_\ast$</td>
<td>$Y$</td>
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<td>$\log \dot{M}_{Q_1}$</td>
<td>$\log \dot{M}_{Q_1}$</td>
<td>$\log \dot{M}_{Q_2}$</td>
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<td>1010</td>
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Notes: Mass-loss rates $\dot{M}$ and terminal velocities $v_\infty$ are in units of $M_\odot$ yr$^{-1}$ and km s$^{-1}$ respectively. $Q_0$, $Q_1$ and $Q_2$ are on a log scale and represent the photon luminosity (s$^{-1}$) in the H$_1$, He$_1$ and He$_{II}$ ionizing continua.
CHAPTER 2. STELLAR ATMOSPHERES

The definition of W-R parameters is still extremely uncertain, as at the time of calculation, the tools for their determination were extremely new and observations quite scarce compared to O stars. In order to define these parameters, we used the results from the most reliable studies of W-R stars using line-blanketed model atmospheres as an analytical tool. We decided on the calculation of two separate grids, a WN and a WC grid, since these subtypes have very different chemistry which may effect the outcome of the emergent continua. These subtypes also have different luminosity and mass-loss properties. For the WN grid, we defined a temperature range of $30000 - 120000K$, to cover all common WN spectral types. In order to mimic the surface chemistries of the Meynet et al. (1994) evolutionary grid, the WN stars were split into two separate groups, in which the late cool WNL types were assigned a decreasing surface hydrogen content from 24 per cent for the coolest star ($30000K$) to 2 per cent for the boundary between WNL and the early WNE subtype at $60000K$. These ratios were quantified by referring to the work of Crowther & Smith (1997). The surface chemistry for all other atoms was determined with reference to the studies of Herald et al. (2001) and Hillier & Miller (1998). All WNE stars were assigned the same surface chemistry, with surface hydrogen content at 2 per cent. In all WN models a luminosity of $L = 3 \times 10^5 L_\odot$ was assumed. WN radii were calculated using a simple temperature–luminosity–radius calibration (see Equation 2.20, §2.3.1). Although luminosities of all W-R stars do in fact usually vary, with the relative intensity of the emission lines changing with luminosity (the Baldwin effect, named after a similar effect in Quasar data (Baldwin 1977)), it was decided that this extra parameter (i.e. luminosity) would increase the number of required models disproportionately to the extra accuracy the data set would gain. This is especially poignant considering the large amount of time required to calculate the W-R grids. Each model took between 6 to 12 hours to converge, with many models having to be rerun due to problems with convergence. This made the total computing time approximately 1440 hours (60 days). Adding an extra dimension to the grid, even with only 3 points would increase this value to half a year of solid processing time. Also, it was not clear whether the different luminosities represented a scatter around a mean value, or if there is a trend to the data.

For the WC grid, a temperature range of $40000 - 140000K$ was chosen. This slightly higher scale reflects the observed range of temperatures. The surface chemistry was set constant for all WC types, with the C/He=2 and C/O=4 by number. Again surface chemistry and radii were taken from several references which have analysed this W-R
CHAPTER 2. STELLAR ATMOSPHERES

subtype, (Hillier & Miller (1999); Dessart et al. (2000) and Crowther et al. (2003b)). The radii were calculated using the temperature–luminosity–radius calibration (see Equation 2.20, §2.3.1). The luminosity was set at $L = 2 \times 10^5 L_\odot$ for all WC models.

The wind density effects in W-R winds are extremely important to the spectral appearance of the star. For this reason great care must be taken in assigning mass–loss and terminal velocity values. Since the winds of W-R stars are believed to have an inhomogeneous or clumped structure, an empirical relationship must reflect this fact. The significance of clumping in this context, is that it leads to a reduction in the determination of W-R mass–loss rates by up to a factor of 5 (Nugis et al. 1998). To derive a representative set of WN mass-loss rates for the WN grid, we used the mass-loss–luminosity–chemical composition relationship of Nugis & Lamers (2000):

$$\log \dot{M} = -13.60 + 1.63 \log \frac{L}{L_\odot} + 2.22 \log Y$$

Equation 2.23 takes account of clumping and Crowther (2001) concludes that it agrees with the results of line-blanketed analyses. Nugis & Lamers (2000) also have an equation analogous for WC stars:

$$\log \dot{M} = -8.30 + 0.84 \log \frac{L}{L_\odot} + 2.04 \log Y + 1.04 \log Z$$

Since we defined constant values for $L$ and $Y$ for all WC subtypes, the WC relation of Nugis & Lamers (2000) was used to give a fixed mass–loss rate of $\log \dot{M} = -4.72 M_\odot \text{yr}^{-1}$ at solar metallicity. In the case of terminal velocities for both WN and WC stars, we used the $v_\infty$–spectral type calibrations of Prinja et al. (1990) as reference for all types. These were again smoothed to give a linear distribution.

As discussed in §2.1.2, there is not yet any clear evidence for the scaling of W-R winds with metallicity, indeed observational evidence for this effect in O star winds is sparse. It is usually assumed that no scaling occurs, since the effect of large amounts of CNO–cycle elements may contribute heavily to an outflow. Recently analysis of the CMFGEN models indicate that iron elements may be responsible for the driving of winds due to a re-evaluation of the importance of the CNO transitions at EUV wavelengths. This seems to indicate that a scaling is indeed applicable although no power-law exponent has yet been suggested.

Observational work has suggested the possibility of a metallicity dependence, such as the work of Hamann & Koesterke (2000) who looked at the parameters of a Galactic
and LMC W-R sample. However, due to large observational and modelling uncertainties and the limit in range of metallicity to a factor of ~ two, a quantitative value on the metallicity dependence was not determined. More recently, an analysis by Crowther et al. (2003b) found that a sample of LMC WC stars had derived wind densities which were ~0.2 dex lower than Galactic WC stars analysed by Dessart et al. (2000) using the same techniques. Other evidence from Guseva et al. (2000), investigating the nebular He II flux from 30 galaxies, suggests that only low metallicity environments \((Z < 0.2 Z_\odot)\) can produce He II \(\lambda 4686\). Since only W-R stars with weak winds can produce significant flux beyond the He\(^+\) ionizing edge, this suggests that some kind of metallicity scaling is being observed. Recent theoretical work in this area is briefly discussed in §2.1.2, and points towards a possible scaling.

As this power law dependency is extremely uncertain, the same values were adopted as used previously for the O star models in § 2.3 (Leitherer et al. 1992). The metallicity effect identified by Crowther et al. (2003b) is consistent with the scaling that we have used.

**Wind Density Effects**

As discussed in the case of the O stars, the emergent ionizing flux is extremely dependent on the mass-loss of a star. The same is applicable to W-R stars, although the results can be far more extreme. Fig. 2.14 shows the SEDs of three WC model atmospheres which are identical \((T_* = 100000\,\text{K})\) with the exception of the below distinctions. The first model shows the SED of a solar metallicity surface chemistry and a mass-loss rate of that expected at \(Z_\odot\). Model 2 is a model with both \(1/10\)th \(Z_\odot\) surface chemistry and mass-loss scaling. Model 3 has the same chemistry as model 2, but has no mass-loss scaling. What is immediately obvious is that chemistry makes a very small difference to the emergent flux, what is far more important is the density of the wind, which is far reduced in the case of model 2. This is due to the recombination that takes place in the dense wind of a W-R star. In a dense wind, the matter is able to cool much more quickly due to the effects of the CNO-cycle elements (Hillier 1989). In this environment, He \(++\) will recombine at some point, so is then able to effectively blanket the wind of the star there on out and blocks radiation with energies greater than 54.4 eV. This emergent flux dependence on mass-loss rate is discussed in Schmutz et al. (1992). Only the most extreme temperature W-R models at low metallicity emit significant flux above 54.4 eV.
Figure 2.14: The flux distribution of model WC#10 ($T_\star = 100\,000\,\text{K}$) for three different assumptions about the mass-loss scaling and abundances: (a) solar mass-loss rate and abundances (solid); (b) mass-loss rate and abundances scaled to $0.05\,Z_\odot$ (dashed); and (c) solar mass-loss and $0.05\,Z_\odot$ abundances (dotted). Only model (b) with wind parameters scaled to $0.05\,Z_\odot$ has a significant flux below the He$^+$ edge at $228\,\text{Å}$ because of the reduced wind density. The flux is in units of ergs cm$^{-2}$ s$^{-1}$ Å$^{-1}$ measured at the stellar surface.
Observations of nebular HeII $\lambda$ 4686 emission is also thought to be associated with low metallicity environments where one may expect to find W-R stars with low density winds (Garnett et al. 1991; Schaerer et al. 1999). The Nugis & Lamers (2000) mass–loss–luminosity–chemical composition relationship was used for this work, although it neglects weak lined stars that are associated with the low density winds that are transparent to the far UV. As a Galactic population was considered, the relationship implies that no W-R stars with weak winds are found in a Galactic environment. This is of course untrue, as a population of W-R stars with weak lines falls some $\sim 0.5$ dex below this calibration, with some extreme cases such as HD 104994, a WN3 being predicted by Crowther et al. (1995) to emit strongly in the He$^+$ continuum. Nebular HeII $\lambda$ 4686 has also been observed around W-R stars in the Galactic environment, such as the nebula associated with the WO star Sand 4 observed by Esteban et al. (1992). Stars like these must have weak winds for the reasons above, and are not included in the Nugis & Lamers (2000) relation, due to their use of radio continuum fluxes for mass-loss determination which is biased towards stars with strong mass-loss. This may not turn out to be problematic, however, as these stars are usually of extreme temperature and are very rare in a Galactic environment. For the case of a starburst, their contribution may be insignificant, since stars with strong winds will dominate in a large population at high metallicity. These stars may well be more important at low metallicity, but this is not a problem, as we have scaled mass-loss accordingly. These stars are not typical of our grid however, as only 18 of the 120 models exhibit this behaviour.

As our mass–loss rates are corrected for clumping, we must be consistent when using an atmosphere model. Fortunately, a clumping mechanism has been introduced into the CMFGEN model atmosphere calculations by adopting a volume filling factor $f$ of 0.1 (Hillier & Miller 1999). This has no effect on the emergent continuum flux but does make a difference to the shape of the W-R line profiles. It is possible to recover the unclumped mass–loss rates by multiplying the values in Tables 2.6 and 2.7 by $1/f^{1/2}$, i.e. $\sqrt{10}$. For all W-R star calculations, a $\beta = 1$ velocity law was assumed (see Equation 2.18).

Figures 2.15 and 2.16 show plots of $Q_0$, $Q_1$ and $Q_2$ as defined in §2.3 for all WN and WC models. $Q_2$ is particularly important as its value shows the transparency of the wind. In the case of WN stars, the winds in the 2.0 $Z_\odot$ models show no tendency to become transparent due to high wind density. In all these models, He$^{++}$ recombines and suppresses the flux below 228 Å. In the highest temperature $Z_\odot$ model, the wind becomes
Figure 2.15: The ionizing properties of the WN grid. Crosses represent 2.0 $Z_\odot$, open triangles represent $Z_\odot$, and filled pentagons, squares and triangles represent 0.4, 0.2 and 0.05 $Z_\odot$ respectively. All 60 models are plotted.
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Figure 2.16: The ionizing properties of the WC grid. Crosses represent $2.0 \, Z_\odot$, open triangles represent $Z_\odot$, and filled pentagons, squares and triangles represent $0.4$, $0.2$ and $0.05 \, Z_\odot$ respectively. All 60 models are plotted.
transparent as is shown by the large increase in $Q_2$ (a factor of $10^8$) between 100 and 120 kK. These extreme stars would contribute almost 100 percent of the ionizing flux below 228 Å of a starburst during certain epochs. The implementation and use of these models is discussed in full in Chapter 4. In the 0.4, 0.2 and 0.05 $Z_\odot$ models the wind becomes transparent at 100kK, 90kK and 80kK respectively. Like the O star models, the WN's $Q_0$ and $Q_1$ values start to converge, in this case at $T \geq 70kK$ ($10^{48.9}$ and $10^{48.2}$). As the wind becomes transparent, the $Q_2$ values should also converge towards a point, although at a much higher temperature.

In the WC grid one notices a large difference between low and high metallicity models at low temperature. At temperatures below 60kK and $Z$ exceeding 0.4 $Z_\odot$ the $Q_2$ and $Q_1$ values are several factors lower than the low $Z$ models. Below 504 Å, the difference can be as much as $10^8$ and below 228 Å this is as much as $10^{10}$, although even the lowest metallicities produce fluxes which are observationally insignificant (see Chapter 4). We again see a trend towards transparent winds at high temperatures, with the wind becoming transparent at 90kK at 0.05 $Z_\odot$, and 120kK at 0.4 and 0.2 $Z_\odot$. In this case 140kK is not hot enough to make the winds in the 2 and 1 $Z_\odot$ transparent.

### 2.5.1 The W-R Star Models

In order to make a meaningful comparison with the only existing W-R atmosphere grid, that of Schmutz et al. (1992), we must be careful, since the Schmutz et al. (1992) grid is parameterised by a very different set of values. The Schmutz et al. (1992) grid, which is implemented into Starburst99 is defined by three parameters: the "transformed radius", $R_T$, a measure of the inverse of the wind density (defined in Chapter 4, Equation 4.3) the effective temperature, $T_e$ and the beta law ($\beta = 1$ or 2). Since the temperatures of this grid range from 25,000K to 300,000K, a straight comparison would not be representative for either the new grid, or the Schmutz et al. (1992) grid. As both grids are intended for use in Starburst99, we compare the two grids by comparing models that are meant to represent the same evolutionary point on the H-R diagram.

The problem of assigning W-R models to evolutionary temperatures is explained in Chapter 4, which we refer the reader to. To make comparisons, we use Starburst99 to select the models, taking into consideration that Schmutz et al. (1992) conclude that a $\beta = 1$ velocity law should be used for $T_e \leq 90,000$K and $\beta = 2$ for higher temperatures.

In Fig. 2.17, we show the WN models that would be chosen in Starburst99 and
Figure 2.17: The emergent fluxes for CMFGEN WN models with $T_\ast = 45, 60$ and 90 kK for $Z = Z_\odot$ (black) and 0.2 $Z_\odot$ (blue). For comparison, the Schmutz et al. (1992) models are also plotted (red) that would be chosen in Starburst99 for the same evolutionary point at 4 Myr in the HRD but corresponding to the hydrostatic core Meynet et al. (1994) temperatures of 56.2, 82.2 and 130.1 kK. The transformed radii $R_t$ of these three models are 17.9, 8.4 and $1.7 R_\odot$. 
Figure 2.18: The emergent fluxes for CMFGEN WC models with $T_\ast = 60$ and 100 kK for $Z = Z_\odot$ (black) and 0.2 $Z_\odot$ (blue). For comparison, the Schmutz et al. (1992) models are also plotted (red) that would be chosen in Starburst99 for the same evolutionary point at 5 Myr in the HRD but corresponding to the hydrostatic Meynet et al. (1994) temperatures of 78.2 and 126.1 kK. The transformed radii $R_t$ of these two models are 2.7 and 1.6 $R_\odot$. 
our updated version at an age of 4 Myr for evolutionary points corresponding to \( T_\star = 45\,000, 60\,000\) and \( 90\,000\) K, for the CMFGEN models and temperatures of 56.2, 82.2 and 130.1 kK for the Schmutz et al. (1992) models. This is due to our choice of comparing the models at temperatures corresponding to \( T_{\text{hyd}} \) for the Schmutz et al. (1992) models and \( T_\star \) (the interpolated value) for the CMFGEN models. Two CMFGEN models correspond to solar and \( 0.2 \times \) solar metallicity. For a temperature of 45 000K, we see that a solar CMFGEN model has a large flux deficit compared to the Schmutz et al. (1992) model. The difference is as large as a factor of \( 10^7 \) just long-wards of the 228 Å edge. This is due to the line blanketing effect which is most important at relatively low temperatures. The solar CMFGEN flux distributions do find far better agreement at 60 000K due to the diminishing importance of line blanketing. This agrees with the finding of Esteban et al. (1993) from photoionization modelling that the Schmutz et al. (1992) models are incorrect at the cool temperature end because of the lack of line blanketing deep in the stellar wind. (§2.2). We can see that for both the 45 000 and 60 000K cases, the low metallicity models converge as expected with the Schmutz et al. (1992) due to the lower wind density. At 90 000K an extremely large difference can be seen between the solar metallicity CMFGEN and the Schmutz et al. (1992) model. This is because a \( \beta = 2 \) model was chosen and the wind has become transparent in the He\,\text{II} continuum. The 0.2 \( Z_\odot \) CMFGEN model also exhibits this behaviour due to the lowering of wind density as a consequence of our adopted metallicity scaling.

In Fig. 2.18, we plot two WC models corresponding to an age of 5 Myr and temperatures of \( T_\star = 60\,000 \) and 100 000 K. For comparison we show the two Schmutz et al. (1992) models of equivalent temperature and radius. The Schmutz et al. (1992) models have temperatures of 78.2 and 126.1 kK. The differences seen for the cooler case are similar to the 45 000K WN model comparison in that the line blanketing at \( Z_\odot \) produces a significant reduction in the flux in the He\,\text{II} continuum. The differences seen at 100 000K can be explained by the wind density determining the flux below 228 Å.

In the next Chapter we explain how these two new grids are applied to the population synthesis code Starburst99.
Chapter 3

Single Star Comparisons

It is extremely important to test the predictions of the new model atmospheres against observations of real stars to ensure their suitability. The far UV is an important region to use as a test since it is dominated by the highest energies where the use of the correct input physics is essential. We cannot however, observe stars in the far UV, due to interstellar extinction from dust and absorption by interstellar hydrogen atoms, so indirect methods of measurement must be sought.

In theory, it is possible to use the emission lines from nebulae associated with single stars as a tracer of the far UV continuum. This is ideal for O stars since many H II regions are known to be associated with them, but nebulae with W-R central stars are far less common. In the first section we deal with the O star sample, using the data of Bresolin et al. (1999) for comparison. §3.2 deals with the detailed comparisons of two individual W-R star models, one extremely hot model, and the other a cool heavily line–blanketed late WN star.

3.1 The CLOUDY Photoionization Code

3.1.1 The Nebular Problem

Modelling the nebulae around O stars has been progressing for a number of years. Using the photoionization model described by Osterbrock (1974), there have been many attempts to reproduce the fine structure line strengths observed in H II regions. The first studies dealing with emission lines associated with H II regions relied on modelling using a single star as a source of photoionization (Shields 1974; Shields 1978; Stasińska 1978). These
models were simple by today’s standards, but nonetheless, densities, temperatures and metallicities of the surrounding nebula could be estimated (e.g. McCall et al. 1985). For our comparisons of nebular line data, we use the photoionization code CLOUDY. This code written by Gary Ferland can model a number of diverse astrophysical environments. There are many situations that can be encountered in which a low density gas is heated by a central ionizing source resulting in the production of the fine structure nebular lines. Under these circumstances it is possible to build a model of the gas which is able to reproduce the intensities and relative ratios of these lines. This can be done by solving the thermal and statistical equilibrium equations simultaneously with those that describe the ionization-neutralization and heating-cooling processes. Osterbrock (1988) and Aller (1984) provide the formulation of basic physical processes to describe the nebula. The workings of the photoionization code CLOUDY are described at length in Ferland (2002), and are compared with other existing codes in the earlier work of Ferland (1995).

CLOUDY is a versatile photoionization code that has been designed to model emission line regions from those found in low density cold, intergalactic medium environments to those of Broad Absorption Line Quasars. CLOUDY has been confirmed to be able to model a range of gas temperatures from 10 to $10^9$ K and densities from $10^{-8}$ cm$^{-3}$ in interstellar medium environments to extremely high densities of up to $10^{13}$ cm$^{-3}$. Since H II regions generally have temperatures in the region of 10kK and densities of 1000 cm$^{-3}$, CLOUDY is suitable for our purpose.

3.1.2 Running CLOUDY for H II Region Purposes

Photoionization analysis can be extremely powerful due to the large number of observable emission lines that can be modelled via the input of a small number of parameters. CLOUDY, for example has the ability to predict the intensities of $10^6$ different emission lines from the handful of different inputs that are listed below.

Continuum Shape

This is specified by the inclusion of a SED from a stellar atmosphere model or a Starburst (SB99) continuum. CLOUDY uses $F_\nu$ (flux in frequency units) versus energy in Rydbergs, so manipulation of our model grids is necessary to obtain a valid result. In our case, this was done via the use of a C code written specifically for the job.
CHAPTER 3. SINGLE STAR COMPARISONS

Continuum Luminosity

The luminosity of a system is normally set by the observations that are to be modelled, since we have an absolute measurement of flux from either photometry or spectroscopy. The input continuum shape is scaled by the specified luminosity.

Chemical Composition

The chemical composition is set by observation, usually via the nebular $\frac{N(O)}{N(H)}$ ratio. If desired, depletions due to dust condensation can be taken into account. Dust itself can also be added.

Geometry

Since CLOUDY is essentially a 1-dimensional code, a spherical geometry is always assumed defined by an inner and outer radius. This can be made effectively plane-parallel by assuming a larger inner radius with a thin wall. The geometry also includes the definition of filling factor and covering factor. The filling factor defines how much of the actual sphere is filled with gas (see Fig. 3.1), since nebulae are known to have inhomogeneous density distributions. The covering factor defines the amount of area on the surface of the sphere that the nebula takes up. The hydrogen density can also be defined. This geometry is a reasonable approximation of many nebulae, although care must be taken to select the most spherically symmetrical cases. Non-spherical cases may not be accurately represented by CLOUDY.

3.2 O Star Comparisons

WM-basic has been tested substantially against optical and UV spectra (above 1200 Å) with excellent results (e.g. Pauldrach et al. (2001), see Fig 2.6, Chapter 2). A more important test for this work is to probe the far UV ionizing continuum using a photoionization analysis to reproduce the nebular line strengths associated with single O stars. Here we compare the new grid of WM-basic O star models with those of CoStar (Schaerer & de Koter 1997) and the LTE (Lejeune et al. 1997) models. We then use the photoionization code CLOUDY (see §3.1 for a description) to predict the nebular line strengths that would be produced by each model’s stellar continuum. We generated models of single star nebula for five O star temperatures (35, 37, 40, 45, 50kK) in the range in which we would
Figure 3.1: A 2-D representation of the CLOUDY assumed spherical geometry. This spherical shell geometry includes dust grains and a filling factor indicated by the clumps of grey ‘gas’ and is a reasonable approximation for most nebulae.
CHAPTER 3. SINGLE STAR COMPARISONS

expect sufficient ionization to produce significant nebular emission line intensities. Five metallicities were considered for the WM-basic and Kurucz (1979) models (0.05, 0.2, 0.4, 1.0 and 2.0 $Z_{\odot}$) and two metallicities for the CoStar models (0.2 and 1.0 $Z_{\odot}$) since we were limited by the model's availability. To simplify the photoionization model we assumed an ionization parameter of $\log U = -3$ and scaled metallicity with that of the atmospheres. U is defined as

$$U = k(Qe^2n_e)^{1/3}$$

(3.1)

where Q is the rate of production of Lyman continuum photons, $e$ is the nebular filling factor, $n_e$ is the electron density and $k$ is a constant. We assumed a hydrogen density of 50 cm$^{-3}$.

H II region studies such as Shields (1978) and Stasińska (1978) point towards the effect of softer ionizing continua at higher metallicity. This relationship was first suggested by Shields & Tinsley (1976) and more recently, the works of Bresolin et al. (1999) and Kennicutt et al. (2000) have found a relationship between an effective temperature decrease and a rise in metallicity. These works discuss the possibility of a lowering of the upper mass cutoff of the IMF with metallicity, although caution is given, as the Kurucz (1979) models that are used in the Bresolin et al. (1999) may have insufficient line blanketing, leading to an ionizing continuum that is too hard.

In Figure 3.2 we plot the predicted He I $\lambda 5876$/H$\beta$ line ratio for the three grids of models over the temperature range defined above, although we omit the 50kK values for clarity since they tend to converge with the 45kK points. We plot the Bresolin et al. (1999) data with determined metallicity over the models as a comparison. These data are predominantly of giant H II regions rather than regions ionized by single stars. However, single star models are applicable when modelling these regions as they represent tens of stars rather than the SB99's representation of $\sim 1000$ or more O type stars. A more realistic way to model these regions would be to use a Monte-Carlo simulation to deal with the small number of O stars that are the ionization source.

As a large proportion of the Bresolin et al. (1999) data is solar or super solar, we can only hint at the suitability of the CoStar models. As expected the non-LTE models predict more He I at high temperature and high metallicity. Our models show that line blanketing has a new effect, not predicted by the Kurucz (1979) models at high metallicity.
Figure 3.2: Comparison of Bresolin et al. (1999) H II region data plotted with several nebula models. The black lines represent single WM-basic models, red represents LTE Kurucz models and blue represents the CoStar models of Schaerer et al. The temperatures are represented by different shaped symbols; 37kK has triangular symbols; 40 kK has square symbol and the 45 kK has pentagonal symbols. The data is presented by a symbol with a star shaped outline (not filled).
The increased effect of line blanketing, enhanced due to the scaling of wind density with metallicity, lowers the \( \text{He I} \lambda 5876 / \text{H} \beta \) flux in stars with effective temperatures of less than 40kK. This effect could be mis-interpreted as a lowering of the effective temperature scale at high metallicity. We see a sharp downward trend in the data, with a lowering of He I intensity with metallicity. This sharp trend may be partly due to line blanketing although the \( \text{He I} \lambda 5876 / \text{H} \beta \) flux falls off so quickly at high metallicity that it may not be wholly attributed to just line blanketing. As suggested by Bresolin et al. (1999) and Kennicutt et al. (2000) this trend could be due to a combination of line-blanketing and other effects such as a lowering of the upper mass cutoff of the IMF. Models with temperatures of 35kK and below exhibit no detectable He I flux.

Next we use the whole Bresolin et al. (1999) dataset (from 2.0 \( Z_\odot \) to 0.2 \( Z_\odot \)) in a \( R_{2,3} \) vs \( \eta' \) comparison with the new models (WM-basic, O star models) and Kurucz (1979) single star models. Costar models are omitted from this comparison due to the limited parameter coverage available for comparison (just 2 metallicities 1 and 0.2 \( Z_\odot \)). These two quantities are used extensively in Chapter 5, therefore the reader is directed to Chapter 5 §5.1 for a discussion of their qualities. \( R_{2,3} \) is the sum of the \([\text{O II}]\lambda 3727\) and the \([\text{O III}]\lambda \lambda 4959, 5007\) nebular line fluxes relative to \( \text{H} \beta \) and is used as a rough metallicity diagnostic. \( \eta' \) is a very useful parameter, since it can be used to probe the relative “softness” of the ionizing source of a nebula. This parameter is defined as \( \eta' = \frac{[\text{O III}] / [\text{O II}]}{[\text{S III}] / [\text{S II}]} \) since these lines are always present in the optical nebular spectrum. Due to the nature of \( \eta \), a soft continuum will have a higher value than a hard one. We present two grids of single star models, the WM-basic and Kurucz (1979) for 5 metallicities (0.05, 0.2, 0.4, 1.0 and 2.0 \( Z_\odot \)) for 4 temperatures (35, 37, 40 and 45kK) plotted with the full Bresolin et al. (1999) dataset in Figure 3.3. Models of differing metallicity, but constant temperature are joined with solid lines. From this figure, the superiority of the WM-basic non-LTE models can clearly be seen. On the top plot, the models ranging from 0.2 to 2.0 \( Z_\odot \), cover the data almost exactly. The 0.05 \( Z_\odot \) models seem to have \( \eta' \) values that are too low, since their ionizing spectra are too hard to reproduce the data. This is due to the lack of very low metallicity \( \text{H II} \) regions represented in the Bresolin et al. (1999) sample. The \( R_{2,3} \) value turns around at low metallicity due to the lowering of the relative intensity the two oxygen lines. In the case of the 2.0 \( Z_\odot \) models (the furthest left on both plots), the \( R_{2,3} \) value is low because of the cooler nature of the nebula. Oxygen exists in a neutral state and \([\text{O I}]\) emission lines are not accounted for in the \( R_{2,3} \) value. The Kurucz (1979) models
Figure 3.3: An \( R_{2,3} \) vs \( \eta' \) comparison of Bresolin et al. (1999) data versus the model predictions of the WM-basic and Kurucz atmosphere models for various temperatures. Lines join models of constant temperature but varying metallicity (0.05, 0.2, 0.4, 1.0 and 2.0 \( Z_\odot \)). The black lines represent single WM-basic models, red represents LTE Kurucz models. The temperatures are represented by different shaped symbols; 35kK has triangular symbols; 37 kK has square symbols; 40 kK has pentagonal symbols and 45kK has hexagonal symbols. The data is presented by a symbol with a hexagonal outline (not filled).
Table 3.1: The parameters of WR124 predicted by a combination of either ISA-wind or CMFGEN with the photoionization code CLOUDY. This table is taken from Crowther et al. (1999) and represents the work therein. $\epsilon$ represents the nebular filling factor.

<table>
<thead>
<tr>
<th>Quantity</th>
<th>Observation</th>
<th>CMFGEN</th>
<th>ISA-wind</th>
</tr>
</thead>
<tbody>
<tr>
<td>$\lambda$ 5876 He I</td>
<td>1.6±0.3</td>
<td>0.0</td>
<td>31.0</td>
</tr>
<tr>
<td>$N^{2+}/N^+$</td>
<td>$\ldots$</td>
<td>0.00</td>
<td>3.0</td>
</tr>
<tr>
<td>$O^{2+}/O^+$</td>
<td>$\leq0.02$</td>
<td>0.00</td>
<td>1.5</td>
</tr>
<tr>
<td>$S^{2+}/S^+$</td>
<td>0.9</td>
<td>1.2</td>
<td>17.0</td>
</tr>
<tr>
<td>$Ar^{2+}/Ar^+$</td>
<td>$\ldots$</td>
<td>0.00</td>
<td>25.0</td>
</tr>
<tr>
<td>$T_e$(K)</td>
<td>6200</td>
<td>6900</td>
<td>8500</td>
</tr>
</tbody>
</table>

are clearly too soft at high metallicity and lower temperature, compared to the non-LTE models which predict the correct $\eta'$ values for cooler O stars. In the LTE case, only the low metallicity case give a hard enough ionizing flux to model the data.

3.3 W-R star comparisions

3.3.1 WR124, a cool W-R star

CMFGEN has been extensively used to model W-R stars in the optical part of the spectrum. Examples of works that produce good model fits to W-R stars are Crowther et al. (2002); De Marco et al. (2000) and Hillier & Miller (1999). Nebulae associated with W-R stars are rare and as a result are often modelled individually since their parameters are poorly defined. W-R model atmospheres are complex, with high density winds, so it is important to test the UV continuum with the use of photoionization models. The Schmutz et al. (1992) models were tested via this method by Esteban et al. (1993), where 8 W-R nebulae were compared. They found that these models had too hard an ionizing continuum to reproduce the nebular emission line intensities for the lower temperature W-R stars, although the hotter W-R stars gave reasonable agreement. Crowther et al. (1999) used
CHAPTER 3. SINGLE STAR COMPARISONS

the cmfgen and ISA-wind codes of Hillier & Miller (1998) and de Koter et al. (1997) in conjunction with cloudy to model the nebula around WR 124, a cool WN8h star with an effective temperature of 33kK. The optical stellar spectral lines produced with both codes fit the observations equally well. Photoionization modelling from this work, however, tells a different story. Table 3.1 shows that cmfgen reproduces the important nebular emission line ratios far better than the ISA-wind model of de Koter et al. (1997). All ratios presented in the Table are over predicted by this model, suggesting that the ionizing continuum is too hard. The cmfgen model, however, predicts a far softer ionizing continuum, and is in agreement with the observations for the N^2+/N^+, O^2+/O^+, S^2+/S^+ and Ar^2+/Ar^+ ratios. It seems that the cmfgen model is far superior to that of ISA-wind for the prediction of the level of the far-UV continuum, for cool W-R stars. This point is crucial for the results of this thesis.

3.3.2 Br2, a hot W-R star

![Image](image.png)

Figure 3.4: Observation of Br2 through Hα filter. This picture is from the work of Chu et al. (1999), taken with the 0.9m telescope at the Cerro Tololo Inter-American Observatory (CTIO).

In general the ISA-wind models have been found to model the UV continua of hot stars with reasonable accuracy. The reason for this is probably due the lesser importance
Table 3.2: The parameters of Br2 predicted by a CMFGEN, CLOUDY model, showing a best fit for an LMC metallicity WN model at 160kK

<table>
<thead>
<tr>
<th>Quantity</th>
<th>Observation</th>
<th>CMFGEN fit # 1</th>
<th>CMFGEN fit # 2</th>
</tr>
</thead>
<tbody>
<tr>
<td>$T_\nu$(kK)</td>
<td>\ldots</td>
<td>160.0</td>
<td>160.0</td>
</tr>
<tr>
<td>$\dot{M}$(10^{-6}$\text{yr}^{-1}$)</td>
<td>\ldots</td>
<td>8.0</td>
<td>9.3</td>
</tr>
<tr>
<td>$\epsilon$</td>
<td>\ldots</td>
<td>0.22</td>
<td>0.036</td>
</tr>
<tr>
<td>$T_e$(K)</td>
<td>14000.0</td>
<td>12200.0</td>
<td>11000.0</td>
</tr>
<tr>
<td>$\lambda$ 4471 He i</td>
<td>1.6\pm1.1</td>
<td>2.10</td>
<td>2.5</td>
</tr>
<tr>
<td>$\lambda$ 4686 He ii</td>
<td>59.0\pm4.0</td>
<td>83.0</td>
<td>55.0</td>
</tr>
<tr>
<td>O^{2+}/O^{+}</td>
<td>7.4 \pm1.0</td>
<td>16.6</td>
<td>11.3</td>
</tr>
<tr>
<td>$\lambda$ 6716 [S ii]</td>
<td>18.0\pm0.1</td>
<td>17.7</td>
<td>19.0</td>
</tr>
<tr>
<td>$\lambda$ 6731 [S ii]</td>
<td>12.0\pm0.1</td>
<td>12.9</td>
<td>13.8</td>
</tr>
</tbody>
</table>

of line-blanketing at high temperatures, where recombination is supressed in the wind of these stars. It is important to test the predictive ability of CMFGEN in the same way, so that we can determine whether it is applicable to the whole range of W-R temperatures. Here we look at the extreme WN2 star Br2, which is associated with the nebula DEM L6 in the LMC (Davies et al. 1976) at metallicity $Z = 0.4 \text{ Z}_\odot$ (see Fig. 3.4). To find the best W-R model for the ionization source we used the optical observations of Crowther (September 2000) from the VLT FORS2 instrument. Figure 3.5 shows the best fit CMFGEN model to the WN2 star. The best fit model has a luminosity of 170,000$L_\odot$, a radius of 0.54$R_\odot$, and a H/He abundance of 0.001 by number. The wind parameters that best fit these observations suggest that the mass loss rate is $8.0 \times 10^{-6} M_\odot\text{yr}^{-1}$ and a $v_\infty$ of 1700 km s$^{-1}$. The mass loss of this star is comparable to that of the hottest 0.4 $Z_\odot$ star in our WN grid. Table 3.2 shows the results of the CLOUDY modelling. This was done by using a spherical geometry with a 40 percent covering factor with an inner radius of 4.0 parsecs and outer radius of 4.4 parsecs. The inner radius was fixed via the observation of Chu et al. (1999) (Fig. 3.4). We set the thickness of the nebular with reference to Esteban et al. (1993) who quote that nebulae of this type have a shell thickness $\Delta R \sim 0.1R$. The CMFGEN model was added as a source continuum and the nebular chemical abundance was set at 0.4 $Z_\odot$. The nebular line intensities of Garnett & Chu (1994)
Figure 3.5: Optical spectra of Br2 from the VLT FORS2 instrument plotted with the best fit CMFGEN model (blue)
were used to give a best fit with CLOUDY's iterative SUPLEX method which scans over a variable parameter range. For this process we allowed variable luminosity, filling factor and hydrogen density. Table 3.2 shows the photoionization modelling from the two models which give good fits to the stellar spectra. One model over estimates the mass-loss and the other is an underestimation. The level of mass-loss is crucial for these models as the flux below 228 Å in the higher mass-loss level is approximately 30 percent of the low mass-loss model. Both these fits give reasonable agreement with the observations. The low mass-loss fit gives a higher nebular temperature, but the high mass-loss model gives a superior fit to the helium and sulphur lines and the O$^{2+}$/O$^+$ ratio. Both fits over predict the line strengths of the oxygen lines by a factor of $\sim 1.5$, although we can conclude that the He II $\lambda 4686$ and [S II] lines can be correctly predicted by these models. Although we see a disagreement in the oxygen ionization balance, we get a reasonable result considering the complex morphology of the system, which cannot entirely be approximated by a 1-D spherically-symmetric geometry used in CLOUDY.
Chapter 4

Evolutionary Synthesis Codes

Evolutionary synthesis codes can be used to model stellar populations to gain a better understanding of various defining parameters such as chemistry, age and stellar content. Clusters, associations and starbursts can all be modelled in this manner. This work is concerned with the environments of newly-formed starburst regions and the properties of their young stellar content. Many of these parameters cannot be observed directly, but can be inferred via the modelling of observable starburst properties from the satellite UV and optical regions of the spectrum. Stellar wind features from the dominant massive star population can be synthesized to provide information on star formation rates, the slope of the IMF and the age of the starburst. In the optical regime, the far UV output of the dominant population creates an ionized nebula around the starburst. The emission lines formed in the nebula are very sensitive to many of the system’s properties including metallicity of the starburst environment, the geometry and dust content of the nebula, and the shape of the starburst’s integrated continuum in the UV (Robert et al. (1993); Leitherer et al. (1995) and de Mello et al. (2000)). The modelling of a nebula surrounding a starburst can be achieved with the combined use of a stand alone photoionization code and integrated starburst model continuum. However, the predictive ability of this method relies heavily on the accuracy of the evolutionary and atmospheric models used in the synthesis code. Nebular geometry can also be problematic when modelling a system with the use of an external photoionization model, which classically uses spherically symmetric or plane-parallel schemes, although this problem can be reduced via clever selection of observational targets. The accuracy of the stellar atmospheres, especially for the young, massive population depends greatly on the physics used to describe them (Chapter 2).
Another source of uncertainty is the current understanding of the evolution of massive stars. Although there has been much recent advancement in understanding of the evolution of massive stars, predictions can still match poorly with observations due to the complexity of these systems (e.g. H-R diagram positions).

Early attempts to model stellar populations within starbursts relied mainly on the Kurucz (1992) LTE atmospheres (e.g. the work of García-Vargas et al. (1995)) for the entire population of stars. It is well known that the LTE approximation breaks down for the most massive stars, by several orders of magnitude, as referred to in Chapter 2. Later evolutionary synthesis models used simple non-LTE models to replace the most extreme examples of the population, the Wolf-Rayet stars. One of the most well known examples of this was the Leitherer et al. (1999) model (Starburst99; SB99), which used the hydrogen-helium non-line blanketed Wolf-Rayet models of Schmutz et al. (1992). Line blanketed non-LTE O star models have since been added to a population synthesis model similar to that of SB99 by Schaerer & de Koter (1997).

Table 4.1: Contemporary population synthesis models

<table>
<thead>
<tr>
<th>group</th>
<th>authors</th>
<th>formation channel</th>
<th>evolution</th>
<th>atmospheres</th>
</tr>
</thead>
<tbody>
<tr>
<td>Baltimore</td>
<td>Leitherer et al</td>
<td>single ± 0 % binary</td>
<td>Geneva</td>
<td>Lejeune et al. (1997)</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>Schmutz et al. (1992)</td>
</tr>
<tr>
<td>Madrid</td>
<td>Cerviño &amp; Mas-Hesse</td>
<td>single ± 30 % binary</td>
<td>Geneva</td>
<td>Kurucz (1992)</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>Brussels</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>Mihalas (1992)</td>
</tr>
<tr>
<td>Quebec</td>
<td>Dionne &amp; Robert</td>
<td>single ± 50 % binary</td>
<td>Geneva</td>
<td>Kurucz (1992)</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>Brussels</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>Schmutz et al. (1992)</td>
</tr>
<tr>
<td>Toulouse</td>
<td>Schaerer &amp; Vacca</td>
<td>single ± 20 % binary</td>
<td>Geneva</td>
<td>Kurucz (1992)</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>Brussels</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>de Koter (1997)</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>Schmutz et al. (1992)</td>
</tr>
</tbody>
</table>

Several competing evolutionary synthesis models now exist, with the differences mostly arising from the evolutionary tracks that they use. Table 4.1 presents the four competing
population synthesis codes reproduced from Leitherer (1999). Although all use the single star evolutionary tracks of Meynet et al. (1994), three of the models choose to include some allowance for binary evolution. The presence of a binary partner can significantly alter the evolutionary cycle of both components of the system. A close binary may experience Roche Lobe overflow from the most massive component to the lower mass star. This may prolong its lifetime and strip the more massive star of its upper atmosphere forcing the star into the W-R phase, even if the star is not massive enough to produce a W-R star in a single system. In a starbursting galaxy, this could have the effect of lowering the mass limit for W-R formation as suspected in many low metallicity environments. In the extreme circumstances found in large stellar populations this effect will also prolong the O star contribution to the starburst due to the mass gained by the lower mass components of some massive binaries.

The model we use in this thesis however, is the Baltimore group’s Starburst99 code. This does not include binary evolution for the following reasons. The main problem with the inclusion of binary evolution is that the binary population is observationally undefined within large stellar populations. Although we expect to find multiple systems in regions of high stellar density, no value has been universally accepted due to a lack of real evidence as these populations remain unresolved by today’s telescopes. This introduces one extra free parameter and two major uncertainties. We need to define the fraction of systems that belong in binaries in a large stellar population, with the only certainty being that the binary fraction lies somewhere between 0 and 1. Two main uncertainties are also introduced in a proper inclusion of binary tracks. Firstly we need some kind of mass spectrum to represent the relative and absolute masses of both binary components. To do this we must introduce some kind of IMF-like function to represent the statistical ensemble. Coupled with this, some function must be defined to represent the distribution of separations between the binary components. Systems that are far apart will not have evolutionary implications on each other, whereas close binaries may be composed of two nuclear burning cores with a common atmospheric envelope. We choose SB99 as the model from which to work, hence binary evolution is ignored hereafter as the main aim of this work is to introduce new stellar atmospheres.

This chapter outlines the structure of an evolutionary synthesis model, with all its components, then introduces the methods for incorporating the new stellar atmospheres and finally compares them to other recent works. In §4.1, the evolutionary tracks of
Meynet et al. (1994) are reviewed, with their application to the HRD and the problems associated with this set of models are discussed. The latest models including rotation are introduced. In §4.2, the part of the Starburst99 model relating to spectral synthesis is described in detail. §4.3 discusses the implementation of the two new grids into the starburst99 model. §4.4 compares the results of the new model with other contemporary work. §4.5 compares the nebular He II predictions of the new model and SB99 with the CoStar O star atmospheres. Finally in §4.6 other outputs relating to the new atmosphere grids are described, such as the mass return and feedback of the new starburst model.

4.1 Evolutionary Tracks

It has been over 60 years since an inhomogeneous model of stellar evolution was proposed by Schönberg & Chandrasekhar (1942). By the mid 1980's advances in computing were sufficient to permit the complex time evolution solution of unified stellar models, allowing for predictions of stellar lifetimes, luminosities and even surface abundances of the chemical products of nuclear fusion. These predictions have a wide range of applications in astrophysics, from the studies of star clusters and larger populations through evolutionary synthesis to the understanding of the chemical evolution of individual stars. The evolutionary tracks that are used in SB99 are those of Meynet et al. (1994), a set of models first described in Schaller et al. (1992) based on the earlier works of Maeder (e.g. Maeder (1983); Maeder & Meynet (1987)).

4.1.1 Physical Ingredients

These models are extremely complex, having not only to explain the dynamical pressure and gravity balances in the stellar core, with accurate description of the nuclear processes therein, but the structure of the star to its photosphere, since convection, radiative outflow and chemical mixing all perturb the inner processes. Like all complex models the assumptions included in the models come from a wide and varied range of scientific disciplines and are described in Schaller et al. (1992). We present a brief description of the main ingredients of the model below.
Opacity

The inclusion of opacity in an evolutionary model is as important as its use in a stellar atmosphere model, since it is coupled with the temperature structure of the stellar interior and restricts the flow of radiation. Older evolutionary models suffer from inaccurate or poorly defined atomic opacity determinations. Until 1991, the Los Alamos opacity libraries (1977) were used as the standard source for the definition of opacity in stellar interiors (Maeder & Meynet 1989). It was known that these libraries were unsatisfactory, and were incorrectly influencing the resulting evolutionary predictions. The problem of opacity was mainly solved by the OPAL project of Iglesias et al. (1987) and OPACITY project of Seaton (1987), where a complete description of all abundant elements was calculated independently. The works of Iglesias et al. (1987) were used in the new grids of the Meynet et al. (1994).

Abundances

Abundances effect the entire evolution of a star in many ways. The reaction rates of the CNO cycle depend on the amount of CNO catalysts in the core, and heavier elements restrict the outflow of radiation due to the effects of opacity. Since the abundance is so crucial to the opacity the Meynet et al. (1994) grids use those provided in the opacity tables of Iglesias et al. (1987).

Nuclear Reactions

Hydrogen, helium and carbon burning phases of stellar evolution are accounted for. For the hydrogen burning phase, both CNO and pp chains are used. The reaction networks used in these processes are described in Maeder (1983). The helium burning network, which follows the hydrogen burning phase is described in Maeder (1987). The carbon burning reactions are taken from Caughlan & Fowler (1988). Since the structure of the model depends on its chemical content, and the chemical content depends on the structure of the model, the model must be iterated to find the correct solution at every time step.

Convection Parameters

The amount of convection in a star is very important for the mixing of H into the nuclear burning core. This enrichment will deliver fresh fuel from the outer envelope of the star
to the core to prolong the lifetime. In massive stars, a phenomenon called overshooting is used. This assumes a small convective layer that occurs close to the stellar core, and causes hydrogen enrichment from the envelope to the nuclear burning core, thus extending the MS lifetime of the star. For instance a $15M_\odot$ stars MS lifetime is prolonged by 18 percent by core overshooting (Lamers & Cassinelli 1999). The Meynet et al. (1994) models are calculated with an overshooting parameter of $d/H_p = 0.2$, where $d$ is the overshooting height and $H_p$ is the pressure scale height. This parameter defines the amount of material in the envelope that can be exchanged via convection.

**Mass–Loss Rates**

Originally, the mass loss rates were set empirically by reference to the work of de Jager et al. (1988) for the solar metallicity track, and a scaling of $\dot{M} \propto Z_\odot^{0.5}$ for all non W-R stars in the other tracks. In order to match the observational lifetimes and numbers of W-R stars, however, the empirical mass–loss rates were enhanced by a factor of two, allowing for a lower initial W-R mass, and echoing the observed suppression of red-ward movement of the most massive stars. This enhancement is applied to all stars apart from the latter stages of W-R evolution; WNE, WC and WO stars, which are not expected to be affected by mass–loss in evolutionary terms in the same way, since their atmospheres are enriched by the products of hydrogen and helium burning (see Chapter 2, §2.1.2).

**Extended Atmospheres**

The standard $T_{\text{eff}}$ in these grids correspond to the hydrostatic core, but it is well known that W-R winds are optically thick. To correct for the non-negligible optical thickness of a W-R wind, an extended atmosphere must be added to simulate the appearance of these stars due to the interaction between radiation and matter in the outflow. The Meynet et al. (1994) models use a scheme modified from the theory of Castor et al. (1975) (see Chapter 2). From this formalism the optical depth can be calculated ($\tau(r) \equiv \tau_c + \tau_{\text{lines}}$; the expressions for $\tau_c$ and $\tau_{\text{lines}}$ can be found in Baschek et al. (1991) and Schaller et al. (1992) respectively). The effective temperature is then calculated using the Stefan–Boltzmann law at an optical depth which corresponds to $\tau(R_{\text{eff}}) = 2/3$. The uncorrected “core” temperature is also given in the evolutionary tracks. The effect of these temperature definitions is discussed later on in this Chapter (see §4.3), as it has important implications for the coupling of the atmospheric models to the evolutionary tracks in the SB99 models.
4.1.2 Evolutionary Implications For Massive Stars

The high mass-loss rates adopted by the Meynet et al. (1994) work have several important implications for the evolution of massive stars. The large mass-loss rates experienced in the main sequence "peels" off the upper atmosphere of high mass stars, allowing for single massive stars to enter the W-R phase. At $1.0 \, Z_{\odot}$ stars more massive than $40 M_{\odot}$ enter the W-R phase. Figure 4.1 shows the predicted HR diagram from the Meynet et al. (1994) tracks. This figure shows mass loss is high enough for stars above $85 M_{\odot}$ to suppress their red-ward movement in the HR diagram and move into the W-R phase without becoming Red Supergiants (RSG). Stars with initial masses of 40 and $60 M_{\odot}$ undergo a RSG phase, while stars on the $25 M_{\odot}$ track finish their lives as RSGs. We see that W-R stars are typified on the HR diagram as low luminosity stars with higher effective temperatures than the O stars. The temperatures plotted in Figure 4.1 are those defined at $\tau(R_{\text{eff}}) = 2/3$, and may therefore be cooler than observable W-R temperatures, or indeed theoretical calculations at a more realistic optical depth.

The effect of the chemical composition on the evolution of a star is large, and thus we can expect significant changes as the metallicity is lowered. Evolutionary tracks for $0.2 \, Z_{\odot}$ are shown in Figure 4.2. Several major differences can be seen. The first obvious difference is that a star must now be more massive to evolve to a W-R star. This is purely due to the lower level of mass-loss as determined by the adopted metallicity scaling in the O star phase. A star of initial mass less than $60 M_{\odot}$ will now not become a W-R star. Less obvious is the fact that lower metallicity stars now move further to the red. Stars on the $85 M_{\odot}$ track become yellow supergiants shortly before entering the W-R phase. Even at $120 M_{\odot}$, the evolutionary track moves into a more red-ward position before becoming a W-R star. This is counter intuitive, as it is well known that a star of constant mass will spend its life in the blue region of an HR diagram (Meynet et al. 1994). It is through the process of mass-loss that a star becomes a RSG. In the case of extreme mass-loss, however, the atmosphere is completely peeled away, leaving the hotter lower atmosphere and core to drive a powerful wind. Table 4.2 shows the time that massive stars spend in the W-R phase for high and low metallicities ($1$ and $0.2 \, Z_{\odot}$). This table clearly shows that higher metallicity stars spend much more time in W-R phase, even the though their nuclear reaction rates should be higher due to a greater number of catalysts in the nuclear burning core in the form of heavier elements. In fact, although the burning rates start off
Figure 4.1: The Meynet et al. (1994) evolutionary tracks at $Z_{\odot}$ for 120, 85, 60, 40 and 25 $M_{\odot}$. The main sequence is depicted with a thick black line, and the W-R phase as a dashed line.
Figure 4.2: The Meynet et al. (1994) evolutionary tracks at 0.2 $Z_\odot$ for 120, 85, 60, 40 and 25 $M_\odot$. The main sequence is depicted with a thick black line, and the W-R phase as a dashed line.
higher, the rate slows down very quickly due to the low core temperatures and pressures in the higher metallicity models caused by the greater mass-loss. Due to the Meynet et al. (1994) inferred metallicity scaling of $\dot{M} \propto Z_\odot^{0.5}$, the $Z_\odot$ mass-loss rate is 2.2 times as great as the 0.2 $Z_\odot$ model. One implication of this is that the high metallicity model actually burns hydrogen for longer than the low metallicity counterpart (3.04 Myr as opposed to 2.90 Myr; Meynet et al. (1994)), although it enters the W-R phase earlier since its mass-loss is high enough to strip away the upper atmosphere of the star. Table 4.2 also shows that low metallicity stars spend a far greater proportion of their time as the less evolved WN stars, whereas the $Z_\odot$ models spend roughly half their time as W-R stars in the form of the WC subtype. This is again due to the high mass-loss, as it tends to keep stripping the atmosphere layers off even during the WNL stage. Although the Meynet et al. (1994) evolutionary tracks are very successful in explaining massive star lifetimes, there are several inconsistencies with observations:

- **The time massive stars dwell on the MS.** Maeder & Meynet (1989) find that the MS lifetimes of stars in clusters is longer than that of the track predictions, a problem that the assumed moderate overshooting cannot account for.

### Table 4.2: W-R star lifetimes, for 1.0 $Z_\odot$ and 0.2 $Z_\odot$ reproduced in part from Maeder & Meynet (1994).

<table>
<thead>
<tr>
<th>Initial Mass ($M_\odot$)</th>
<th>t(WR) ($10^5$yr)</th>
<th>t(WNL) ($10^5$yr)</th>
<th>t(WNE) ($10^5$yr)</th>
<th>t(WC) ($10^5$yr)</th>
</tr>
</thead>
<tbody>
<tr>
<td><strong>1.0 $Z_\odot$</strong></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>120</td>
<td>19.586</td>
<td>11.504</td>
<td>4.395</td>
<td>3.687</td>
</tr>
<tr>
<td>85</td>
<td>8.520</td>
<td>2.818</td>
<td>0.149</td>
<td>5.552</td>
</tr>
<tr>
<td>60</td>
<td>5.130</td>
<td>0.789</td>
<td>0.171</td>
<td>4.171</td>
</tr>
<tr>
<td>40</td>
<td>5.443</td>
<td>0.336</td>
<td>0.525</td>
<td>4.581</td>
</tr>
<tr>
<td><strong>0.2 $Z_\odot$</strong></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>120</td>
<td>5.783</td>
<td>2.352</td>
<td>0.005</td>
<td>3.426</td>
</tr>
<tr>
<td>85</td>
<td>4.100</td>
<td>1.324</td>
<td>0.039</td>
<td>2.736</td>
</tr>
<tr>
<td>60</td>
<td>1.967</td>
<td>1.315</td>
<td>0.132</td>
<td>2.747</td>
</tr>
</tbody>
</table>
• **Surface abundances of massive stars.** Several works find that O, B and A supergiants are observed to be overabundant in the products of CNO cycle burning compared to the Meynet *et al.* (1994) evolutionary tracks. Herrero *et al.* (1992) and Villamariz *et al.* (2002) find an overabundance of He in fast rotating O stars; Walborn (1976), McErlean *et al.* (1999) and Smartt *et al.* (2002) find an overabundance of C and N in a sample of OB supergiants; Venn (1999) (1999; 1995) find nitrogen variations in 10 A supergiants in the Small Magellanic Cloud which cannot be predicted by the evolutionary tracks.

• **The ratio of blue to red supergiants at high metallicity.** Langer & Maeder (1995) notice that although evolutionary tracks can account for the ratio of blue to red supergiants at low Z, the number of blue supergiants predicted at high Z is far too low to agree with observations.

• **W-R star predictions.** There are several problems with W-R star evolutionary models, most importantly, the W-R to O ratios can only be produced by using mass-loss rates at 2 times the observed value. Other problems include discontinuities in the transition from hydrogen rich to hydrogen poor atmospheres, which is observed to be continuous in the W-R population, as well as the under-prediction of WN/WC transition stars (1 per cent as opposed to 4-5 per cent observed; Conti & Massey (1989)).

Recently, these problems have been solved via the inclusion of rotation, which was previously considered a second order effect (Meynet & Maeder 2000). Improvements to the theory of transport mechanisms induced by rotation (Zahn 1992) have allowed for new evolutionary grids to be calculated. A series of publications by Maeder and Meynet have introduced these new evolutionary grids (starting with Maeder & Meynet 1997). In the review of Maeder & Meynet (2000), the main improvements of using rotation are explained. These new grids predict rotating stars to be more luminous than their non-rotating counterparts. The lifetimes of these stars is also increased due to the mixing of the outer stellar envelope with the core, prolonging H and He burning lifetimes. The W-R phase of the star is also increased, for example a 60$M_\odot$ star will have a W-R lifetime a factor of 3 greater than a non-rotator, due to higher mass-loss and chemical mixing of unprocessed material.
(Maeder & Meynet 2000). The surface chemistry of rotating stars is now altered due to the degree of rotation that the star experiences accounting for the observed enrichment. The lower mass-loss limit for a WN star is now decreased. At 1.0 $Z_\odot$, instead of a 40 $M_\odot$ limit, we have the significantly lower value of 25 $M_\odot$. Due to this effect, the need to increase mass-loss rates is no longer necessary. Unfortunately, these tracks were not available for use with SB99, so we must use the Meynet et al. (1994) considering the above caveats.

4.2 The Starburst99 Model

4.2.1 An Overview of the Model

Leitherer et al. (1999) present the evolutionary synthesis code SB99 which predicts the observable properties of galaxies undergoing active star formation, and is tailored to the analysis of massive star populations. The code is an improved version of that previously published by Leitherer & Heckman (1995). It is based on the evolutionary tracks of Meynet et al. (1994) for enhanced mass loss and the model atmosphere grid compiled by Lejeune et al. (1997), supplemented by the pure helium W-R atmospheres of Schmutz et al. (1992). The SB99 model relies on two major areas of stellar physics, that of evolutionary stellar physics and that of stellar atmospheric physics. Both sets of theories are extremely complex, making it impractical to model the population straight from theory. Instead, the obvious answer is to use pre-calculated data in the form of model grids and with careful interpolation, these can be used to cover all the required parameter space. The evolutionary tracks used in SB99 are those of the Meynet et al. (1994) grid as discussed in the previous section. Here we will discuss the part of SB99 concerned with the generation of synthetic starburst SEDs. This output is by far the most useful single output, since it can be applied to photoionization codes, and can be used as a basis for the calculation of a number of other important observables, such as synthetic spectra and emission line equivalent widths. The program requires 21 starting variables in order to define the population. These are read in from a text input file. Some of these variables are unimportant as they correspond to other outputs not explored in this thesis. These outputs, unchanged by this work are discussed in detail in Leitherer et al. (1999). The variables are described in Table 4.3.

From this table we can see that the number of free parameters is intentionally extremely
Figure 4.3: A flowchart of the main spectral synthesis procedure
# CHAPTER 4. EVOLUTIONARY SYNTHESIS CODES

Table 4.3: Specified variables in SB99

<table>
<thead>
<tr>
<th>Variable</th>
<th>Significance</th>
<th>Used in spectral synthesis?</th>
</tr>
</thead>
<tbody>
<tr>
<td>$i_{sf}$</td>
<td>Star formation channel switch.</td>
<td>✓</td>
</tr>
<tr>
<td>$M_{tot}$</td>
<td>Total mass of starburst (instantaneous burst only)</td>
<td>✓</td>
</tr>
<tr>
<td>$\dot{N}$</td>
<td>Star formation rate (continuous burst only)</td>
<td>✓</td>
</tr>
<tr>
<td>$\alpha$</td>
<td>IMF power-law exponent</td>
<td>✓</td>
</tr>
<tr>
<td>$M_{up}$</td>
<td>IMF upper mass cutoff</td>
<td>✓</td>
</tr>
<tr>
<td>$M_{low}$</td>
<td>IMF lower mass cutoff</td>
<td>✓</td>
</tr>
<tr>
<td>$M_{SN}$</td>
<td>Supernova mass cutoff</td>
<td>×</td>
</tr>
<tr>
<td>$M_{BH}$</td>
<td>Black hole mass cutoff</td>
<td>×</td>
</tr>
<tr>
<td>$Z_M$</td>
<td>Metallicity selection, specifying 1 or $2 \times \dot{M}$</td>
<td>✓</td>
</tr>
<tr>
<td>$i_{wind}$</td>
<td>Wind model switch</td>
<td>✓</td>
</tr>
<tr>
<td>$T_{min}$</td>
<td>Initial time</td>
<td>✓</td>
</tr>
<tr>
<td>$\Delta t$</td>
<td>Discrete time step for calculation</td>
<td>✓</td>
</tr>
<tr>
<td>$T_{max}$</td>
<td>Last time step</td>
<td>✓</td>
</tr>
<tr>
<td>$i_{grid}$</td>
<td>Switch specifying resolution of the mass grid</td>
<td>✓</td>
</tr>
<tr>
<td>$L_{min}, L_{max}$</td>
<td>Switches specifying track mass boundaries</td>
<td>✓</td>
</tr>
<tr>
<td>$i_{atmos}$</td>
<td>Switch specifying atmosphere grids to be used</td>
<td>✓</td>
</tr>
<tr>
<td>$i_{line}$</td>
<td>Metallicity of UV line spectrum</td>
<td>×</td>
</tr>
<tr>
<td>$v_{RSG}$</td>
<td>Red Supergiant micro turbulence velocity</td>
<td>×</td>
</tr>
<tr>
<td>$i_{Z_{RSG}}$</td>
<td>Switch specifying Red Supergiant abundance</td>
<td>×</td>
</tr>
</tbody>
</table>

limited. The only free parameters we use in a spectral synthesis model define the IMF shape ($M_{tot}$, $M_{up}$, $M_{low}$ and $\alpha$) and the chemistry of the starburst ($Z_M$). In fact this number can be reduced further by assuming values for $M_{up}$ and $M_{low}$ of usually $100M_\odot$ and $1.0M_\odot$ respectively. A discussion of the IMF is given in Chapter 1. By reducing the number of degrees of freedom, the model’s sensitivity to the remainder may be explored more closely.
The SB99 Model

Due to the reliance on model grids, SB99 typically takes only a couple of minutes to complete its calculation. The flowchart in Fig 4.3 gives a representation of the important processes used in the calculation of a time series set of synthetic starburst continua. The first operation is to define the zero age population in terms of numbers in discrete mass bins. For this, an IMF is used, satisfying total mass considerations, and being bound by upper and lower masses. The population is then defined as in Equation 1.4 in Chapter 1 with a negative power-law slope, such that the low mass population outnumbers the high mass stars by many factors. The equation is solved analytically to give the number density at every mass point:

\[
N(m) = \left[ \frac{C}{1 - \alpha} \right] \left( (m + \Delta m)^{1-\alpha} - (m - \Delta m)^{1-\alpha} \right)_{\alpha \neq 1}
\]

\[
if \alpha = 1; \Rightarrow N(m) = C(\log (m + \Delta m) - \log (m - \Delta m))
\]

where \(C = \frac{M_{\text{tot}}(2 - \alpha)}{M_{\text{up}}^{2-\alpha} - M_{\text{low}}^{2-\alpha}}\)

\[
if \alpha = 2; \Rightarrow C = \frac{M_{\text{tot}}}{\log M_{\text{up}} - \log M_{\text{low}}}
\]

Next the evolutionary tracks are read in from the Meynet et al. (1994) grids, using the correct metallicity and mass-loss required by the input. As explained above the 2 \( \times \) \( \dot{M} \) tracks are used as default as these give the correct observed W-R lifetimes. The tracks give information on the mass, surface chemistry, temperature, luminosity and mass-loss rate of a star at a given age. Once the evolutionary tracks are read from file, they are manipulated so that they are used to parameterise the mass bins in terms of not only number, but the other relevant quantities from the tracks (surface chemistry, luminosity, temperature, mass remnant and mass loss etc.). The grid provided by Meynet et al. (1994) is not linearly spaced in either mass or age, due to the non-linear nature of stellar evolution where large changes occur over short time periods, but SB99 works in a linear fashion with age and mass, and can cope with a high resolution in both these dimensions. For these reasons, the tracks are interpolated between both mass and age, using a linear interpolation routine where a star resides on the main sequence (where its properties change slowly) and a more complicated spline fit off the main sequence, where changes occur more rapidly. The results of this interpolation are illustrated in Figure 4.4 for the case of a 50\(M_\odot\) star.

The interpolated tracks are only calculated at the discrete time intervals required for
Figure 4.4: The 60$M_\odot$ and 40$M_\odot$ solar metallicity Meynet et al. (1994) models (dashed and dotted lines), and a 50$M_\odot$ (unbroken line) model produced by SB99
the SB99 calculation, and therefore the $50M_\odot$ model does not follow the two models used to calculate it into a yellow supergiant phase. This is because this phase lasts for $< 40,000$ years (Meynet et al. (1994); $60M_\odot$ model), much smaller than a typical time interval used in SB99, in which it makes sense to use time steps of $> 100,000$ years for computational reasons, such as time and memory. In Figure 4.4, the time step is sufficiently large that the red-ward movement of the star is completely omitted from the calculation, so does not appear as a point on the track. However, since these these yellow supergiants/hypergiants are extremely rare in nature, compared to an O star due to their brief lifetimes, this presents no problem to the overall ensemble. Note also, that the end of the W-R phase is missing in this track. Again the star evolves very quickly in the WO stage near supernova. A more general outline of stellar evolution can be found in the previous section.

Once the grids have been interpolated to the relevant resolution, they are assigned to several two dimensional vectors, holding the information of mass, temperature, abundances, mass-loss and luminosities as a function of age and zero age mass. These vectors are all important in SB99, and are used (once rescaled at each time step to insure a normalised IMF) as representative stellar components of the starburst. To generate a synthetic continuum, these components are manipulated in the following way, although all starburst outputs are generated via the integration of all components in a certain time period.

**Generation of Synthetic Spectra**

The most important process in the calculation of the spectra is the coupling of a single stellar component with the grids of atmospheres used to represent the stars. SB99 uses the compiled atmospheres of Lejeune et al. (1997) to represent all stars from M to O, dwarf to supergiant. This is done by using the temperature vector from the evolutionary tracks and creating a surface gravity vector. This is calculated from the following simple formula:

$$\log g = \log m(t) + 4.0 \log T_{\text{eff}} - \log \frac{L}{L_\odot} - 10.6 \quad (4.2)$$

using the temperature ($T_{\text{eff}}$), luminosity ($L$) and mass ($m(t)$), the time dependant vectors from the evolutionary tracks. Since the Lejeune et al. (1997) model grid has been calculated with two dimensions, temperature and luminosity, a nearest fit is then calculated using a simple algorithm and associated with each stellar component. Only the
hottest, more evolved O stars fail to match well with this grid of atmospheres, where a 50kK supergiant of typically log($g$) $\sim$ 3.0 is represented by an atmosphere of log($g$) = 5.0. For comparisons between the properties of the two surface gravities, we refer the reader to Figure 2.3 of Chapter 2, which shows the difference in ionizing properties of the two luminosity classes.

The Wolf-Rayet stars are coupled in a different way, due to the use of the Schmutz et al. (1992) pure helium non-LTE grid to represent these stars. SB99 uses the W-R atmospheres when certain conditions are met; the surface hydrogen of a stellar component has dropped below 40 per cent, and the effective temperature is above 25,200 K. Once these conditions are satisfied, the W-R grids are used.

Two grids are used in SB99 to model the different velocity fields found in W-R winds, expressed as a function of radius as in Equation 2.18 of Chapter 2 simplified by the assumptions that $V_0 = 0$ and $\beta_1 = \beta_2$. This simplified form was first suggested by de Loore et al. (1982). The first grid has a $\beta$ exponent of 1, to represent dense, slower winds, and the second grid using the $\beta = 2$ exponent to model the less dense faster winds in hotter more extreme stars. The differences in the velocities fields are represented in Fig. 4.5 showing the far slower convergence of the velocity law with the larger $\beta$ component. The velocity law of a model may have a large effect on the ionizing output of a star. For instance, Figure 4.6 shows the difference between two 90 kK models, both with the same radius. The $\beta = 2$ model shows a huge excess of flux below 228 Å compared to the $\beta = 1$ model. This is because the steeper velocity field allows He$^+$ to recombine at some point in the atmosphere, whereas the $\beta = 2$ model allows no such recombination, therefore no line absorption can take place. For this reason Schmutz et al. (1992) choose to represent stars of $T_{\text{eff}} < 90$kK with the $\beta = 1$ grid and stars of $T_{\text{eff}} \geq 90$kK with the $\beta = 2$ grid.

Schmutz et al. (1992) decided that their W-R grid should have two dimensions, effective temperature and radius, to represent a scatter in not only temperature but luminosity values, crucial for the ionizing properties in these models. This coupling is far more complicated than for the Lejeune et al. (1997) case. The transformed radius is used to couple the stellar component to the evolutionary tracks, giving a temperature at some point in the W-R wind. The transformed radius is more meaningful in a W-R wind because it is scaled with the density of an outflow and can define the point where matter and radiation cease to interact (the Meynet et al. (1994) grids assume this happens at a scale height of
Figure 4.5: Comparison of the velocity field with a $\beta$ exponent of 1 (unbroken line) and of $\beta = 2$ (dashed line). The figure shows a model of $v_{\infty} = 2500$ km s$^{-1}$.
Figure 4.6: Comparison of two 90kK Schmutz et al. (1992) models with different velocity laws. The $\beta = 1$ model is represented by an unbroken line and the $\beta = 2$ model is represented by a dashed line. Both have the same radius.
\[ \tau = \frac{2}{3} \]. The transformed radius is (Schmutz et al. 1989):

\[ R_t = R_s \left( \frac{0.4 M_\odot \text{yr}^{-1}}{dM/dt} \frac{V_\infty}{2500 \text{km s}^{-1}} \right)^{\frac{3}{5}} \]

(4.3)

where the mass-loss, \( \frac{dM}{dt} \) and \( V_\infty \) are calculated from either empirical data versus W-R subtype, from the evolutionary track (\( \dot{M} \) only) or a theoretical relationship. The default relation is the theoretical option (explained in §4.6). This new radius is then used to find a nearest fit to the grid.

The grid then needs to be coupled with the temperature of the stellar component to ensure an accurate representation of its atmosphere. Since W-R stars have dense winds, where the photosphere of the star is effectively hidden, this process is not straightforward. For this reason, the Meynet et al. (1994) grids come with two different measures of temperature, to solve the problem of temperature diagnostics in stars with dense winds. The problem was addressed by Maeder (1990) and Meynet et al. (1994), with the solution being to correct the hydrostatic \( T_{\text{hyd}} \) value by assuming a velocity law to give the radius corresponding to an optical depth of \( 2/3 \) (i.e. far out in the wind) and hence a lower temperature \( T_{2/3} \). This does not effect stars with optically thin winds, since their winds are too weak to interact at this depth, but does not solve the problem entirely because the corresponding \( T_{2/3} \) temperatures derived from W-R model atmospheres analyses are generally far too low to accurately describe the problem. The matter within a W-R atmosphere will interact with the emergent radiation from the core. Processes such as absorption and reemission alter the appearance of the spectrum, cooling the wind. Schmutz et al. (1992) claim that a W-R model atmosphere is best characterized by the core temperature \( T_* \) which is closer to the radius where the optical depth is \( \approx 10 \) (corresponding to a thermalization depth of unity for the continuum opacity which is dominated by electron scattering). This core temperature is used to couple W-R stars to the Schmutz et al. (1992) model grid. However, as we shall see in §4.3, this temperature is also unsatisfactory, and yields an unphysically high temperature scale for W-R stars.

Once the correct atmosphere has been assigned to a stellar component, care must be taken to assure that the atmosphere is normalised, so that the total energy is conserved in the star. Since no interpolation between grid models is performed, the atmosphere must be scaled to fit the luminosity of the star that it represents. This is done in very
much the same way as previously described in Chapter 2, Equation 2.19, using a Planck function as the renormalization factor. The reason that no interpolation is performed, is due to the rapid change in the appearance of neighbouring atmospheres in the EUV continuum region, which may change from an undetectable level, to a level higher by many magnitudes. A careless interpolation in these cases may result in excess flux being predicted in the starburst, as winds suddenly become transparent at a certain threshold due to the energy requirements of complete He\(^+\) ionization without recombination. In fact, early versions of SB99 included this interpolation.

Since all stellar components can now have a model atmosphere assigned to them, the emergent starburst continuum is simply calculated via the summation of all individual components, and a time series created by the systematic repetition of this process at every time step, with careful consideration of the re-initialising of the flux variable at the start of each time step.

This method is applied to the calculation of many different observables of the model starburst, and are all calculated linearly on the same time step. This Chapter only outlines the outputs applicable to this new set of atmospheric grids.

### 4.3 Implementation of the Atmosphere Grids

#### 4.3.1 O Star Grids

The first attempt at replacing the inappropriate LTE atmospheres of Lejeune et al. (1997) was by Schaerer & Vacca (1998), who replaced the models pertaining to the O and early B stars with the CoStar model atmospheres of Schaerer et al. (1996). To integrate our new O star grid, we have simply replaced the Lejeune et al. (1997) LTE models by the WM-basic models for effective temperatures above 25000 K, re-mapped to 1221 points over the wavelength range \(\lambda = 91 \text{ to } 1.6 \times 10^6 \text{ Å} \), as required by the code. To ensure a smooth transition to the LTE models, we adopt the method of Schaerer & Vacca (1998): the WM-basic models are restricted to the \(T_{\text{eff}}-\log g\) domain defined by \(\log g \geq 2.2\) and \(\log g < 5.71 \times \log T_{\text{eff}} - 21.95\). This limits the use of the early B supergiants, which do not converge smoothly to the flux level of the LTE models at lower temperatures (~ 25kK).

The difference between the LTE and non-LTE models at this temperature is illustrated in Figure 4.7. This figure clearly shows the non-LTE model to predict far more ionizing flux below 228 Å and 504 Å. We adopt the same method used by Leitherer et al. (1999) in
Figure 4.7: An illustration of the differences between an LTE and non-LTE model at 25kK. Plotted in red (dashed line) is a CMFGEN 25kK model. A Lejeune et al. (1997) 25kK supergiant model is plotted against it in a black (solid) line.
choosing the best model atmosphere to represent an evolutionary point by using the nearest fit in $\log g$ and $T_{\text{eff}}$, but interpolation is not used as this may break flux conservation. A Planck function is used to re-normalize the atmospheres to the correct luminosity, since O star radii from the evolutionary tracks may deviate from the new grid values (Chapter 2).

### 4.3.2 W-R Star Grids

The new W-R grid is based on current best estimates of W-R temperatures which are considerably lower than those of the Schmutz *et al.* (1992) grid, which is coupled to the $T_{\text{hyd}}$ evolutionary model core temperatures. Although no formal temperature-spectral type calibration exists, it is obvious to the observer that the range of temperatures used in the implementation of the Schmutz *et al.* (1992) models is too high, with extreme models reaching values greater than 200kK. The hottest examples of early W-R stars reach as much as 160kK (e.g. Br2, Chapter 3), but observationally these are extremely rare objects.

The problem we face is therefore how to match a W-R model atmosphere defined by $T_*$ to an evolutionary model described by two temperatures: $T_{\text{hyd}}$ and $T_{2/3}$. As both these temperatures are inadequate, since one is too hot and the other too cool, we must look at the available evidence in the form of observed W-R star temperatures for inspiration. If we look at Figure 4.8 we can see the temperature distribution of W-R stars from the Galactic survey of van der Hucht *et al.* (1981) using a rough W-R calibration scale (Crowther, private communication), which shows a peak in the WN population at roughly 80kK and a bias towards cooler stars for the WC subtype, with a typical temperature being 50-70kK. Although this type of distribution may not be representative of a starburst environment, where the IMF may have an extended upper mass limit compared to field star formation, it is currently the only evidence we can look at to derive a temperature distribution at all. This evidence must be considered with extreme caution, but does point towards the hot, extreme subtypes being rather rare. Figure 4.9 shows the W-R populations of a solar metallicity starburst from SB99 using three different conditions; the two evolutionary temperatures $T_{\text{hyd}}$ and $T_{2/3}$, and a new one $T_*$, a linear interpolation at some point between the 2 values. These histograms were calculated by placing the components of
Figure 4.8: Histogram plots of the temperature distribution of Galactic WN and WC subtypes. The top plot shows the distribution of WN stars at solar metallicity and the bottom plot show the WC populations. These plots were made using the Galactic survey of van der Hucht et al. (1981) with the temperature calibrations from Crowther (2002; priv. comm.).
Figure 4.9: Histogram plots of the temperature distribution for the WN and WC phases summed over the lifetime of the W-R phase as given by SB99 with the Meynet et al. (1994) evolutionary tracks at solar metallicity for an instantaneous burst with $M = 10^6 M_\odot$, $\alpha = 2.35$ and an upper mass of $100 M_\odot$. For each W-R type, we show the temperature distributions corresponding to the hydrostatic temperature $T_{\text{hyd}}$ (light grey), the corrected hydrostatic temperature $T_{2/3}$ (dark grey) of Meynet et al. (1994), and the adopted $T_* = 0.6 T_{\text{hyd}} + 0.4 T_{2/3}$ (black). The temperature bin sizes are 10 kK.
an instantaneous burst model into temperature bins depending on temperature and W-R type defined by the three temperature scales and surface chemistry, and integrated over the entire starburst lifetime. If we first consider the pale grey bars in Figure 4.9, we see the population according to the corrected $T_{\text{hyd}}$ temperature values. One can easily see that the WN population is double peaked, a feature that does not appear in the observation, although the average temperature is quite close to that of the Galactic distribution. Looking at the $T_{\text{hyd}}$ distribution in the WC population, we can see that it is far too hot, with temperatures extending out to greater than 175kK, hotter than the most extreme WO star. We also see a peak at 155kK, far beyond that of the data. It is safe to say that this temperature distribution yields a population that is far too hot.

At the other extreme, plotted in black we see the $T_{2/3}$ W-R temperatures, which are calculated far out in the dense W-R wind (see §4.1). The top plot of Figure 4.9 shows that the WN temperatures are indeed predicted to be too cool. The peak population is at roughly 45kK, compared to the observation at 80kK, and although the WC population yields a cooler distribution, we now find that it is indeed too cool. There are now no WC stars predicted above 65kK, although the peak coincides quite well with the data.

Since the new W-R model grids have been calculated to mimic a realistic population of W-R stars, both these distributions are unsatisfactory. The most thorough way of solving this problem would be to abandon the Meynet et al. (1994) evolutionary tracks in favour of the new more accurate models with rotation. However, this is beyond the scope of this thesis, so a compromise must be made instead. We know that a more accurate representation of the temperature distribution should lie at some point between the two extremes, and that the correct temperature would originate at an optical depth of $\sim 10-20$.

To create a satisfactory distribution, $T_*$ was calculated to lie in between $T_{\text{hyd}}$ and $T_{2/3}$. Using a simple linear interpolation technique to determine the point between the two temperatures, the new distribution was repetitively calculated until a point was found that reproduced the observations as well as possible. This point was eventually set at $T_* = 0.6T_{\text{hyd}} + 0.4T_{2/3}$. Figure 4.9 shows this distribution plotted in dark grey. The WN population was set at point just cool enough to inhibit the double peak present in the $T_{\text{hyd}}$ population, although it can be seen that the temperatures are too low on average. The WC population however presents the opposite problem, as it peaks at temperatures that are too high to be representative of the real WC population, where a 50-70kK peak
is found rather than at 115kK. Since raising the temperature of the distribution towards
the hydrostatic value results in an increase in both the WN and WC populations, thus
yielding a WC population that is even hotter, and lowering the temperature towards $T_{2/3}$
yields an even cooler WN population, this level was deemed to be the closest possible fit
with the available tracks. Originally it was decided to deal with the temperatures of the
two populations separately, due to their different positions on the HRD, but this notion
was eventually rejected in favour of consistency between the two populations.

Although the adopted temperatures do not seem to exactly match that of the van der
Hucht survey, the temperatures do span the range that has been discovered by several
recent works (e.g. Hillier & Miller (1998); Herald et al. (2001); Crowther et al. (2003b)).
Also in favour of the adopted values is the fact that the WN and WC temperature peaks
 correspond to late WN and early WC stars. Studies of W-R galaxies (e.g. Guseva, Izotov
& Thuan 2000) show that the composite W-R ‘bump’ at $\sim 4686$ Å is usually dominated
by these W-R subtypes. Chapter 5 looks closely at the synthesis of these features, which
gives an idea of the W-R populations.

Next, some switch is needed to differentiate between a stellar component that is a W-R
star and O and B stars. This cannot simply be done via age constraints since this changes
with metallicity, but by using knowledge of the surface chemistry of the star. Furthermore
the W-R grid has been split into WN and WC subtypes, due to the differences in mass-loss
and chemistry. The most obvious way of selecting the correct grid is to use a switch in SB99
developed by Leitherer et al. (1999) to count the population densities of each evolutionary
point, including W-R stars. This switch uses several limiting factors to calculate WN and
WC species and bins them into either late or early subtypes using the following method:

- **WN L** $0.1 < X(H) < 0.4$
- **WNE** $X(H) < 0.1$ and $\frac{X(C_{12})}{X(N_{14})} < 10.0$
- **WCL** $X(H) < 0.1$ and $\frac{X(C_{12})}{X(N_{14})} > 10.0$ and $\frac{X(C_{12}) + X(O_{16})}{X(H)_{\lambda 4026}} < 0.5$
- **WCE** $X(H) < 0.1$ and $\frac{X(C_{12})}{X(N_{14})} > 10.0$ and $\frac{X(C_{12}) + X(O_{16})}{X(H)_{\lambda 4026}} < 1.0$
- **WO** $X(H) < 0.1$ and $\frac{X(C_{12})}{X(N_{14})} > 10.0$ and $\frac{X(C_{12}) + X(O_{16})}{X(H)_{\lambda 4026}} > 1.0$

Here $X(Z)$ represents the surface abundance of each element $Z$. This formalism means
that WN stars will only be selected when nitrogen is present in reasonable abundance
in the star, and WC stars when no nitrogen is present, as is required by observation. Currently, there is only a need to differentiate between WN and WC stars, although this method makes it easy for future updates if observations can more tightly tie down W-R parameters, especially mass-loss with early and late subtypes.

Finally, a switch is included to restrict the W-R grids to the solar metallicity part if required by the user for any starburst metallicity. This is added since a mass-loss scaling with metallicity is indeed very uncertain, with no strong evidence available at present, only observational hints. As mentioned in Chapter 2, the ionizing flux level of a stellar atmosphere is overwhelmingly dependent on wind density, rather than the input chemistry, hence a solar metallicity atmosphere will mimic another atmosphere very well at the same mass-loss but different metallicity. This is added as a flag in the main body of the code but not added to the input file, although it may be changed by the experienced user. A constant mass-loss vs metallicity scenario is not discussed in this thesis due to the belief that scaling is in fact appropriate (for discussion on this matter we refer the reader to Chapter 2).

4.4 Comparisons With Other Work

To show the usefulness of the new atmospheric grids applied to a population synthesis model, it is important to compare like with like in order to establish differences. In order to add the new grids to the SB99 model, we used SB99 v3.1, running complete with LTE Lejeune et al. (1997) grids to represent all stars from M to O and the Schmutz et al. (1992) grids to represent W-R stars. However, since Schaerer & Vacca (1998) included the Costar models of Schaerer et al. (1996) into their evolutionary synthesis code, it is prudent to compare with this more recent work as well, which replaces the LTE O star atmospheres with the non-LTE CoStar models. The population synthesis code used here however, contains some differences to the Schaerer & Vacca (1998) work; in their work a 20 percent binary fraction has been added to reproduce the W-R to O ratios found at lower metallicity, also, the evolutionary tracks are not interpolated between the lowest mass W-R star and that of the next track down, which is not massive enough to form a W-R. Although this feature preserves the integrity of the evolutionary tracks, since the W-R and non W-R behaviour is quite different, this problem only occurs over a relatively short period of time for a small mass range (i.e. $25M_\odot$ to $40M_\odot$ for solar metallicity),
so that the change is small. Due to the unavailability of the Schaefer & Vacca (1998) code\(^1\), it was decided that the CoStar models would be implemented into SB99 for the sake of comparison, since two grids \((Z_\odot\text{ and } 0.2 Z_\odot)\) are freely available on the internet at http://webast.ast.obs-mip.fr/people/schaerer/SdK96.html. These were implemented into the Starburst99 model in the same way as our grid described above, for consistency, since the evolutionary track interpolation method of Schaefer & Vacca (1998) may hide trends caused solely by the difference in O star models. The CoStar models are described in Chapter 2. This method also gave an excellent standard with which to test the consistency of the method we used to integrate the atmospheres into SB99, since we used the models presented in the Schaefer & Vacca (1998) as our standard.

There are two phases of star formation available to use in SB99, an instantaneous burst model and continuous star formation. Both these modes of star formation are important, so are compared below, with different results. The instantaneous burst model assumes all stars are formed at the same time, and hence evolve together, so the effect of different components of the population can be seen with the evolution of the starburst. The continuous mode assumes a constant, but less cataclysmic star formation rate, so adds on a small amount of mass, divided up into bins with an IMF at every time-step, and hence stars evolve at different time periods and at some time the output converges to an equilibrium population. This mode is dominated by the most massive stars formed in the IMF, the early O and W-R stars. Both these modes are observed, although a starburst is never exactly instantaneous and may last \(\sim 10\) Myr (Mouhcine & Contini 2002).

4.4.1 Instantaneous Bursts

Differences in Synthesis Models

In all models the adopted total mass is \(10^6 M_\odot\), with a IMF consisting of an upper mass, \(M_{up} = 100 M_\odot\), a lower mass of \(M_{low} = 1 M_\odot\) and a Salpeter IMF of \(\alpha = 2.35\). Although it is unclear whether these limits satisfactorily describe a starburst, all models use the same assumption which allows the most massive atmospheric models to be used in the comparison. Figures 4.10 & 4.11 show the comparison in SEDs of the new atmospheres compared to that calculated by the SB99 model of Leitherer et al. (1999). At solar

\(^1\)The outputs to certain starburst models are now available at http://webast.ast.obs-mip.fr/people/schaerer/frame.main.html# Our models, but were not at the time of comparison.
metallicity (Figure 4.10) several crucial differences between the two plots can be noticed. Starting at the earliest age, where early O V stars dominate the ionizing flux of the starburst, it can be seen that the Leitherer et al. (1999) model has negligible ionizing flux below 228 Å (log λ = 2.36). This is not the case for the new models, since they now predict some flux below the He$^{+}$ continuum edge due to the addition of the WM-basic non-LTE models. One also notices that the continuum now appears some what more ‘spiky’ due to the rigorous inclusion of non-LTE line blanketing.

Later on the differences in the instantaneous burst models become more apparent, when the W-R phase of the starburst arises at an age of between 3 and 5 million years. The Leitherer et al. (1999) model predicts a large amount of flux below the 228 Å continuum edge due to the unblanketed Schmutz et al. (1992) models, whereas the new CMFGEN model W-R stars of Hillier & Miller (1998) exhibit an entirely different behaviour. These models predict a negligible level of flux below 228 Å , and as seen below in the ionizing quanta plots, the difference between the two models can be as much as a factor of $10^8$, dropping far below any observational limit. This is entirely due to the blanketing effect of the dense W-R winds which was unaccounted for in the Schmutz et al. (1992) models. By 10 Myr, the earliest stars have disappeared from the starburst and the models converge to a single result, since the Lejeune et al. (1997) models are now used to represent the remaining population in both cases.

The differences at lower metallicity are less extreme. Figure 4.11 shows the case for a starburst in a Small Magellanic Cloud (SMC) like environment with a metallicity 0.2 $Z_{\odot}$. At very early ages ($\lesssim 1$ Myr) the new O star atmospheres result in a larger amount of ionizing flux below the 228 Å continuum edge than the Leitherer et al. (1999) model. This is due to the non-LTE effects in the hotter atmosphere allowing less He$^{+}$ to recombine therefore ionizing flux to escape the less dense O star winds. However with regard to the W-R phase, which at low metallicity is much shorter due to the higher values of mass required to create a W-R star (Meynet 1995), it can be seen that the two models agree far more closely in the continuum below 228 Å , although the new models are still a factor of 100 lower. The lower wind density allows for less recombination in the W-R wind at low metallicity, and hence is more transparent to the far UV ionizing flux. Again we see consistency in the models at ages of greater than 10 Myr since the model atmospheres invoked are the same.
Figure 4.10: Spectral energy distributions of the new models (top) compared with SB99 (bottom) at time intervals of 1, 3, 4, 5 and 10 Myr for an instantaneous burst at $Z_\odot$. 
Figure 4.11: Spectral energy distributions of the new models (top) compared with SB99 (bottom) at time intervals of 1, 3, 4, 5 and 10 Myr for an instantaneous burst at $0.2 Z_{\odot}$. 
Figure 4.12: The evolution of the photon luminosity for ages of 1-8 Myr in the ionizing continua of hydrogen ($\log Q_0$), He I ($\log Q_1$) and He II ($\log Q_2$) at $Z=Z_\odot$ for an instantaneous burst at time steps of 0.1 Myr. The new models (black, solid) are compared to the ionizing fluxes of: Leitherer et al. (1999) (red, dotted); Schaerer & Vacca (1998) (blue, dashed)
It is perhaps easier to see the effects of the new atmospheres if we turn to the $Q_2$, $Q_1$ and $Q_0$ ionizing fluxes defined in Equation 2.17 in Chapter 2, which give the sum of the fluxes below 228 Å, 504 Å and 912 Å respectively. Figure 4.12 shows the $Q$ values for a solar metallicity model, showing the predictions of the new work, Leitherer et al. (1999) as well as that of SB99 with the Schaerer et al. (1996) CoStar models, our analogue to the work of Schaerer & Vacca (1998) (from now on SV98). Concentrating on the hardest $Q$ value, $Q_2$, we can easily see a huge drop in the value in the new models during the W-R phase. Since both the SV98 models and the Leitherer et al. (1999) (from now on SB99) models both use the Schmutz et al. (1992) W-R grid, they exhibit no differences in the W-R phase. It can be seen that at early epochs, when O stars dominate that there is a clear departure from the SB99 for both SV98 and the new models (from here on SB2002). The CoStar O star models clearly show the hardest ionizing continuum, with the WM-basic models providing a lower level of far UV flux, although much closer to the SV98 value. In the earliest time period the difference in flux between the two non-LTE models is about a factor of 10, compared to the LTE models which predicts a flux of almost a factor of $10^4$ less. The difference between the two non-LTE models lies in the derivation of line-blanketing, in which the WM-basic model does a far more thorough job (Chapter 2). At 2 Myr the difference in these two models is exaggerated even further, due to the inclusion of O supergiants, in which the departure is more extreme.

If one considers the $Q_1$ plot, one can see that again the CMFGEN models predicts less flux than the Schmutz et al. (1992) W-R based models. Line-blanketing is still clearly important in this region, although the departure is less profound. This region is particularly important for the formation of nebula lines, since many atomic species have ionization edges here, such as O$^+$, S$^{++}$ and Ar$^+$ (see Chapter 5). Here the SB2002 model acts much more like the LTE models at early years due to the line-blanketing. At approximately 4.5 Myr, the models almost converge. This is due the domination of WC stars, since they are hotter, and thus have more transparent winds, so line blanketing is less efficient in this domain. By 6 Myr, the late WC population dominates, this time being produced mainly by lower mass stars which spend only a short time as WN stars before evolving to WC star (c.f. Table 4.2). These are again cooler, but metal rich and thus the CMFGEN line-blanketing is more efficient. For observational evidence of the advantages of a CMFGEN models the reader should refer to Chapter 3.
Figure 4.13: The evolution of the photon luminosity for ages of 1-8 Myr in the ionizing continua of hydrogen ($\log Q_0$), He I ($\log Q_1$) and He II ($\log Q_2$) at 0.2 $Z_\odot$ for an instantaneous burst at time steps of 0.1 Myr. The new models (black, solid) are compared to the ionizing fluxes of: Leitherer et al. (1999) (red, dotted); Schaerer & Vacca (1998) (blue, dashed)
The $Q_0$ plot shows little difference between all three sets of models, although the SB99 models give a slightly lower flux during O star domination. Theoretically, the difference should be almost undetectable, since line-blanketing is less dominant in this region as the neutral and singly ionized atoms that would dominate this region in terms of blanketing do not exist in hot star atmospheres. The difference arises due to the use of high gravity ($\log g = 5.0$) O star models in the SB99 case, even for supergiants. These models are not really adequate to represent a supergiant (or in some high temperature cases even a dwarf star), as a lack of flux is observed in these models due to the smaller assumed radius at high gravity.

General Trend with Metallicity

The appearance and underlying characteristics of a starburst depend heavily on metallicity. The new SB2002 models attempt to give a genuine idea of the changes that are likely to be seen in different chemical environments. Figures 4.14 & 4.15 show the $\log(Q_1/Q_0)$ and $\log(Q_2/Q_0)$ ratios (defined in Equation 2.17 in Chapter 2) for all 5 metallicities which compare the general trends that are experienced as starbursts move from high to low metallicity. Considering first, Figure 4.14 showing $\log(Q_1/Q_0)$ as one moves from the top of the figure (2.0 $Z_\odot$), we can see the W-R phase becomes less and less significant for all the models. This is due to the raising of the upper mass cutoff for W-Rs combined with lower density winds, which also has the effect of shortening the W-R phase, as can be seen throughout the plot. However, we see one other important trend in this region of the starburst's evolution, due to the inclusion of the new models. Although there is a quantifiable difference between the SB2002 models and the others during the W-R phase at 2.0 $Z_\odot$, this reduces as the metallicity decreases, to a point when at 0.2 $Z_\odot$ the difference is minimal. At this metallicity and lower, however, another factor comes into play. This is best seen in the very low metallicity 0.05 $Z_\odot$ models. Late O and early B stars start to add significantly to the continuum, since they spend much longer on the MS at low metallicity due to their reduced mass-loss rates. We can see that the SB99 models now predict a lower ratio than the other models, due to the problem with high gravity supergiants mentioned above. Generally however, the difference in the predicted ratio for the SB2002 model is due to the effect of the atomic lines in the atmosphere models. Since wind density effects reduce with metallicity, we see a convergence towards the SV98 and SB99 models at low
Figure 4.14: The time evolution of the hardness of the ionizing flux as shown by \( \log(Q_1/Q_0) \) at five metallicities for an instantaneous burst over 1–8 Myr. The new models (black, solid) are compared to the ionizing fluxes of: SB99 (red, dotted); SV98 (blue, dashed).
Figure 4.15: The time evolution of the hardness of the ionizing flux as shown by 
$log(Q_2/Q_0)$ at five metallicities for an instantaneous burst over 1–8 Myr. The new 
models (black, solid) are compared to the ionizing fluxes of: SB99 (red, dotted); 
SV98 (blue, dashed).
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metallicity.

If we now look at Figure 4.15, depicting the log(Q₂/Q₀), ratio we see a much more exaggerated effect. Considering the SB2002 model at 2.0 \( Z_\odot \) we now only see a small contribution to the ratio at the very earliest times. This lack of flux has important overtures in the prediction of nebular He \textsc{ii} \( \lambda 4686 \) flux (see §4.5). Again as we move further down the plot, we see the contribution returning at lower metallicity. Even at 0.05 \( Z_\odot \) there is some difference between the \textsc{cmfgen} and Schmutz \textit{et al.} (1992) models. An important feature to note is that of the secondary peak in the 0.4 \( Z_\odot \) plot. This peak seems to cross the other models, indicating a greater ratio for a short period. The peak is due to the final lower mass generation of WNE stars, which have winds which have become transparent due to the lack of He \textsc{ii} recombination somewhere in the outflow. If one considers Figure 2.15 (Chapter 2) it can be seen that this is due to the use of the 100kK and 120kK WN models. Since no interpolation can be performed between transparent and non-transparent wind models, the ionizing flux simply jumps to the higher level with the use of the hotter model.

4.4.2 Continuous Star formation

Continuous star formation is also very important, since instantaneous starbursts are often far too idealised to properly represent a stellar population. Metal rich galaxies are consistent with extended bursts, which may last as long as 50-100Myr (Mouhcine & Contini (2002); Kilgard \textit{et al.} (2002); Meurer (1999)). Figure 4.16 shows the \( Z_\odot \) metallicity continuous star formation SED for the \( 1 M_\odot \) yr\(^{-1} \) and the same IMF parameters as the instantaneous case. Here we see the continuum level rise as the number of stars increases over time. The effect of the W-R population is most obvious between 3 Myr when the W-R population first appears and 10 Myr, where the W-R population is in equilibrium. The SB2002 models show a very small increase in ionizing flux below 228 Å as characterised by the small increase in flux at the absorption feature at log \( \lambda = 2.23 \) (170 Å ). The higher level continuum tail at 10 Myr is due to a dominant red supergiant population. The SB99 model behaves, as expected very differently at UV wavelengths, since the W-R population totally dominates, contributing significant flux below 228 Å at all times after the beginning of the first population starts its W-R phase. At a metallicity of 0.2 \( Z_\odot \) (Figure 4.17) we see the new O star models giving an increased flux in the He \textsc{ii} continuum due to the reduction in wind density. The supergiant contribution can be
distinguished from the main sequence population at 3 Myr, as the continuum is slightly higher in the far UV. Although we expected the supergiant models to have a deficit of flux at \( Z_\odot \) from the \( Q \) plots in Chapter 2, the metallicity scaling in the wind has made the wind more transparent, and we now see a positive contribution. We can see again that the first population of W-R stars at 4 Myr now starts to converge to that of the Schmutz et al. (1992) grid due to the lower wind density of the models.

If we now consider the ionizing properties of the continuous burst for the same two metallicities (Figs. 4.18 and 4.19), we can see the time evolution of the three models with more accuracy. For a \( Z_\odot \) burst we see that the \( Q_0 \) values for all three models are extremely similar. In particular, the SV98 and SB2002 models have almost identical continuum levels, although we can see the difference due to the LTE supergiant gravities in the SB99 model. In the \( Q_1 \) plot, we see that the CoStar O star models are providing a significant increase in flux. We know it must be these models, as the SB99 and SV98 W-R models are identical. A careful look at the differences between SB2002 and the SB99 work, shows that the LTE models give a larger flux level at early times.

Turning to Fig. 2.3, we can see that the deficit is due to the dominance of a 30 to 35kK population. The SB2002 model \( Q_1 \) value becomes larger at roughly 2.5Myr due to the influence of the giant and supergiant WM-basic population. The far UV (\( Q_2 \) bottom, Figure 4.19) plot as always shows the largest difference in the models, and illustrates the differences between the two W-R atmosphere grids. As the W-R phase starts, we see a large jump in \( Q_2 \) for both the SB99 and SV98 models. The SB2002 model shows no extra contribution from the W-R population as the dense winds have He\(^+\) recombination, thus block the He\(\text{II}\) continuum at this metallicity. This is a very different prediction to earlier work, as we can see that the O stars are the only contributors to the He\(\text{II}\) continuum, contrary to previous ideas. It is important to point out that a solar metallicity WN model with a transparent wind does exist, but the adopted temperature coupling to the evolutionary tracks inhibits the use of this model. The highest temperature WN model that is chosen has a temperature of around 105kK, and does not have significant He\(\text{II}\) ionizing flux.

Turning to the ratios of \( \log(Q_1/Q_0) \) and \( \log(Q_2/Q_0) \) illustrated in Figs. 4.20 and 4.21 we can see the metallicity dependence of the models. Plotting these ratios is a good way of exploring the evolution of nebular lines that may be produced around a starburst. If
Figure 4.16: Spectral energy distributions of the new models (top) compared with SB99 (bottom) at time intervals of 1, 3, 4, 5 and 10 Myr for a continuous burst at $Z_\odot$. 
Figure 4.17: Spectral energy distributions of the new models (top) compared with SB99 (bottom) at time intervals of 1, 3, 4, 5 and 10 Myr for a continuous burst at $0.2\ Z_{\odot}$. 
we look from high to low metallicity (top to bottom on the plot), we see that the SB99 has a higher ratio than the SB2002 model again due to the Schmutz et al. (1992) W-R models. Looking at the SV98 model, we can see that the O star models make a significant contribution, since the CoStar models have a harder ionizing continuum at all epochs. At 0.2 $Z_\odot$, the SB2002 model has almost converged with the SB99 model at later ages, when the steady state solution of the starburst has been reached. This is again due to the CMFGEN metallicity scaling due to the lower wind density at low metallicity. The SV98 model only starts to converge with the other models at 0.05 $Z_\odot$ however, since it underestimates line blanketing (Chapter 2 §2.2) which is more rigorously accounted for in both the other two codes.

The log($Q_2/Q_0$) plot (Fig. 4.21) illustrates the SB2002 trend towards the SV98 model at low metallicity, mimicking the general shape of the SV98 plot. However the SB2002 model still exhibits a deficit at extremely low metallicity in the steady state solution, when we may expect transparent W-R winds. The new CMFGEN models do have significant metals at low metallicity as they assume that the products of helium burning are present in a W-R wind, accounting for the softer continuum, mainly due to the cooling effects of C and O allowing for recombination of He$^+$. Even at early epochs (1-2Myr) where no W-R stars are present, a difference between the CoStar models and WM-basic models can clearly be seen. This is probably due to the differences in line blanketing deep in the stellar wind (see Chapter 2, §2.2). At high metallicity, the SB2002 models show a softer flux when W-R stars are present, than at early epochs dominated by O stars. Comparing this with the low metallicity environment, where a hardening of the continuum occurs during the W-R phase by $\sim$ a factor of 10, we conclude that W-R winds are far more sensitive to metallicity and hence wind density than O stars due to cooling by C and O.

4.5 The Nebular He $\text{II}$ Line

Nebular He $\text{II}$ $\lambda$ 4686 has long been associated with H $\text{II}$ regions surrounding starbursts. The existence of such a line associated with young stellar populations has been a source of much debate, as the energy required to create the optical line exceeds 4 Rydbergs. Ionizing energy this hard is not expected to be produced by O stars, since they are too cool to
Figure 4.18: The evolution of the photon luminosity for ages of 1–10 Myr in the ionizing continua of hydrogen (log $Q_0$), He I (log $Q_1$) and He II (log $Q_2$) at $Z_\odot$ for a continuous burst at time steps of 0.1 Myr. The new models (black) are compared to the ionizing fluxes of: SB99 (red); SV98 (blue).
Figure 4.19: The evolution of the photon luminosity for ages of 1–10 Myr in the ionizing continua of hydrogen (log $Q_0$), He I (log $Q_1$) and He II (log $Q_2$) at 0.2 $Z_\odot$ for a continuous burst at time steps of 0.1 Myr. The new models (black) are compared to the ionizing fluxes of: SB99 (red); SV98 (blue).
Figure 4.20: The time evolution of the hardness of the ionizing flux as shown by the \( \log(Q_1/Q_0) \) at five metallicities for a continuous burst over 1–10 Myr. The new models (black) are compared to the ionizing fluxes of: SB99 (red); SV98 (blue).
Figure 4.21: The time evolution of the hardness of the ionizing flux as shown by the $\log(Q_z/Q_0)$ at five metallicities for a continuous burst over 1–10 Myr. The new models (black, solid) are compared to the ionizing fluxes of: SB99 (red, dotted); SV98 (blue, dashed).
produce significant amounts (Garrett et al. 1991). W-R stars have instead been proposed as the ionization source powering this line because of their high effective temperatures and presence of broad stellar He II in their winds. The work of SV98 seems to prove this theory, as they predict significant amounts of nebular He II during the W-R phase of a starburst. Here we re-evaluate that claim with the use of the new SB2002 model with the grids of realistic ionizing fluxes.

### 4.5.1 Model comparisons of He II Flux

Instead of using a full photoionization model for calculating the nebular He II flux, we use the fact that the nebular line strength is proportional to the level of flux in the He$^+$ ionizing continuum ($Q_2$). A full use of the CLOUDY photoionization model is applied to SB2002 in Chapter 5. Using the equalities defined in SV98, we define the $H\beta$ flux:

\[
L(H\beta) = 4.76 \times 10^{-13} f_\gamma Q_0
\]  

(4.4)

Where $f_\gamma$ is the fraction of ionizing photons absorbed by the nebula and is assumed to be 1 in an ionization-bounded nebula. The He II $\lambda$ 4686 flux can be defined in a similar way:

\[
L(\lambda 4686) = 1.02 \times 10^{-12} f_\gamma Q_2
\]  

(4.5)

To allow us to compare the relative intensity of He II $\lambda$ 4686 to $H\beta$ we simply take the quotient:

\[
I(\lambda 4686)/I(H\beta) = 2.14 \times Q_2/Q_0
\]  

(4.6)

We also assume that for the He II line to be detected, it must be at least one percent of the nebular $H\beta$ line intensity for confident detection above the spectral noise level.

Figure 4.22 shows the predictions of the SB2002 and SV98 models. The SV98 model clearly predicts that there will be significant nebular He II line production during the W-R phase of all starburst models, independent of metallicity. During the 0.4 $Z_\odot$ burst the He II line intensity reaches 10 percent of the $H\beta$ line. The new models show a different picture, mainly due to the inclusion of a more complex chemistry in the new CMFGEN atmospheres. These models predict no nebular He II production at any metallicity. The 0.2 $Z_\odot$ model comes closest to the 1 percent detection limit, and could exceed this with an altered IMF, with a higher upper mass limit or shallower IMF slope.
Figure 4.22: Predictions of nebular He II line intensities for SB2002 (top) and SV98 (bottom). All are instantaneous burst models and have \( M_{up} = 100M_\odot \), \( M_{low} = 1M_\odot \) and a power law slope of \( \alpha = 2.35 \). The dotted line shows the assumed detection limit for He II \( \lambda 4686 = 1 \) percent H\( \beta \). The large difference is mainly due to the inclusion of the CMFGEN line blanketed W-R models.
It seems that nebular He II cannot be produced from the ionizing continua of a young stellar population. Guseva et al. (2000) look at the nebular properties of 30 H II galaxies and find that they do indeed detect significant He II fluxes in agreement with SV98. Many of the sample galaxies, however, have nebular He II emission but have no sign of the W-R star broad stellar lines and moreover nebular He II λ 4686 is only found in galaxies with $Z < 0.2 Z_\odot$. They conclude that the ionizing continuum from W-R stars cannot be the only source of the He II line and suggest that other mechanisms such as collisional shocks could provide the answer. They also note that the work of SV98 shows that the strongest He II fluxes would be predicted at metallicities of solar and above, in clear contradiction to the observations.

4.6 Rate of Mass and Energy Return to the ISM

Super-winds from starburst galaxies have long been thought to be the consequence of massive star winds combined with supernovae (Heckman et al. 1990). These winds have been suggested as the process for heating the ISM (De Young 1978), which in turn could be an important mechanism to explain why runaway star formation is suppressed (Dekel & Silk 1986; Larson 1974; White & Rees 1978). Information on the feedback properties of starbursts is extremely important to the success of cosmological models. As we have updated the mass-loss rates and ionizing properties of all stars that dwell in the upper HR Diagram, it is important to compare these new data to that of the original SB99 model. Since the evolutionary tracks have been untouched in SB2002, the energy released via supernovae will remain unchanged, so we concentrate on the nuclear burning massive star population.

4.6.1 Mass–Loss from Stellar Winds

Fig. 4.23 shows the predicted mass–loss rate of the SB99 and SB2002 models for an instantaneous model with the IMF parameters defined in §4.4.1. The SB99 model is calculated using the following prescription for the mass–loss and $v_\infty$ of the various stellar components:

- Mass–loss from O stars:

The mass-loss rates and terminal velocities are from a multidimensional fit performed
Figure 4.23: The predicted mass-loss rates from the SB2002 (solid, black) and SB99 (dashed, red) models as a function of age and metallicity.
by Leitherer et al. (1992) on a grid of models based on a radiation-hydrodynamics code (Leitherer et al. 1989) using the formalism of Castor et al. (1975).

- **Mass–loss from cool stars:**
  
The relation of Reimers (1975) is used to define the mass–loss from Red Supergiants. A generic value of 30 km s\(^{-1}\) is assumed for the \(v_\infty\) of all stars (Drake 1986).

- **Mass–loss from LBVs**
  
  LBVs are defined on the Meynet et al. (1994) evolutionary tracks as having \(\log \dot{M} > -3.5\) \(M_\odot\) yr\(^{-1}\) and an effective temperature of between 5600K and 25000K. The adopted mass–loss is given by Lamers (1989) as \(\log \dot{M} > -3.9\) \(M_\odot\) yr\(^{-1}\). A value for \(v_\infty\) of 200 km s\(^{-1}\) is assumed for all LBVs (Lamers 1989).

- **Mass–loss from W-R’s**
  
  A W-R star is defined such that its effective temperature is greater than 25kK, the surface hydrogen abundance of less than 40 per cent and the stellar mass is greater than that of the minimum W-R mass. The average values of van der Hucht et al. (1986) are used to define a mass–loss for each of the WNL, WNE, WCL, WCE and WO subtypes. The terminal velocity values of Prinja et al. (1990) are used to supply an average for each of the subtypes defined above.

All terminal velocities and mass–loss rates are scaled with metallicity using \(v_\infty \propto Z^{0.13}\) and \(\dot{M} \propto Z^{0.80}\) except for W-R stars which are not scaled with metallicity.

Our approach to the mass return determination differs in the following way: All O and W-R star mass-loss rates and terminal velocities are taken from the values used in our grids, with the method used to match the evolutionary stellar component being exactly the same as for the atmosphere models, described above in §4.3. In our case, all mass-loss rates are scaled using the above prescription including W-R stars.

If we turn to Figure 4.23, we see the differences between the models. At early times, the SB2002 O star models have a greater mass-loss than the SB99 model for all metallicities. The W-R stars return less mass than the SB99 model however, for all metallicities with an increasing difference at lower metallicity due to our scaling. If we consider the 2.0 \(Z_\odot\) model, we see agreement at ages after 3 Myr. This is due to a trade off between higher O star mass–loss rates and lower W-R mass-loss rates of the newer models. At 1.0 and 0.4 \(Z_\odot\) the O stars no longer lose enough mass to maintain this equality, so we see a deficit in the mass return from the new SB2002 models.
However, in the SB2002 model at 0.2 and 0.05 $Z_\odot$, the underlying late O and early B star population contribute considerably to the mass-loss rate of the system, swamping the SB99 OB population at 5 – 10 Myr. Why do the two populations exhibit such different behaviour? Figure 4.24 shows the comparison of the two models’ mass-loss rates. If the mass-loss rates agree, the points should lie on the solid diagonal line. Stars with high mass-loss, especially the supergiant population lie in a tight concentration just below the diagonal line indicating a good agreement between the two rates, although the mass-loss is slightly higher in the SB2002 determination. The most striking trend is at low mass-loss, where the dwarf population has far higher values than Leitherer et al. (1992). Observationally, even early B stars have significant mass-loss properties, and a difference of a factor of $10^{2.5}$ can be seen between the two values at the low end. It is this difference which is causing the OB population in the SB2002 model to give a much larger prediction at low metallicity. These stars now evolve more slowly due to the low metallicity, and are still on the main sequence after the W-R phase has finished (rather than becoming supergiants). At high metallicity these stars become supergiants before the end of the W-R phase.

We can now see what effect the new mass-loss determination has on the overall power output from the SB2002 model. Figure 4.25 shows a comparison between the SB2002 and SB99 models for all five metallicities. In this figure, we see that there is indeed a tradeoff between the two models. The SB2002 model has higher O and early B star power output due to the increased mass-loss which offsets the difference in W-R star mass-loss rates. At very high metallicity, where W-R stars dominate the mass return, there is only a small difference between the two models as both W-R grids have high wind densities. At the other extreme, although there is a large difference in W-R star winds, the greater B star mass-loss rates in the SB2002 model dominate the power output. We see that now W-R winds no longer dominate the contribution to stellar wind power at low $Z$.

Overall in this chapter, we see that the new models predict a very different role for the stellar populations. W-R stars have become a far less dominant ionizing source, mostly due to the trade off between their rarity at low $Z$ and their weak ionizing behaviour at higher $Z$. They have ionizing properties much more like O stars, due to the inclusion of line blanketing. W-R stars now contribute less to the mass return than previously thought, due to the metallicity scaling of the wind and the use of contemporary mass-loss rates which include clumping factors. W-R stars are far more sensitive to metallicity changes.
Figure 4.24: A comparison of the mass-loss rates predicted by the theoretical prediction of Leitherer et al. (1992) and that of the empirical relationship presented here.
Figure 4.25: The predicted energy return from both the SB99 and SB2002 models for 5 metallicities. The solid line represents the SB2002 values and the dashed the SB99 model.
than their younger relatives, although these predictions show that their signatures are harder to find than previously thought. In the next chapter, we explore the observational comparison with the new models and look at the spectral indications of an underlying W-R population.
Chapter 5

Observational Comparisons

Theoretical work needs to be validated by comparing its predictions with that of observation. Photoionization models employing single source continua have been used to model nebulae for 20 years, since the pioneering work of Osterbrock (1974). This method was not only used to model nebulae around single stars, but as a way of comprehending the environment around starbursts, associations and clusters. Even though analysis of the nebular environment could be attempted via this technique, information on the underlying stellar population is lost due to a single star assumption. Recently, population synthesis has become reliable enough to be used as the underlying input for photoionization models, allowing modelling of the population content of a starburst or cluster (Leitherer & Heckman 1995; Stasińska & Schaerer 1997). This chapter outlines the methods of analysis of nebular emission lines via the use of the continuum outputs discussed in Chapter 4 and the use of the photoionization code CLOUDY. Using optical and infra-red data for a sample of H II regions, we attempt to obtain consistent model fits and information on the ionizing stellar populations. With this information we then apply a new technique, developed with the help of our new atmosphere grids, which can generate a synthetic spectrum in the optical of the “W-R bump” features sometimes observed in starburst galaxy spectra, indicating the existence of an underlying W-R population. We find that this feature can be used to gain information on the W-R subtypes within the starburst and the age and metallicity of the burst. §5.1 introduces a method to compare many H II regions at once to the new model predictions, §5.2 describes the new technique to synthesise the “W-R bump” features observed in the spectra of W-R galaxies, and §5.3 presents the first results of using a combined photoionization and spectral synthesis technique to parameterise five
CHAPTER 5. OBSERVATIONAL COMPARISONS

W-R galaxies.

5.1 H II Region Diagnostics

To illustrate the usefulness of the new atmospheres, we will first look at a sample of previously observed H II regions and compare them with the equivalent model predictions. In order to do this fairly, we must try to make the nebular line intensities as parameter independent as possible, so that no biases can occur. In this comparison we use two large data sets, that described in Bresolin et al. (1999) of a large sample of extragalactic H II regions in the optical region, and the mid-infrared ISO dataset of Giveon et al. (2002) for Galactic H II regions.

To complement the observed nebular intensities we calculated two simple photoionisation model grids, one with the new Starburst99 (SB2002) atmospheres (CMGEN for W-R stars and WM-basic for O stars) and that using Starburst99 including the CoStar atmospheres similar to that used in Schaerer & Vacca (1998) (from now on SV98). We used models of three metallicities, 0.2, 1.0 and 2.0 $Z_{\odot}$, to span the metallicity range of the data and modelled bursts in 1 Myr intervals up to 6 Myr. This created a grid of 36 unique models that could be compared with the data. Once the continua were generated, they were prepared as an input into the photoionization code CLOUDY as described in Chapter 3. The following parameters were used to define our model nebulae. The most simple model geometry was chosen, that of a plane-parallel assumption and an identical metallicity to the Starburst99 grids was assigned to be consistent with the continuum. A hydrogen density of 50 cm$^{-3}$ was also assumed and, although dust grains were not included in the model, the depletions listed in Table 5.1 were included (Dopita et al. 2000). Ionization parameters of log $U = -2$ and $-3$ were assumed (defined in Equation 5.3 in this subsection), since most H II regions in the Bresolin et al. (1999) dataset (which was primarily used for the comparison) lie in the region $-3.5 \leq \log U \leq -1.5$. CLOUDY can model far more complicated systems, including spherical shell nebulae, with dust and complex density structures. However, since a large number of H II regions are being modelled all with different properties, it was felt that the most simplistic approach would suffice.
Table 5.1: Assumed depletion factors for the CLOUDY model grid (Dopita et al. 2000)

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* 1.3 & 0.9 for 2 & 0.2 Z\(_\odot\) respectively

5.1.1 Optical datasets

The dataset of Bresolin et al. (1999) has enough information on the relative intensities of the emission lines to use an essentially parameterless quantity known as \(\eta\) defined in Vilchez & Pagel (1988):

\[
\eta = \frac{O^+ / O^{++}}{S^+ / S^{++}}
\]

The balance between ionization levels of an element can be written as (Osterbrock 1974):

\[
\frac{n(X^{i+1})}{n(X^i)} = \frac{U \epsilon}{\beta(X^i)} \frac{\int_{\nu(X^i)}^{\infty} F(\nu) \sigma^i(\nu) d\nu / h\nu}{\int_{\nu(H\alpha)}^{\infty} F(\nu) \sigma^i(\nu) d\nu / h\nu}
\]

where \(F(\nu)\) is the flux at \(\nu\), \(\sigma^i(\nu)\) is the cross section of the \(i^{th}\) ionization state. The top integral is defined in the range between the ionization edge of the \(i^{th}\) ion and \(\nu = \infty\) and the denominator is integrated between the \(H^0\) ionization edge and \(\nu = \infty\). \(\beta(X^i)\) is the recombination coefficient and \(U\), the ionization parameter, is defined as:

\[
U = k_\epsilon (Q \epsilon n_e)^{1/3} \equiv \frac{Q}{4\pi r_0^2 n_e c}
\]

where \(Q\) is the rate of production of Lyman continuum photons, \(\epsilon\) is the nebular filling factor, \(n_e\) is the electron density, \(r_0\) is the separation between the centre of the source of
ionizing radiation and the illuminated face of the cloud, \( c \) is the speed of light and \( k \) is a constant. Looking at these equations we can see that dividing one ionization fraction by another, we can remove the \( U \) dependence, and hence the dependence on the nebula filling factor and density. In a more realistic situation however, this does not remove the \( U \) dependence entirely, so in datasets where there is a scatter of \( U \) values, more than one \( U \) value must be considered. In the case of our models, where radius and density are kept constant, \( U \) can be thought of as a measure of the ionizing photons from the source continuum (i.e. \( U \propto Q^3 \)). We plot the observational quantity \( \eta' \) analogous to \( \eta \) defined as:

\[
\eta' = \frac{[O II] \lambda 3726,3729/[O III] \lambda 4959,5007}{[S II] \lambda 6717,6731/[S III] \lambda 9069,9539}
\]

against the \( R_{2.3} \) parameter defined as \(([O II] \lambda 3727 + [O III] \lambda 5007,4959)/H\beta\) which is approximately proportional to metallicity. This is due to the assumption that a nebula will have a large percentage of its oxygen in the ionized \( O^+ \) and \( O^{2+} \) states, therefore the intensity of these lines should change in a way that is roughly correlated to metallicity. The \( H II \) region data of Bresolin \textit{et al.} (1999) is presented in Table 5.2. This dataset consists of 96 extragalactic \( H II \) regions covering a metallicity range of \( 0.2 Z_\odot < Z < 2.0 Z_\odot \). These data were obtained between 1988 to 1992 on the Steward Observatory’s 2.3m telescope as part of a survey by Oey & Kennicutt (1993). These data were supplemented by several other datasets referenced in Bresolin \textit{et al.} (1999). Figures 5.1 and 5.2 present an \( R_{2.3} \) vs \( \eta' \) plot of the Bresolin \textit{et al.} (1999) data with the 12 SB2002 and SV98 models. Two \( \log U \) values are considered due to the scatter of the Bresolin \textit{et al.} (1999) data. The \( \eta' \) quantity has low values for hard ionizing continua and higher values for continua that are softer. This plot shows that the SB2002 models give a better fit to the data than the older SV98 models, which appear to be too hard in the ionizing continuum. The new models tend to go beyond the data for the higher metallicity models (three models extend just off the plot in Figure 5.1 and two in Figure 5.2. This indicates that the 2.0 \( Z_\odot \) models may be too soft in the far UV in fit the data, especially in the W-R phase. Figure 5.2 is the starburst analogue to Figure 3.3, where the single star source functions are replaced in favour of an output Starburst99 continuum. At early starburst ages (of order 1Myr) these models fall in the same loci, as demonstrated in Figure 5.3, due to the starburst’s population being dominated by and early O star population. Differences occur at later ages, especially at high metallicity, where the W-R star population dominates and the
Table 5.2: Optical spectra from Bresolin et al. 1999

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Notes: \(^a\) Relative to galaxy nucleus, positive in east and north directions, in arcseconds.
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Notes: 

- Relative to galaxy nucleus, positive in east and north directions, in arcseconds.
ionizing properties of the population changes significantly. For example at 2 and 3 Myr, the W-R and O supergiant populations cannot reproduce the hardness of the ionizing continuum produced in the single star models or the Bresolin et al. (1999) dataset.

5.1.2 IR datasets

To complement the optical dataset of Bresolin et al. (1999), we have obtained IR data in order to probe different regions of the predicted far UV of the SB2002 and SV98 models. We have used the mid-IR dataset of Giveon et al. (2002) from the Infrared Space Observatory Short-Wavelength Spectrometer (ISO-SWS). This dataset contains the fine structure emission line intensities for two ionization species each of Ne, S and Ar as well as the Brackett \( \alpha \) and \( \beta \) hydrogen lines for 112 Galactic H II regions and is presented in Table 5.3.

For this comparison, we define a “radiation softness” parameter \( \zeta' \) to be analogous to that of the \( \eta' \) parameter of Vilchez & Pagel (1988). Since there are only two elements with strong lines from two different ionization stages present in the optical spectral region, this is the only combination that can be defined. In the mid-IR, we are gifted with three line pairs, allowing us to define two analogous \( \eta' \) quantities, with slightly different properties. These parameters are defined as:

\[
\zeta_1 = \frac{S^{3+}/S^{2+}}{Ne^{2+}/Ne^+} \Rightarrow \zeta'_1 = \frac{[S IV] \lambda 10.5 \mu m/[S III] \lambda 18.7 \mu m}{[Ne III] \lambda 15.6 \mu m/[Ne II] \lambda 12.8 \mu m}
\]

\[
\zeta_2 = \frac{Ar^{2+}/Ar^+}{Ne^{2+}/Ne^+} \Rightarrow \zeta'_2 = \frac{[Ar III] \lambda 8.99 \mu m/[Ar II] \lambda 6.98 \mu m}{[Ne III] \lambda 15.6 \mu m/[Ne II] \lambda 12.8 \mu m}
\]

(5.5)

What is the exact relevance of the “softness parameters”? The answer lies in the source continuum. The intensity of a line depends on the number of excited species in the nebula which can recombine and release radiation. For this reason, the line intensity relies on the amount of flux beyond the ionizing potential of the atomic species. For example, the line intensity of the forbidden oxygen line \([O III]\) \( \lambda 5007 \) relies on the ionizing source’s flux just shortward of the \( O^+ \) continuum edge at 2.58 Rydberg where \( O^{++} \) is produced. If the continuum is negligible at this point, then there will be no source to ionize oxygen to a doubly ionized state and hence no observable \([O III]\) line. Figure 5.4 plots the ionization edges of some of the ions important for nebular line analysis. This plot shows which parts of the continuum we actually probe when we consider the softness parameters.
Figure 5.1: \( \eta' \) versus \( R_{2,3} \) plot with \( \log U = -2 \) of Bresolin et al. (1999) data against the Schaerer & Vacca (1998) and SB2002 models between 1 and 4 Myr for 0.2, 1.0 and 2.0 \( Z_{\odot} \).
Figure 5.2: $\eta'$ versus $R_{2,3}$ plot with $\log U = -3$ of Bresolin et al. (1999) data against the Schaerer & Vacca (1998) and SB2002 models between 1 and 4 Myr for 0.2, 1.0 and 2.0 $Z_\odot$. 
Figure 5.3: $\eta'$ versus $R_{2,3}$ plot with $\log U = -3$ comparison of 0.2, 1.0 and 2.0 $Z_\odot$ 45kK WM-basic models and 1Myr SB2002 models against the Bresolin et al. (1999) data.
Table 5.3: ISO Galactic HII regions of Giveon et al. (2002). The fluxes are dereddened.

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Notes: Some values omitted due to zero $\text{Br} \alpha$ or $\text{Br} \beta$. Fluxes are in units $\text{W cm}^{-2}$. Values for Extinction from Li & Draine (2001).  

*a*: Distance from the galactic centre.
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Notes: Some values omitted due to zero Br$\alpha$ or Br$\beta$. Fluxes are in units W cm$^{-2}$. Values for Extinction from Li & Draine (2001).

a: Distance from the galactic centre.
CHAPTER 5. OBSERVATIONAL COMPARISONS

Figure 5.4: The ionization edges of some of the important ions used in nebular emission line analysis. Emission lines are sensitive to the continuum of the ionizing source especially around the wavelengths of these edges. Plotted with these edges are the continua from the updated Starburst99 (solid line) and the old version including the pure helium W-R models of Schmutz et al. (1992) (dashed line).
Table 5.4: The dominant continuum edges for the three softness parameters, \( \eta' \), \( \zeta_1 \) and \( \zeta_2 \)

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<th>Dominant ionization edge of numerator</th>
<th>Dominant ionization edge (Rydbergs)</th>
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<td>( \zeta_1 )</td>
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<td>( Ne^+ \rightarrow Ne^{2+} ) (3.01 Ryd.)</td>
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<td>( \zeta_2 )</td>
<td>( Ar^+ \rightarrow Ar^{2+} ) (2.03 Ryd.)</td>
<td>( Ne^+ \rightarrow Ne^{2+} ) (3.01 Ryd.)</td>
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If we first consider \( \zeta_1 \), which is essentially dependent on the ionizing flux above the \( S^{2+} \) and \( Ne^+ \) ionizing edges, we see that it probes a very hard region of the source continuum, because of the high ionizing potential of \( Ne^+ \). A soft continuum will yield little \( Ne \) and hence the \( \zeta_1 \) value will be large. Generally the softness parameter will be most sensitive to the line which is produced from the ion with the highest ionization potential. Only in extreme cases (i.e. Br2, Chapter 3) will the flux peak in the neutral helium ionizing continuum.

The parameter \( \zeta_2 \) is like \( \zeta_1 \) parameter although it reaches down to lower energies. It should also be noted that with all the IR softness parameters, the important ionization edges reside in the energy range \( 1.0 < E < 4.04 \) Ryd. This means that there is no contribution from a state above the Lyman edge. The optical \( \eta' \) value does however include \( S^0 \), shown in Figure 5.4 which is above the Lyman break at 0.76 Rydberg. This gives a large difference in the continuum levels that only the \( Ne^+ \) continuum edge can match. One other pair of IR lines exist, that of:

\[
\zeta_2 = \frac{[SIV]\lambda10.5\mu m/[SIII]\lambda18.7\mu m}{[ArIII]\lambda8.99\mu m/[ArII]\lambda6.98\mu m}
\]  
(5.6)

This parameter depends on the flux levels at the \( S^{2+} \) and \( Ar^+ \) continuum edges (2.56 and 2.03 Rydbergs), and is therefore the softest of all the parameters due to the position of the \( S^{2+} \) ionization edge at 2.56 Rybergs. Considering Figure 5.4 we can see that the new SB2002 continua can have a higher flux level over this region. A problem with this parameter is illustrated in this figure, as the difference in flux at the lowest continuum edge compared to the highest can be very small and can be heavily influenced by the low resolution caused by the SB99 atmosphere wavelength grid, which can change the continuum level by 0.5 dex. We have decided to omit this parameter from our discussions as it is unlikely to yield any useful information about the shape of the predicted continua.
In order to gain some measure of metallicity to plot the data against we use a parameter analogous to that of $R_{2,3}$, where we define $R_{2,3}^{IR}$ as either ([Ne II] + [Ne III]/Br $\alpha$) or ([Ar II] + [Ar III])/Br $\alpha$ assuming that all neon or argon atoms are in these ionic states within the nebula, so that the total intensity represents the contribution of all ions and is therefore correlated with metallicity. The sulphur line $R_{2,3}^{IR}$ ([S III] + [S IV])/Br $\alpha$ cannot be relied upon because of significant amounts of S$^+$ being present in these nebula environments (20 percent or more), which can be seen in the optical as [S II]. Tests with CLOUDY find that neon should follow the nebular abundance most closely, as under 1 percent of neon is in a form other than Ne$^+$ and Ne$^{2+}$ (CLOUDY model with SB2002 3Myr $Z_{\odot}$ input continua, $T_e$ = 9000K, $N_e$=800cm$^{-3}$). Argon is less useful as 5 percent of the total argon can be in the form of Ar$^{3+}$. Figures 5.5 and 5.6 show the two model grids of SB2002 and SV98 plotted over the ISO H II region data. Each of the two models are plotted for 3 metallicities (0.2, 1.0 and 2.0 $Z_{\odot}$) and over 5 epochs (1, 2, 3, 4 and 5 Myr) with log $U = -2$. If we first consider Figure 5.5, the $\zeta_1$ versus $R_{2,3}^{IR}$ plot, we see that our new updated atmospheres fit the data more accurately than SV98, due to a softer continuum. The sample data is best fit by the softer 1.0 and 2.0 $Z_{\odot}$ metallicity models, rather than the harder 0.2 $Z_{\odot}$ models, because of the lower wind density assumed by metallicity scaling (see Chapter 2). In fact, the SV98 models do not overlap the data at all. Although the new SB2002 models only fit the lower part of the data, this is not recognised to be a problem of the new models since it must be noted that some of the H II regions in this sample only have one or a few stars associated with them. These are unlikely to be fitted well by a starburst model.

### Table 5.5: Summary of nebular diagnostic parameters

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<th>Symbol</th>
<th>Definition</th>
<th>Explanation</th>
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<td>$\zeta_2$</td>
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<td>$R_{2,3}^{IR}$</td>
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Figure 5.5: A $\zeta'$ versus $R_{3,3}^{LP}$ comparison of the CoStar and SB2002 Starburst99 models for $\log U = -2$ and ISO HII region data. The Starburst99 models are plotted for 3 metallicities (0.2, 1.0 and 2.0 $Z_\odot$) and over 5 epochs (1, 2, 3, 4 and 5 Myr).
Figure 5.6: A $c'_2$ versus $R^{HI}_{2,3}$ comparison of the CoStar and SB2002 Starburst99 models for log $U = -2$ and ISO Hii region data. The Starburst99 models are plotted for 3 metallicities (0.2, 1.0 and 2.0 $Z_\odot$) and over 5 epochs (1, 2, 3, 4 and 5 Myr).
and will have on average softer ionizing source continua, since they will mostly represent cooler O and B stars (as these are preferentially created over more extreme stars due to the slope of the IMF and the small number of stars involved), whereas a starburst population will always contain the most extreme ionizing stars in its stellar inventory, although this does depend on age. In these cases, nebular modelling using a single star source function approximation, or alternatively a starburst continuum are inadequate. A more accurate method would be to generate a small stellar population via a Monti-Carlo technique and fit individual models, which is beyond the scope of this Thesis.

The $\zeta_2$ versus $R_{4.5}^{15\mu m}$ plot (Fig.5.6) shows the argon to neon ratio. The two sets of models make very different predictions in their $\zeta_2$ values. The SB2002 models, in particular the solar metallicity models, fit the data quite well, being coincident with a large number of data points. The 2.0 $Z_\odot$ points seem to be too soft during the early W-R phase to model the IR data. The SV98 model is shifted much further down the plot, indicating a much harder ionizing source spectrum at all ages. Comparing the different $\zeta'$ plots we can see that the $\zeta'_2$ definition gives the largest shift between the two models. $\zeta'_2$ is the hardest probe of the ionizing continuum, where the difference between the two model sets is greatest, due to the very hard Schmutz et al. (1992) W-R models and CoStar O star models.

5.1.3 Conclusion from H II Analysis

Overall we see that the new models match the observations much better than the previous generation of atmospheric models. In the optical we see a large improvement over the SV98 models, where the SB2002 models now show a correct match to the data, being much softer over the range probed by $\eta'$ of 1.72 to 2.58 Rydbergs.

This conclusion is confirmed in the mid-IR where we see the new softer continuum matching the data far more satisfactorily than the SV98 models. $\zeta'_1$ illustrates this point well, with the new models overlapping the harder H II region data points. If we consider the SV98 predictions for the same data, we see that the continua are too hard to coincide with any of the observations. Although $\zeta''_2$ is less convincing because both model grids can reproduce the $\zeta'_2$ values of at least part of the Giveon et al. (2002) dataset, there is a reasonable amount of evidence to suggest that the new models predict continua soft enough to agree with observation.
5.2 W-R Feature Spectral Synthesis

Although a bulk consideration of the ionization properties of the new starburst continua is important, it is also worthwhile to model W-R galaxy spectra in a more rigorous way. In this section we model five observations of starburst galaxies with underlying W-R spectral features. First we use the photoionization code CLOUDY to model the nebular emission lines to determine the age of the starburst, assuming an instantaneous case. Using the parameters determined in the photoionization modelling, we then use a new technique, developed for the use of the new stellar atmospheres. We use Starburst99 to sum the contribution of all stars in the optical region between 3700 and 6500 Å to simulate the W-R features in the starburst spectra. These features can then be compared to the observational data with respect to the shape and intensity of the so called “W-R” bumps, which appear around 4686 and 5808 Å.

5.2.1 The Spectral Synthesis Code

To complement the “traditional” photoionization modelling of the nebula surrounding a starburst, we have utilised the raw W-R CMFGEN stellar atmospheres in order to create a synthetic spectrum of the same spectral region. This region is of particular interest due to the existence of observed W-R bump spectral features distinguishable above the underlying starburst continuum. The first object in which a W-R feature was recognised was He 2-10 by Allen et al. (1976). Since that time, galaxies which exhibit these broad features, observable at \( \lambda 4686 \) and \( \lambda 5808 \), are known as Wolf-Rayet galaxies to signify that they contain W-R stars. This phrase was first coined by Conti (1991).

To construct the synthetic spectrum we used the raw CMFGEN atmospheres to represent the W-R stars. A grid of synthetic spectra was created covering all SB99 metallicities and all W-R temperatures by re-binning the original fluxes onto a 1 Å grid over the wavelength range of 3700 to 6500 Å. This resolution was designed to be higher than that of a typical W-R galaxy observation, but low enough so that the memory required to create the grid was not restrictive. These models are presented in Figures 5.7, 5.8, 5.9, 5.10 and 5.11. We plot a cool and a hot WN star (WNL and WNE) of temperatures 40 and 80kK with WCL and WCE models at 60 and 100kK. We note at very low metallicity, the W-R
Figure 5.7: The $Z_{\odot}$ W-R models from the CMFGEN grid. Marked on the plot are WN 40 and 80kK models and WC 60 and 100kK models.
Figure 5.8: The $0.4 \, Z_\odot$ W-R models from the CMFGEN grid. Marked on the plot WN 40 and 80kK models and WC 60 and 100kK models.
Figure 5.9: The $0.2 Z_\odot$ W-R models from the CMFGEN grid. Marked on the plot are WN 40 and 80kK models and WC 60 and 100kK models.
Figure 5.10: The $0.05\,\text{Z}_{\odot}$ W-R models from the \textsc{cmfgen} grid. Marked on the plot are WN 40 and 80kK models and WC 60 and 100kK models.
Figure 5.11: The $2.0 \ Z_\odot$ W-R models from the CMFGEN grid. Marked on the plot are WN 40 and 80kK models and WC 60 and 100kK models.
bumps are extremely weak compared to the moderate and high metallicity models, due to the comparatively weak wind. The $\text{N III} \lambda 4640$ line in the late WN model is heavily affected by wind density (Abbott, private communication), and therefore is an indication of metallicity if mass-loss scaling is valid. In the WCE models, we see the balance between the $\text{C III}$ and $\text{C IV}$ lines changing with metallicity. The strong $\text{C IV}$ is indicative of a lower density wind, where the ionization balance favours the $\text{C}^{3+}$ ion over the $\text{C}^{2+}$. At higher metallicities we see cooling effects which alter this ratio in favour of the $\text{C}^{2+}$ state.

However, W-R stars are not the only contributors to the blue W-R bump spectral features. Schmutz & Vacca (1999) noticed that in some W-R features, the nitrogen emission was too strong to be just due to W-R stars. They conclude that evolved O supergiants must also contribute to the feature. In order to create an accurate optical representation of O star supergiants, the WM-basic models could not be used to produce the emission line features, but were used to renormalise the O star continua (c.f. Figure 5.12). These models were calculated to represent the ionizing fluxes produced by O stars, as W-R synthesis was not foreseen when these models were originally calculated, therefore spectral lines were not calculated at wavelengths long-ward of 1200 Å. Tests of WM-basic also indicate that it performs poorly in reproducing W-R emission lines, so a new grid of O supergiant models was instead calculated using CMFGEN, modelling only those stars which have a significant contribution to the W-R bump features. We ran a grid of O supergiant models between 2.0 and 0.2 $Z_\odot$, from 51 to 30kK, since models outside this parameter space have low density winds and yield insignificant contributions to the W-R feature. A total of 32 models were calculated with parameters equivalent to those defined in Table 2.3, Chapter 2.

The general method of the calculation of the synthetic spectrums follows that of the continuum calculation described in Chapter 4, where various stellar components are added to create a final synthetic spectra. Within this framework however, we found that we had to be careful in describing each component, since most continua (i.e. M to B stars) were still covered by only 1221 points. These models had to be re-binned during the calculation so that they could be added to the starburst spectra. The alternate method of using the 1221 points in the grid was found to be unsatisfactory, since this resolution was too low to accurately represent the shape of the W-R bumps at earlier epochs, where the lines produced are narrow, coming from O supergiants and from WNL stars. The flowchart in Figure 5.12 shows a representation of the basic structure of the subroutine which creates
the synthetic spectra. The subroutine steps through all the components of the starburst, binned by mass as in the main routine. It again determines the type of star that the component represents.

If the star is a cool, low or intermediate mass star, a Lejeune atmosphere is used to represent it, by calling the relevant subroutine used in the continuum calculation. This atmosphere is gridded on the low resolution Starburst99 wavelength grid, so the atmosphere must be re-gridded to fit on the high resolution grid consisting of 2800 points over the wavelength range of between 3700 Å to 6500 Å. If the star is an O star, the surface gravity and temperature determine whether it is hot enough and evolved enough to be represented by the O supergiant grid. If the star has an effective temperature above 30kK, and the log surface gravity is less than 3.5, or if \( T_{\text{eff}} > 40 \text{kK} \) and the log surface gravity is less than 3.88, then the star is selected as an O supergiant. This avoids contamination by the giant population. The lowest metallicity case does not have an associated \texttt{CMFGEN} grid since test cases showed that the wind features were too weak to create any contribution to the integrated starburst spectrum. If the \texttt{CMFGEN} grid was not chosen, the correct WM-basic grid was assigned in the way described in Chapter 2. This flux is then re-gridded to correspond to the higher resolution grid used in the subroutine. A supergiant model is associated with one of the high-resolution atmospheres, and normalised to the corresponding WM-basic supergiant atmosphere for flux conservation. A more rigorous normalisation procedure would use a Planck blackbody, and summation over the entire wavelength range of the continuum, but this is not possible since the high resolution grid is only defined over the 3700 Å to 6500 Å range.

The W-R star atmospheres are fitted in a similar way, although \texttt{CMFGEN} atmospheres are used to represent all W-R stars. To determine whether a star is a W-R star, the same criteria are used as in previous chapters, i.e. the surface hydrogen abundance must be less then 0.4, and the effective temperature must be greater then 25kK. WN and WC stars are differentiated by their surface chemistry as before (a WN star must have a non negligible nitrogen abundance (see Chapter 4)). All stars are normalised to their low resolution continuum counterparts as in the O star case.

Once an atmosphere has been assigned to a component, it is added to the total. This procedure is repeated until all components are accounted for. Once the total stellar component is calculated, the nebular component is included. This is crucial due to the nature of observations of W-R bumps, which are always observed with an associated
CHAPTER 5. OBSERVATIONAL COMPARISONS

What type of star?
- Other
- W-R
- O star

Does $m = M_{up}$?
- yes
- no

Choose correct Lejeune Atmosphere

Choose correct WM-basic Atmosphere

Choose hi-res Atmosphere

Write output

end

start

Read in $WN, WC, OIf$ Atmospheres

$m = M_{low}$

Is $T_{eff} > 30$ kK?
- yes
- no

Regrid atmosphere over correct spectral range

Normalise to low resolution atmospheric grid

$\tilde{\tau}_{tot} = \tilde{\tau}_{tot} + \tilde{\tau}_{m_1}$

Figure 5.12: A flowchart of the W-R bump spectral synthesis procedure
nebula. The continuum is added via a subroutine from the original SB99 code, which calculates the nebular continuum by taking into account the hydrogen and helium free–free and bound–free emission, and the two photon contribution. This contribution is re-gridded onto the high resolution grid and added to the stellar component. The resultant spectrum is written to an output file and the process is repeated for all required starburst ages.

5.2.2 Results of the Spectral Synthesis Code

It is well known that the W-R bump in a spectrum can look very different from starburst to starburst. Here we present the first results from modelling these features, looking at the variability of the profile between bursts of different ages and metallicities. We will concentrate on the instantaneous burst models in this section due to the fast changing nature of the feature. For a constant star forming rate the population will reach an equilibrium which has a single synthetic spectrum associated with it. Figure 5.13 shows a solar metallicity spectrum for a series of ages during the W-R phase of the burst. At the start of the W-R phase, at 2.5 Myr, the blue W-R bump has a sharp peaked morphology, with a strong, relatively narrow λ 4686 HeII line, and a NIII blend of λλ 4640 and 4634 on its blue side. The nitrogen bump is almost as intense as the helium bump at this stage, with two clear narrow emission lines distinguishable at the top of the feature. At this epoch, O supergiants contribute significantly to the two features, and it is their narrow lines that cause the two separate NIII features. Contributing to the feature at this time is also the first W-R star population, consisting entirely of late WN stars, which have also quite narrow emission line features. We also see He I λ 5876 with a P Cygni profile at this young age.

As we advance to 3.0 Myr, the He II emission line becomes much stronger than the nitrogen line, due to the combination of extra WNL stars, and fewer O supergiants. As the most massive O supergiants have now evolved into W-R stars, the narrow nitrogen lines disappear. We can also see a broadening of the helium line due to WNE stars appearing in the population. At this age, clearly WNL stars are dominant, with many helium lines showing some emission. By 3.5 Myr, all supergiants are gone, and the helium line is extremely broad, with the N III and He II features producing a blend. This type of feature can only be the product of the hotter, fast winds from a WNE population. The WNL population is now only represented by a small feature at the top of the broad He II bump.
Figure 5.13: Time series of an instantaneous burst at $Z_\odot$ between 2.5 and 5.5 Myr showing evolution of both the red and blue W-R bumps.
At this time, WC stars are also appearing in the spectrum, as shown by the \text{Civ} \lambda 5808 bump. Looking at the entire time series, we see that this is the peak of both the WN and WC star features. Over the next 2 Myr, these features gradually fade as the most massive stars explode as supernovae. At 5.5 Myr we see that the last W-R stars to die are again WNL stars, since these are created by the lowest mass progenitors.

If we turn to lower metallicities, looking at Figures 5.14 and 5.15, we see a shortened W-R phase, as dictated by the Meynet et al. (1994) evolutionary tracks. However, both these metallicities show the blue W-R feature with very different morphology, enabling us to differentiate between the metallicities by the shape of the features alone. At 0.4 \( Z_\odot \) we instantly see that the nitrogen line is very different. Although we still see the very narrow double line feature of the solar metallicity case, there is little or no underlying broader contribution from WNL stars. The ratio of intensities between these two lines is very different to the higher metallicity case. Another difference is that the \text{Civ} red W-R bump is predicted to be as strong as the blue feature at 3.5 Myr, since the lowering of metallicity has not weakened it as much as the \text{He}{\,\!\!\,II} \lambda 4686 bump. Overall, we expect weaker features at low metallicity however, due to the raising of the lowest mass W-R progenitor and hence fewer W-R stars. At 0.2 \( Z_\odot \), this is now 60 rather than the 40 \( M_\odot \) at solar metallicity. We see also at lower metallicity that the contribution of late type W-R stars is not swamped, due to the existence of a higher WNL/WNE ratio. At 0.2 \( Z_\odot \) we see a narrower WN feature, which is a consequence of slower (\( V_\infty \) is scaled with metallicity) and more transparent winds. The wind transparency also contributes to the weakening of the emission lines. The broader WC bump is caused by the evolutionary tracks favouring hot WC stars at low metallicity.

Given that metallicity and age change the intensity and shape of the W-R bumps, is it possible to determine the age and W-R composition of the starburst from just the consideration of these features? In order to determine this, we must also look at the effect the IMF has on the features. Obviously they will change with the IMF, since it constrains the number of the most massive stars. The upper mass cutoff limits the mass available to create W-R stars, since a star must be more massive than the W-R cutoff mass to evolve into a W-R star. The power exponent also has a large effect as it limits the mass which is available to form massive stars. Figure 5.16 shows the effect of varying the exponent, \( \alpha \) and the upper mass cutoff on the blue \( \lambda 4686 \) W-R feature of a solar metallicity burst of 3.0 Myr. Lowering the exponent obviously allows more mass to be used in high mass stars,
Figure 5.14: Time series of an instantaneous burst at 0.4 $Z_\odot$ between 2.5 and 5.5 Myr showing evolution of both the red and blue W-R bumps.
Figure 5.15: Time series of an instantaneous burst at 0.2 \( Z_\odot \) between 2.5 and 5.5 Myr showing evolution of both the red and blue W-R bumps
hence the bumps are much stronger, with the same profile. However, lowering the upper mass limits the different W-R species, which are created from stars of different masses. Here we see that a burst with an upper mass limit of 80$M_\odot$ gives the highest N \textsc{iii} to He \textsc{ii} ratio, since stars more massive than this contribute more to the He \textsc{ii} feature than the N \textsc{iii} feature. At 60$M_\odot$, N \textsc{iii} is absent due to a total lack of O supergiants. We see that the He \textsc{ii} line is broader with a high upper mass limit ($M_{\text{up}} > 100M_\odot$), since the most massive stars create the hot WNE stars with faster flowing winds. At an upper mass limit of 40$M_\odot$ of course, there are no stars massive enough to create W-R stars and no W-R bump feature is present. From 80 to 120$M_\odot$, however, the feature looks very similar, leading us to the conclusion, that over this range, at least, a W-R bump has a unique shape for a certain age and metallicity. The change in intensity of the bump due to the power-law exponent $\alpha$ is useful, however, as it allows us to constrain the IMF of a burst more precisely. This could be a direct way in which to measure the massive star population within a starburst, if the evolutionary tracks used in the model are reliable. However, due to resolution limits and the introduction of noise, it is possible that bumps at different metallicities and ages could appear extremely similar. In this case, we can use nebular emission lines to constrain metallicity and put limits on the age of a starburst. This is the method used in the next section.

5.3 Comparisons of W-R Features

Here we present a preliminary analysis of five W-R galaxies using datasets kindly donated by colleagues. This section is intended to outline the usefulness of the new atmosphere grids and show applications for the new W-R bump spectral synthesis routine.

5.3.1 Observations

Two sets of observations were used. The first set of observations of three W-R galaxies was conducted by Crowther & Abbott, taken on the 1st to 3rd of August 2002 using the 4.2 metre William Hershel Telescope (WHT) and ISIS double beam spectrograph. All galaxies were observed with the EEV12 CCD on the blue arm, with an array of a 4096 $\times$ 2048 (13.5$\mu$m pixels) and the red arm with the Marconi2 CCD with a 4700 $\times$ 2148 array (also 13.5$\mu$m pixels). In each case we used a 1$''$ slit centred on the nucleus of the galaxy in question, and exposures of 1200 seconds were taken. The observational setup used a 6100 Å dichroic filter to generate a spectrum of blue light from 3500 to 6100 Å and a red
Figure 5.16: The effect of the IMF on the blue $\lambda$ 4686 feature, for 5 different upper mass cutoffs and 3 different $\alpha$ exponents. The upper mass is plotted vertically from 120 to 40 $M_\odot$ and the IMF slope is plotted horizontally.
Table 5.6: W-R galaxies used in photoionization and spectral fitting

<table>
<thead>
<tr>
<th>Name</th>
<th>Position (1950) α</th>
<th>δ</th>
<th>Morphology*</th>
<th>m_b (mag)</th>
<th>V(rad)</th>
<th>Mpc</th>
<th>Reference</th>
</tr>
</thead>
<tbody>
<tr>
<td>NGC6764</td>
<td>19 07 01.2</td>
<td>+50 51 08</td>
<td>SB(s)bc</td>
<td>12.6</td>
<td>2308</td>
<td>30</td>
<td>1</td>
</tr>
<tr>
<td>VII Zw 611</td>
<td>15 33 04.1</td>
<td>+57 27 00</td>
<td>—</td>
<td>—</td>
<td>3267</td>
<td>44</td>
<td>1</td>
</tr>
<tr>
<td>NGC6500</td>
<td>17 53 48.1</td>
<td>+18 20 41</td>
<td>SAab</td>
<td>13.1</td>
<td>2788</td>
<td>37</td>
<td>1</td>
</tr>
<tr>
<td>Mrk 1094</td>
<td>05 08 17.4</td>
<td>−02 44 33</td>
<td>I0 pec?</td>
<td>14.1</td>
<td>2700</td>
<td>36</td>
<td>2</td>
</tr>
<tr>
<td>NGC3049</td>
<td>09 52 10.2</td>
<td>+09 30 32</td>
<td>SB(rs)ab</td>
<td>13.0</td>
<td>1440</td>
<td>19</td>
<td>2</td>
</tr>
<tr>
<td>NGC3125</td>
<td>10 04 19.1</td>
<td>−29 41 30</td>
<td>S</td>
<td>13.5</td>
<td>900</td>
<td>12</td>
<td>2</td>
</tr>
<tr>
<td>II Zw 40</td>
<td>05 53 05.0</td>
<td>+03 23 03</td>
<td>BCD;Irr</td>
<td>15.5</td>
<td>690</td>
<td>9</td>
<td>2</td>
</tr>
</tbody>
</table>

References: 1, Crowther & Abbott 2002; 2, Vacca & Conti (1992); Notes: * Morphology from Schaerer et al. (1999)

The spectrum from 6100 to 8700 Å in the blue and red beams respectively. The data were reduced and calibrated by Abbott (private communication, 2002) using standard IRAF procedures. This dataset has a resolution of 6 Å and a signal to noise of 37. Additional data for four W-R galaxies were obtained from Dr. W. Vacca (Vacca & Conti 1992). The observational dataset is summarized in Table 5.6.

### 5.3.2 Experimental Procedure

Two of the three spectra obtained by Crowther and Abbott showed no real evidence for W-R features. Both of these objects have been designated as W-R galaxies in the past. NGC 6500 was recognised as a W-R galaxy by Barth et al. (1997) and VII Zw 611 was signalled to have features by Guseva et al. (2000) although an earlier work by Izotov et al. (1997) mentions no broad features. These two spectra were excluded from the work due to the lack of evidence of W-R features in our data. The negative detection in VII Zw 611 could well have been due to a misalignment of the slit during the observation, since the spectra is of extremely high quality, although the dataset of Guseva et al. (2000) is extremely. The NGC 6500 spectra, however, was extremely noisy, due to observing conditions, therefore it was impossible to give a positive detection to the blue W-R feature. The spectrum of NGC 6764 was de-reddened in the manner suggested by Osterbrock (1974) via the readjustment of the relative strengths of Hβ to Hα, Hβ to Hγ and Hβ to Hδ. The metallicity was
The ionic abundances were determined for O$^+$ and O$^{2+}$ by using the code EQUIB written by S. Adams and I. D. Howarth to convert relative [O III] $\lambda$5007 and [O II] $\lambda$3727 into abundance. An electron temperature of 10,000K was assumed since $\lambda$4363 was not detected and a density of 300cm$^{-3}$ was used in accordance with Eckart et al. (1996). For the methodology of the determination of parameters in the other datasets, the reader should consider the references given in Table 5.6. The parameters for all the observations are summarised in Table 5.7.

In order to test the validity of the W-R bump spectral synthesis, we model the starburst using the model continuum from the updated Starburst99 code as an input to the photoionization code CLOUDY. For CLOUDY, we used the metallicity most appropriate to model the nebula, generating six input files, each representing the continuum of a different starburst age, starting with 1 Myr, and finishing at 6 Myr, (after the W-R phase had finished) in time steps of 1 Myr. As in the fashion of CLOUDY, the continuum was input in units of Rydbergs vs Energy (eV) using all 1221 Starburst99 data points (10.08 to 0.00057 Rydbergs) for the maximum possible resolution. We used a spherical geometry with a radius of 1 kpc to ensure that all the ionizing flux was absorbed and re-emitted, a covering factor of 1.0, and the default CLOUDY abundances for all metals and dust grains and scaled these values to the metallicity observed in the nebula (presented in Table 5.8). A constant density was assumed throughout the nebula.

To ensure a realistic fit to the observations, we used the observed line intensities from eight lines, if available and the intrinsic luminosity of H$\beta$. These were matched by CLOUDY using its built-in optimization routine using the SUPLEX method, which varies the parameters which have been allowed to remain free until a satisfactory fit is obtained, or 20 iterations were reached. In this case the best fit is calculated. In our models we decided that the input luminosity, the filling factor and the hydrogen densities should be varied, as these quantities all relate to the ionization parameter U, which alters the ionization balance of the nebula. These parameters were chosen since the filling factor was unknown, and the luminosities and hydrogen densities are uncertain. All other quantities were kept constant.

When a best fit was obtained, we used the synthetic spectra to model the blue $\lambda$4686 W-R bump, and where possible the red $\lambda$5808 W-R feature. Using the age constraints determined from the nebular modelling, and the atmosphere grid metallicity that was
### Table 5.7: Parameters of W-R galaxies

<table>
<thead>
<tr>
<th>Object Name</th>
<th>Extinction factor $c(H\beta)$</th>
<th>$n_e$ cm$^{-3}$</th>
<th>$T_e$ (K)</th>
<th>$L(H\beta)$ erg s$^{-1}$</th>
<th>$Z/Z_\odot$</th>
</tr>
</thead>
<tbody>
<tr>
<td>Mrk 1094</td>
<td>0.03</td>
<td>90</td>
<td>9440</td>
<td>$5.7 \times 10^{39}$</td>
<td>0.32</td>
</tr>
<tr>
<td>NGC3049</td>
<td>0.34</td>
<td>259</td>
<td>10000*</td>
<td>$8.4 \times 10^{39}$</td>
<td>1.48</td>
</tr>
<tr>
<td>NGC3125</td>
<td>0.4</td>
<td>140</td>
<td>10200</td>
<td>$1.3 \times 10^{40}$</td>
<td>0.30</td>
</tr>
<tr>
<td>II Zw 40</td>
<td>0.35</td>
<td>50</td>
<td>12600</td>
<td>$1.9 \times 10^{40}$</td>
<td>0.16</td>
</tr>
<tr>
<td>NGC 6764</td>
<td>1.24</td>
<td>300**</td>
<td>10000**</td>
<td>$1.6 \times 10^{40***}$</td>
<td>0.44</td>
</tr>
</tbody>
</table>

Notes: * Vacca & Conti (1992) assumed value; ** Eckart et al. (1996) values; *** value assuming $H_0 = 60 \text{km s}^{-1} \text{Mpc}^{-1}$

closest to that of the nebula, we matched the synthetic spectra to the observations with the following procedure. Care was taken to deredden the spectra using the values quoted in existing works, since large Balmer jumps in the observations indicate that all have contamination by an older population of less massive stars ($\geq 10$ Myr). As a synthetic spectrum without the addition of an older population yielded reddenings significantly in excess of values quoted by earlier reference work, we added the older continuum to the W-R bump synthetic spectrum in different ratios in order to reduce the reddening to an acceptable value, with consideration of the size of the Balmer jump in the observation. This method, however, has the undesirable effect that a W-R feature predicted by the combination of the two spectra can be considerably diluted. Although strengthening this feature is possible through altering the IMF slope, we have to keep the IMF constant, since we have no direct evidence that an altered IMF is applicable.

#### 5.3.3 Results

**NGC 6764**

W-R bumps were first found in the spectrum of this galaxy by Osterbrock & Cohen (1982). A spectrum obtained by Eckart et al. (1996) confirmed the W-R nature of this galaxy. This analysis found that the galaxy is still undergoing star formation on a time-scale of $10^7$ yrs. We determine an extinction coefficient of 1.24. Our photoionization modelling gives a best fit for a 3.0 Myr burst at $0.4 \times Z_\odot$ metallicity. In the model presented in
Table 5.8: Observation vs model results

| Object   | Origin | [O II] | [O III] | [O III] | He I | [N II] | [S II] | Age (Myr) | Z | Log L(H/β) (ergs s⁻¹) | log n_H (cm⁻³) | χ²|^a |
|----------|--------|--------|---------|---------|------|--------|--------|------------|---|---------------------|----------------|-----|
| NGC 6764 | obs    | 370    | 17.4    | 55.4    | 8.4  | 146    | 44.1   | 37.1       |   | 0.44                | 40.2            | ... |
|          | Model  | 295    | 20.9    | 60.4    | 7.8  | 99.9   | 44.8   | 50.3       | 3.0| 0.40                | 40.5            | 2.94 |
| Mrk 1094 | obs    | 241    | 98.1    |         | 12.9 | 26.9   | 42.6   | 31.6       |   | 0.32                | 39.8            | ... |
|          | Model  | 264    | 73.2    |         | 8.3  | 53.6   | 26.7   | 19.4       | 4.0| 0.40                | 39.8            | 1.98 |
| NGC 3049 | obs    | 126    | ...     | 32.0    | 10.4 | 108    | 32.0   | 24.9       |   | ≥1.0                | 40.3            | ... |
|          | Model  | 107    | ...     | 84.5    | 8.3  | 86.9   | 20.7   | 24.9       | 4.0| 2.0                 | 40.44           | 2.95 |
| NGC 3125 | obs    | 193    | 3.4     | 170     | 7.8  | 8.9    | 10.0   | 7.8        |   | 0.30                | 40.1            | ... |
|          | Model  | 126.7  | 5.3     | 233     | 9.4  | 17.4   | 7.0    | 6.4        | 2.0| 0.40                | 40.2            | 2.63 |
| III Zw 40| obs    | 68.2   | 8.8     | 238     | 9.1  | 6.9    | 6.7    | 4.9        |   | 0.16                | 39.9            | ... |
|          | Model  | 60.7   | 9.6     | 241     | 10.3 | 9.8    | 4.1    | 4.1        | 2.0| 0.20                | 39.9            | 2.82 |

^aχ² value is the sum of the average χ² relative intensity and Log L(H/β) values. (χ² = \[\sum_{i=1}^{n} (x_i - \bar{x})^2/(n-1)\])

^bMetallicity of model is that of the input Starburst99 continuum. A metallicity equal to that of the observations was used in the photoionization fitting.
Figure 5.17: The spectrum of NGC 6764 fitted with a 2.5 Myr, $Z_\odot$ metallicity instantaneous burst model.
Table 5.8 we see a good $\chi^2$ fit, with excellent agreement in the oxygen lines, especially the [O III] lines. [OII] is slightly under predicted with L(H\beta) being over predicted in order to increase the ionizing flux to the required level, suggesting that the continuum is slightly too soft to exactly reproduce the nebular emission. The derived age of this starburst fits well with the W-R phase for the metallicity of this environment.

If we now consider the synthetic spectrum fit presented in Figure 5.17, we see that consistency between the synthesis and photoionization models can be reached. The best fit is a 2.5 Myr model $Z_\odot$ model. An $E(B-V)$ value of 0.55 was obtained for the stellar continuum reddening, which although, high is in accordance with the $0.4 \pm 0.2$ value of Eckart et al. (1996). This synthetic spectrum includes no contribution from an older population. The higher metallicity was used to reproduce the N III $\lambda\lambda$ 4634, 4640 features, since at $0.4 \times Z_\odot$ the feature is extremely weak (see the time series of spectra in Fig.5.14). At an age of 2.0 Myr, no W-R stars exist, so the only stellar emission is from the weaker-lined O supergiant population, which cannot produce intense enough emission lines. A closer look at the He II $\lambda$4686 feature reveals a complex structure. The W-R bumps feature is composed of two separate resolved features, the He li $\lambda$4686 feature and the N iii $\lambda\lambda$ 4634, 4640 blend. In between these features is the nebular [Fe III] line. This line cannot possibly be produced directly from the stellar population, and has a FWHM comparable to the H\beta nebular emission line. The nitrogen bump can clearly be identified as the combination of $\lambda\lambda$ 4634 and 4640 in both the observation and model, with a hint of N v $\lambda$4609 being present in both. The model underestimates the nitrogen line contributions, and overestimates the He II line. In fact the N III line appears stronger than the helium line, which is not observed in any single stars. This phenomenon is well known. In fact Schmutz & Vacca (1999) showed that this feature cannot possibly be produced solely by WN stars, but also must include an O supergiant population. Since we include the O supergiant contribution, which tends to add very narrow lines, we manage to reproduce the lines with some success, although they are still underestimated. This under prediction may well be solved with the addition of rotation in the evolutionary tracks, which will increase the number of blue supergiants and therefore the N III line strength.

This tends to point towards the younger age that our synthetic spectra suggest, rather than the slightly older photoionization determination.

Although the observation was taken between 3000 to 9000 Å we are not able to fit the red 5808 Å bump, due to the existence of a strong He i $\lambda$5876 nebular emission line and
an interstellar absorption feature (Na I 5890) which drowns the weak WC contribution. Using the stellar population information from Starburst99, and information on absolute distance from the WR galaxy survey of Conti (1991), we measure an absolute mass of the system and hence determine the W-R population. The W-R population for all of the galaxies is presented in Table 5.9. Our value of 3766 WR stars is extremely close to the value quoted in Eckart et al. (1996) of 3600.

Mrk 1094 \equiv II Zw 33

These data are from work published by Vacca & Conti (1992). They determine the metallicity to be 0.32 \text{Z}_\odot assuming the Edmunds & Pagel (1984) relationship between oxygen line intensities and oxygen abundance. A very low extinction correction was found, at \text{c(H/}\beta) = 0.03. The emission line intensities were extremely high in the data set, hence [O\text{III}] \lambda5007 was saturated and was not included in the line list. Some success was achieved in the photoionization modelling of this galaxy, especially with the oxygen and helium lines, suggesting a good agreement with the predicted level of the ionizing flux from the input continuum. The [N\text{II}] \lambda6584 line was over-predicted by a factor of two and the [S\text{II}] lines were under-predicted by nearly 40 per cent. These lines were however modelled to be in the correct ratio. To add to this, a perfect match for the H\beta line luminosity was found, giving overall a satisfactory fit. An age of 4.0 Myr and hydrogen density of 95 cm$^{-3}$ was determined from the model fit.

The spectrum of this object, presented in Figure 5.18 unfortunately only covers the spectral range up to 5400 Å, so no red bump can be modelled. This spectrum is also very noisy at the continuum level, but there is definitely a He II \lambda 4686 feature present. Matching a synthetic spectrum with the observation, we can see that the W-R feature could be represented by the model, but due to the high level of noise, ages from 3.0 to 4.5 Myr could fit the data. This age spread can be narrowed down considerably if data existed for the red W-R bump, since WC stars dominate for a much smaller time period (< 1 Myr).

We determined a WN population of \sim 240 stars as opposed to the 130 Vacca & Conti (1992) estimated. To add to this, the later age of this burst means that the WC population should be \sim 2500, although no data is available for this region. Our reddening value of 0.51 may account for the excess in WN stars predicted as the Vacca & Conti (1992) value (0.03) is extremely low. Recent observations by Méndez et al. (1999) puts the reddening
Figure 5.18: The spectrum of Mrk 1094 fitted with a 4.0 Myr, 0.4 Z\(_{\odot}\) metallicity instantaneous burst model.
at about 0.52, consistent with our value.

\textit{NGC 3049} $\equiv$ Mrk 710

We initially used the Vacca & Conti (1992) values for the basis of our photoionization model. Using these values, we were only able to obtain poor model fits, with the best model having a $\chi^2$ value of 500. We then considered the new values from a recent study of this galaxy by González Delgado et al. (2003). These new values were then used as the input to model the nebula and are presented in Table 5.8 instead of the Vacca & Conti (1992) data. Although González Delgado et al. (2003) mention that this object has a super-solar metallicity, no formal value is given, so we use the 1.48 $Z_\odot$ value from Vacca & Conti (1992). The new model allows us to obtain an age of 4.0 Myr, in agreement with the 3.5 Myr quoted in the new work, calculated from spectral synthesis of the UV wind lines arising from O stars. Although our fit is much improved, we still have some problem with the overprediction of the [O III] $\lambda$5007 line, indicating that the ionizing continua is still too hard. However, a 5.0 Myr model gives a similar $\chi^2$ value, but correctly estimates the [O III] $\lambda$5007 line and underestimates the [O II] $\lambda$3727 line indicating a continua that is too soft. A probable best fit age will therefore lie between 4 and 5 Myr.

The spectrum (Fig. 5.19) shows good evidence for the blue W-R bump features, showing some agreement with our 2.5 Myr $Z_\odot$ model. As with the NGC 6764 dataset the HeII bump seems to be overestimated, while the NIII bump seems to fit extremely well. This early age seems to indicate that a further softening of the ionizing continuum is required for this object.

From the spectral fit we find that the WN population is 280, compared to the Vacca & Conti (1992) value of 380.

\textit{NGC 3125} $\equiv$ Tololo 3.

The photoionization model of NGC 3125 gives reasonable agreement with the Vacca & Conti (1992) observations. A 0.4 $Z_\odot$ model was used to provide the ionizing continuum, although it is unclear whether a lower metallicity of 0.2 $Z_\odot$ should be used due to its borderline metallicity of 0.3 $Z_\odot$. We get a reasonable agreement with the observations, although the oxygen ionization balance is too high, with too little [OII] and too much
NGC 3049 (Vacca & Conti)
1.0$Z_\odot$ model 2.5 Myr

Figure 5.19: The spectrum of NGC 3049 fitted with a 2.5 Myr, 1.0$Z_\odot$ metallicity instantaneous burst model.
Figure 5.20: The spectrum of NGC 3125 fitted with a 3.0 Myr, 0.4 $Z_\odot$ metallicity instantaneous burst model.
[O\textsc{iii}]. The nitrogen line is again over-predicted by 200 per cent, and the sulphur is under-predicted, but in the correct ratio.

When we turn to the comparison of the spectrum however (Fig. 5.20), we see excellent agreement with the observation for the W-R features. The model predicts almost the exact intensity of He \textsc{ii} observed with the same equivalent width. In fact the fit is so good that we can pick out the nebular [Fe \textsc{iii}] and [Ar \textsc{iv}] lines. Unfortunately, this fit requires a slightly older age burst than that predicted by the photoionization modelling, which has a best fit at 2 Myr. This age is actually too early for any W-R features at all since it gives O stars too little time to evolve into W-R stars. A 2 Myr burst is dominated by O supergiants and resembles the 2.5 Myr plot in Figure 5.14, where the nitrogen bump is more intense than the helium bump, although far narrower.

Vacca & Conti (1992) determine the number of WN stars to be approximately 390, whereas our method gives a lower value of 130.

\textit{II Zw 40.}

II Zw 40 is a good example of a successful photoionization model fit. The model of this galaxy gives excellent agreement to the observations with all lines falling in the tolerances set by the observations with the possible exception of the nitrogen line which is overpredicted by just over 40 per cent and the [S \textsc{ii}] \(\lambda 6716\) line which is underestimated. Since this fit was achieved via a \(\chi^2\) fitting method testing a wide parameter space, it may be possible to get an exact match for this galaxy if special care is taken.

The spectrum (Fig. 5.21) is in agreement with the 2.0 Myr fit from the model, as no He \textsc{ii} is predicted at this age. The spectrum, however, is extremely noisy, due to contamination from the strong [Ar \textsc{iv}] \(\lambda\lambda 4714\) and 4740 lines and the [Fe \textsc{iii}] \(\lambda 4658\) line, and it is impossible to exclude any age within the W-R phase. Plotted on the figure is the 3.5 Myr fit showing a possible W-R bump feature underlying the four marked nebular lines. The spectrum also appears to show He \textsc{ii} \(\lambda 4686\), although since this line is narrow, it is more likely to be of a nebular origin than from W-R stars. Clearly, higher signal to noise data are required for this object.

\textbf{5.3.4 Summary}

Comparisons with the optical extragalactic dataset of Bresolin \textit{et al.} (1999) and the mid-IR galactic dataset of Giveon \textit{et al.} (2002) clearly favour the ionizing continua of the new
Figure 5.21: The spectrum of II Zw 40 fitted with a 3.5 Myr, 0.2 $Z_{\odot}$ metallicity instantaneous burst model.
Table 5.9: The W-R population in the programme galaxies

<table>
<thead>
<tr>
<th>Name</th>
<th>$E_{(B-V)}$</th>
<th>Age (Myr)</th>
<th>$M_{\odot}$ ($10^6$)</th>
<th>$N_{O}$</th>
<th>$N_{WN}$</th>
<th>$N_{WC}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>NGC6764</td>
<td>0.55</td>
<td>2.5</td>
<td>33.3</td>
<td>153200</td>
<td>3766</td>
<td>...</td>
</tr>
<tr>
<td>Mrk 1094</td>
<td>0.51</td>
<td>4.0</td>
<td>6.7</td>
<td>26000</td>
<td>240</td>
<td>2500</td>
</tr>
<tr>
<td>NGC3049</td>
<td>0.50</td>
<td>2.5</td>
<td>5.0</td>
<td>7000</td>
<td>280</td>
<td>...</td>
</tr>
<tr>
<td>NGC3125</td>
<td>0.50</td>
<td>3.0</td>
<td>2.0</td>
<td>5800</td>
<td>130</td>
<td>...</td>
</tr>
</tbody>
</table>

SB2002 models. In all three of the softness parameter plots the new models overlap some of the observations, as opposed to the SV98 models which are too hard in two of the three plots. This clearly points towards the greater suitability of the new ionizing continua to model starburst populations in the important UV spectral region. To add to this evidence, we can see from Chapter 3 that individual O and W-R star models from our new grid are superior to older work.

A long standing problem with matching the predictions of evolutionary synthesis models to observations is that the age derived from the UV stellar wind lines is always much less than that derived from the nebular emission line analysis (e.g. González Delgado et al. 2003). The most likely explanation is that the ionizing continua are too hard, particularly during the W-R phase, and thus ages > 5 Myr are needed to fit nebular emission line ratios. We have shown in this chapter that for the first time, consistent ages for the nebular and stellar components of W-R galaxies can be obtained with the new grid of line-blanketed, non-LTE models incorporated in SB99.

Using the observations of five W-R galaxies, we are able to demonstrate that the new evolutionary synthesis models can not only model the ionizing continua of complex systems, but can predict W-R spectral features that can match the data. This new tool will allow us to probe the W-R population within a starburst and provide a clear indication of the metallicity and age of the population.

Although we are unable to make predictions of the IMF parameters from these W-R
spectral features, it may be possible to do so by combining this work with that of González Delgado et al. (1999) who have synthesised the Balmer and He I lines in the same spectral region. They maintain that these lines can give an indication of the slope of the IMF. A combination of this and the W-R features may allow us to constrain the IMF and age of the burst.
Chapter 6

Summary and Conclusions

We review the main points of the work in this Chapter, summarising the main conclusions that have been reached. The final section is devoted to the description of work that must be done as a follow on to develop the conclusions of this thesis.

6.1 Summary

6.1.1 Stellar Atmosphere Grids

We present a grid of 165 non-LTE line blanketed O and early B star atmosphere models calculated using the WM-basic code of Pauldrach et al. (2001) to replace the older LTE atmosphere grid of Lejeune et al. (1997) in Starburst99. These models span five metallicities (0.05, 0.2, 0.4, 1.0 and 2.0$Z_{\odot}$) and cover temperatures from 25 to 50kK with three surface gravities corresponding to the dwarf, giant and supergiant luminosity classes. The atmospheres represent spectral types O3 to B1 and are presented in Tables 2.1, 2.2 and 2.3, §2.3, Chapter 2. Fundamental parameters are defined using the methods described in Chapter 2 §2.3. The new models are compared to the LTE grid of Lejeune et al. (1997) and CoStar non-LTE models of Schaerer et al. (1996). We find that the new atmospheres predict more ionizing flux than the Lejeune et al. (1997) work, due to the microscopic assumptions of the model, but predict less ionizing flux on average than the CoStar grid due to a more complete incorporation of line blanketing.

To complement the new O and early B star models we present a new grid of Wolf-Rayet models calculated with the non-LTE, co-moving frame code CMFGEN of Hillier & Miller (1998). We calculate 120 W-R models, spanning the same range of metallicity as
the OB grid. The grid is split into two separate components, due to the surface chemistry of the star, and hence a WN and a WC grid is calculated. Definition of the fundamental parameters of both grids are found in §2.5, Chapter 2. The WN grid spans a temperature range of 30-120kK, with the WC grid being slightly hotter with a temperature range of 40-140kK. Comparisons with the existing Schmutz et al. (1992) pure helium non-LTE grid show that the inclusion of line blanketing reduces the ionizing flux by many factors in some stars, especially the cooler models. We assume a mass-loss scaling with metallicity of all new atmosphere grids including W-R stars. Mass-loss is scaled as \( \dot{M} \propto Z^{0.8} \) and \( V_\infty \propto Z^{0.13} \).

Chapter 3 compares the new OB and Lejeune et al. (1997) grids with the H ii region dataset of Bresolin et al. (1999), using the photoionization code CLOUDY to predict nebular line strengths. We use two diagnostic plots, a metallicity versus He i plot in which we include the limited CoStar atmosphere grid and a \( R_{2,3} \) versus \( \eta' \) comparison. In the first plot, the WM-basic grid shows a downturn in He i at high metallicity which mimics the trend in the data, unlike the Costar and LTE models. The \( \eta' \) plot shows WM-basic to be superior to Lejeune et al. (1997) in reproducing the observational data. §3.3 presents the photoionization modeling of two W-R nebulae (one cool WN9 and a hot WN2 star) using CLOUDY.

### 6.1.2 Implementing the New Atmospheres in Starburst99

In Chapter 4 we present a brief overview of the Starburst99 evolutionary synthesis model. We explain the implementation of the WM-basic OB star grid, and the CMFGEN W-R star grid. We deal with the coupling of the W-R star grid with the evolutionary tracks, explaining our choice of interpolating between the \( T_{hpd} \) and \( T_{2/3} \). We set the temperature at:

\[
T_* = 0.6T_{hpd} + 0.4T_{2/3}
\] (6.1)

In §4.4 we compare the predictions of the SED and ionizing fluxes from our new models with those of Leitherer et al. (1999) and a hybrid Starburst99 model containing the CoStar O star models. We find that the new models predict \( 10^4 \) times more He\(^+\) ionizing flux than the SB99 model in the early O star phase of the starburst, but 10 times less than the CoStar model for a solar metallicity instantaneous burst. In the W-R phase of this burst model we predict \( \sim 10^9 \) times less He\(^+\) ionizing flux due to the heavy blanketing
of the CMFGEN W-R models. In another comparison (Fig. 4.15), we see that the new models start to converge with both the SV98 and SB99 predictions at low metallicity, due to the assumed W-R wind scaling. We look at the case of continuous star formation. Our models show little change in the He$^+$ ionizing flux in high to intermediate metallicity models, compared to the sharp change in flux observed in the other models as the first W-R stars are generated. In all comparisons the ionizing flux is lower than the older starburst models.

As an extension to this, we look at the predicted nebular He II flux from the CoStar and new SB2002 models. We find that contrary to existing predictions of the Schaerer & Vacca (1998) model, no detectable (above 1 percent of H$\beta$) He II λ 4686 flux is predicted at any metallicity during any phase of the starburst. We postulate that the ionizing source of the line observed in starburst nebulae must arise from another source, possibly from X-ray shocks.

We end Chapter 4 by considering the implications of the new mass-loss rates assumed in the production of the model grids. We find that although the W-R mass-loss rates are lower, especially at low metallicity due to scaling, the O star mass-loss rates are higher, and balance out the effect at high metallicity giving the same overall result. At low metallicity, the early B star mass-loss rates are high enough to balance the W-R star deficit, due to inaccuracies in the theoretical Castor et al. (1975) B star mass-loss rates.

6.1.3 Comparisons with Data and Further Applications of the Atmosphere Grids

In Chapter 5, we apply the new SB2002 model to two datasets, optical data from Bresolin et al. (1999) and IR data from Giveon et al. (2002). We use the CoStar Starburst99 model as a comparison.

For the optical dataset we use the $R_{2,3}$ versus $\eta'$ plot as a diagnostic and find that the new SB2002 models fit the data more satisfactorily than the older CoStar models, due to the softer ionizing continuum of SB2002. In the IR we define two $\eta'$-like parameters from combinations of the ratios of [S IV]/[S III], [Ne III]/[Ne II] and [Ar III]/[Ar II]. We compare the CoStar and SB2002 models with these $\eta'$-like parameters, defined in Chapter 5, §5.1, named $\zeta_1'$ and $\zeta_2'$. We find that the SB2002 models are coincident with the observations which have harder softness parameters in a $\zeta_1'$ comparison and cover the main body of the data in the $\zeta_2'$ plot. The new SB2002 models fit the data far better than the SV98 models.
which have ionizing continua that are too hard. We conclude that the SB2002 models
give a much more realistic ionizing distribution than SV98 and are able to reproduce the
observational values with success.

In §5.2 we discuss a new modelling tool developed with the newly calculated cmfgen
atmospheres. We introduce the method in which we synthesize the red and blue optical
broad emission line features found in some starburst galaxies due to the presence of the
W-R population. In order to do this we calculate 40 O supergiant models with cmfgen
to reproduce the emission line features found in Of star spectra. Examples of the time
and IMF dependence of these features are given.

In §5.3 we model the optical nebular line intensities of four W-R galaxies using the
photoionisation code cloudy. We are able to reproduce most of the nebular line features
in all of these galaxies with some low \( \chi^2 \) fits. The new high metallicity SEDs allow us
to reproduce the relative emission line intensities for the super solar metallicity galaxy
NGC 3049 for the first time, as previous attempts have been forced to use starburst ages
far older than UV stellar wind lines suggest. Our determination is only 1 Myr older
than the 3.5 Myr wind line value of González Delgado et al. (2003). In tandem with
this technique, we use the newly developed W-R feature synthesis routine to fit the W-R
features found in the optical spectrum. Due to the nature of the data, we could only fit the
blue He II \( \lambda 4686 \) bump. In several of the observations we are able to obtain good fits, with
three of the fits matching the age determinations of the photoionization modelling. One
spectrum in particular, that of NGC 3125 has an excellent W-R bump fit. The W-R fits
allow us to quantify the W-R population of the parent starburst via use of the evolutionary
tracks, in which we obtain consistent results with previous studies.

6.1.4 Conclusions

We briefly list the main conclusions of this thesis.

- The new OB atmospheres represent a significant improvement over the old LTE
  atmospheres of Lejeune et al. (1997), and a reasonable improvement, due to the
  more complete inclusion of line blanketing over the CoStar models.

- The new W-R stars are a vast improvement over the Schmutz et al. (1992) pure
  helium models.

- The W-R phase of the new model starburst is not the dominant source of ionizing
flux at high metallicity, instead O stars are more important. At low metallicity W-R stars become more dominant. However, there are too few W-R stars present to contribute significantly to the ionizing flux.

- No He II line emission is now predicted to originate from the ionizing flux of the young stellar population within a starburst.

- The mass and energy returned by the starburst into the ISM is now predicted to be dominated at low metallicity by late O and early B stars.

- The ionizing fluxes predicted by SB2002 are a vast improvement over existing models allowing for age determinations via photoionization modelling to coincide with determinations from other methods such as observations of W-R bumps.

- The new W-R bump spectral synthesis routine allows for the determination of the W-R population of the parent starburst, and may help to constrain the IMF.

6.2 Future Work

The addition of realistic ionizing fluxes has allowed the Starburst99 evolutionary synthesis model to become a reliable tool for modelling starbursts via their nebular line ratios for the first time. However, improvements in the model can still be made to increase the accuracy of the predictions. The area in need of biggest improvement is in the evolutionary tracks. These tracks cannot be wholly relied upon for predictions of W-R star evolution. The new evolutionary tracks including rotation represent a significant improvement over the old Meynet et al. (1994) tracks, although they are not as yet on general release. An inclusion of these tracks will have obstacles to overcome however, as the tracks represent a variety of different initial rotational velocities for which a realistic statistical distribution must be defined. The inclusion of binary tracks must also be considered, for which the binary fraction of massive stellar populations must first be constrained. Due to recent interest in zero and very low metallicity environments, this work could be extended to include these tracks and extra stellar atmospheres. The new W-R synthesis routine will benefit from more realistic evolutionary tracks as it will solve the problem of coupling the atmospheres to the tracks. The addition of synthetic Balmer line profiles will increase the power of this tool further, allowing us to not only constrain W-R populations within the burst, but possibly to also start to constrain the shape of the IMF. The next step will be
to obtain high resolution, high signal to noise data for a large sample of W-R galaxies. Higher resolution will allow us to differentiate between nebular and stellar components of the blue W-R feature, and define the shape of both bumps more closely. As observations of the red W-R bump do exist, it will be crucial to obtain spectra of W-R galaxies in this region. In the more immediate future, the resolution of the Starburst99 atmosphere grid needs to be increased in the high energy range. Due to the larger memory capacity of contemporary computers, the resolution could be raised by a factor of 10 without any loss of performance, as demonstrated by the higher resolution W-R bump spectral synthesis grid.

Modelling the far UV between 912 to 1200 Å could also be extremely important as this can be a strong indicator of the stellar population within a starburst, especially the O supergiant population (González Delgado et al. 1997). The Far Ultra-violet Spectroscopic Explorer (FUSE) satellite covers this spectral range with good signal to noise and excellent spectral resolution. Modelling has recently been carried out by Robert et al. (2003) using observed template spectra to cover 3 metallicities (0.2, 0.4 and 1.0Z☉). However, in this spectral region there is no unique W-R spectral feature, therefore their study does not duplicate the work in this thesis. W-R spectral features do exist in the far UV, but long ward of the FUSE spectral range. Instead we need to use the older Internal Ultra-violet Explorer satellite with its spectra range from 1150 - 2000 Å. Large 8-10m telescopes can be used to obtain rest frame spectra of z ~ 2–3 galaxies. Spectral synthesis of this region can be performed by either using observation templates as in the Robert et al. (2003) work, or as in this work, with the use of a grid of W-R models.

As we have added new mass-loss rates to Starburst99, it is also important to use the predictions to explore the chemical yield and energy release from a burst, as this has a large influence on the feedback of the system which is extremely important in cosmological models. At low metallicity we now predict much higher levels from the early B star population which may help this problem.

The new models are available to download at www.star.ucl.ac.uk/starburst.html
List of Publications

Interactions of Young Starbursts With Their Environments: Realistic Ionizing Fluxes From 0.05 to $2 \times Z_{\odot}$.

Realistic Ionizing Fluxes For Young Stellar Populations From 0.05 to $2 \times Z_{\odot}$.

Wolf-Rayet Populations in Starburst Galaxies.

New line-blanketed model atmospheres and their impact on synthesis models
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