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Abstract

This thesis starts with an introduction to the solar atmosphere and the physics that governs its behaviour. The formation processes of spectral lines are presented followed by an explanation of employed plasma diagnostic techniques and line broadening mechanisms. The current understanding on some principle concepts of flare physics are reviewed and the topics of flare homology and non-thermal line broadening are introduced.

The many solar satellites and instrumentation that were utilised during this thesis are described. Analysis techniques for some instruments are also presented. A series of solar flares that conform to the literature definition for homologous flares are examined. The apparent homology is shown to be caused by emerging flux rather than continual stressing of a single, or group of, magnetic structures. The implications for flare homology are discussed.

The analysis of a solar flare with a rise and peak in the observed non-thermal X-ray line broadening \((V_{nt})\) is then performed. The location of the hot plasma within the flare area is determined and consequently the source of \(V_{nt}\) is located to be within and above the flare loops. The flare footpoints are therefore discarded as a possible source location. Viable source locations are discussed with a view to determining the dominant mechanism for the generation of line broadening.

The timing relationships between the hard X-ray (HXR) flux and \(V_{nt}\) in many solar flares are then examined. I show that there is a causal relationship between these two parameters and that the HXR rise time is related to the time delay between the maxima of HXR flux and \(V_{nt}\). The temporal evolution of \(V_{nt}\) is shown
to be dependent upon the shape of the HXR burst. The implications of these results are discussed in terms of determining the line broadening mechanism and the limitations of the data.

A summary of the results in this thesis is then presented together with suggestions for future research.
Chapter 1

Introduction

The Sun has been an object of mystery and intrigue for many centuries. From the earliest solar observations the Sun has posed challenging questions to many areas of physics. This has never been more true than at present. There are many solar phenomena that remain unresolved, for example the solar dynamo and 11 year cycle, coronal heating, the solar flare trigger and particle acceleration mechanism and the acceleration of the solar wind.

The Sun displays a wide range of dynamic phenomena including flares, coronal mass ejections and coronal waves. In this thesis I study two aspects of solar flares that have been the subject of many investigations and discussions in recent times: homology and the non-thermal broadening of X-ray emission lines.

In this chapter I present an overview of the solar atmosphere and introduce the physics that governs its behaviour. I then give a brief description of the physics of spectral line formation and diagnostic techniques, followed by an overview of line broadening mechanisms. I then give a summary of the present understanding on some of the major issues in solar flare studies that will be referred to in subsequent chapters and I conclude with an introduction to the topics of homology and non-thermal line broadening describing their importance in our understanding of solar flares.
1.1 Solar Atmosphere

The Sun can be divided into two principal regions, the interior and the atmosphere. The interior consists of the core, where the fusion reactions occur, and the radiative and convective zones, named due to the dominant mechanism transporting heat outwards from the core.

The solar atmosphere is divided into four zones determined by the temperature \( T_e \) of the emitting plasma. These zones are the photosphere \( (T_e \approx 10^4 K) \), chromosphere \( (T_e \approx 10^5 K) \), transition region \( (T_e \approx 5 \times 10^5 K) \) and the corona \( (T_e > 10^6 K) \). These four zones have traditionally been thought to be layered, forming a plane parallel atmosphere with the temperature increasing and density decreasing radially from the solar surface (Figure 1.1).

![Figure 1.1: The temperature and density structure of the solar atmosphere.](image)

The photosphere is the visible 'surface' of the Sun. It is a very thin layer from which the bulk of the energy is radiated in the visible and Infra-Red (IR) wavelength ranges. The structure of the photosphere is dominated by the solar granulation. This structure takes the form of small, bright polygonal patches of order of 1100 km across that are surrounded by thin dark lanes. This granulation
pattern is formed by rising convection cells of hotter plasma from the hotter layers below, while the darker lanes are the descending currents of cooler material.

The magnetic field strength at the photosphere can be determined using the Zeeman effect; the splitting of certain spectral lines in the presence of a magnetic field (Chapter 1.3.3). In the hotter outer layers of the atmosphere the thermal width of the line smears out the Zeeman splitting, hence direct coronal magnetic field measurements cannot be obtained using this method. The field in the corona can be inferred from the plasma structures as will be shown in the next sub-section.

The chromosphere is the layer between the temperature minimum and the corona where the temperature begins to rise. The chromosphere can be seen as a pink flash at the moment just before totality during a solar eclipse. The emission from the chromosphere is dominated by the $\text{H}\alpha$ line at 6563Å which gives it the characteristic pink colour. The transition region lies between the chromosphere and the corona. In this thin layer the temperature increases rapidly from a few tens of thousand to greater than one million Kelvin.

The Solar Corona appears as a blue-ish white halo during total solar eclipses. It is an extremely hot tenuous layer that is highly structured by the magnetic field. The high temperature of the corona is evident from the emission lines it produces. The first identified emission line of a highly ionized species was that of a forbidden transition of $\text{Fe XIV}$ (Eldén 1942) later known as the ‘green line’. The presence of $\text{Fe XIV}$ implied the corona was hot $\approx 2 \times 10^6 K$ and because the transition was forbidden, the density in the corona had to be low. The details of the coronal heating mechanism are still presently disputed and is a strong topic of research, although it is widely believed that the heating is related to the magnetic field. Due to the high temperatures in the corona, emission is dominated by emission lines from highly ionized species and thermal Bremsstrahlung. These emissions are at frequencies that are absorbed by the Earth’s atmosphere, hence coronal emission observations must be made from space. Observations of the white light corona, visible during solar eclipses, can be accomplished using both ground based and
space borne coronagraphs. The white light emission is photospheric radiation that is scattered by electrons and dust grains.

Recent results from the Solar and Heliospheric Observatory (SoHO) are challenging this traditional picture of a plane parallel atmosphere stratified in temperature and density. Observations of coronal structures at a range of temperatures from chromospheric to coronal show loop structures present over a range of heights (e.g. Brekke et al. 1997; Fludra et al. 1997; Matthews and Harra-Murnion 1997; Harra-Murnion et al. 1999)

1.2 Coronal Structure

For a plasma contained in a magnetic field, the degree to which the magnetic field dominates over the plasma is given by the ratio of their respective pressures. This quantity is called the plasma $\beta$ and is given by,

$$\beta = \frac{16\pi nkT}{B^2}$$  \hspace{1cm} (1.1)

(Tandberg-Hanssen and Emslie 1988), where $n$ is the electron number density, $k$ the Boltzmann constant, $T$ the plasma temperature and $B$ the magnetic field strength. In the photosphere the plasma $\beta$ is high, hence the plasma motions shape the magnetic field structure. In the corona where the density is much lower, the magnetic field dominates over the plasma.

The temporal variations of the solar magnetic field are given by the induction equation (Priest 1982),

$$\frac{\partial B}{\partial t} = \nabla \times (\mathbf{v} \times \mathbf{B}) + \eta \nabla^2 \mathbf{B},$$  \hspace{1cm} (1.2)

where $\mathbf{B}$ is the magnetic field, $\mathbf{v}$ the plasma velocity and $\eta$ the magnetic diffusivity. The first term on the right hand side describes the advection of magnetic field lines carried by plasma motions, the second term describes the diffusion of the magnetic field through the plasma. The magnetic diffusivity is defined as $1/\sigma \mu$ where $\sigma$ is
the electrical conductivity and \( \mu \) the permeability of free space. For a fully ionized collisionally dominated plasma, like the corona, the magnetic diffusivity can be written as,

\[
\eta = \frac{m_e}{\mu n_e e^2 \tau_{ei}} = 5.2 \times 10^7 \ln \Lambda T^{-3/2} \text{m}^{-2}\text{s} \tag{1.3}
\]

where \( \ln \Lambda \) is the Coulomb logarithm and \( \tau_{ei} \) the effective electron ion collision time. The typical value for \( \eta \) in the solar corona is \( 10^9 T^{-3/2} \text{m}^{-2}\text{s} \).

The ratio of the advection and diffusion time-scales, in the induction equation (1.2) defines the magnetic Reynolds number,

\[
R_m = \frac{l_0 V_0}{\eta}, \tag{1.4}
\]

where \( l_0 \) and \( V_0 \) are typical length and velocity scales. A fundamental supposition of magneto-hydrodynamics is that electro-magnetic variations are non-relativistic, or quasi-steady, therefore

\[
V_0 \ll c, \tag{1.5}
\]

and \( V_0 = l_0 / t_0 \) (Priest 1982). Therefore, ‘typical’ scale lengths in the corona are those which satisfy equation 1.5. In terms of observations, large coronal structures may extend to \( 10^7 \text{m} \), but there evolutionary time-scales are on the order of hours, whereas brightenings on scales of a few seconds only occur in small compact loops, thus satisfying \( V_0 \ll c \).

In the solar atmosphere \( R_m = 10^6 \rightarrow 10^{12} \), therefore the diffusion term in the induction equation is negligible and the induction equation reduces to the perfectly conducting limit,

\[
\frac{\partial \mathbf{B}}{\partial t} = \nabla \times (\mathbf{v} \times \mathbf{B}). \tag{1.6}
\]

In this perfectly conducting limit the frozen-flux limit of Alfvén (1942) holds. This theory states that in a perfectly conducting plasma the magnetic field lines behave as if they move with the plasma. The magnetic field is therefore frozen into the plasma.

In the corona where the plasma \( \beta \) is small and the field frozen in, the magnetic field will confine and determine the emitting plasma structures. We can therefore...
use observations of Soft X-ray structures to study the morphology of the coronal magnetic field. However, it is not possible by this method to determine the field direction or strength.

1.3 Spectral Line Formation

In this section I describe the formation process of spectral emission lines in the solar corona and how, through the direct observations of these emission lines, we can estimate the temperature and emission measure of the emitting plasmas. Mechanisms that broaden spectral lines are also described. These processes form an integral part of the data analysis techniques for the Bragg Crystal Spectrometer (BCS) that were undertaken in Chapters 4 and 5.

1.3.1 Line Intensities

In the solar corona the electron density \( n_e \) is low, with typical values of \( 10^9 \text{cm}^{-3} \) in the quiet Sun to \( 10^{13} \text{cm}^{-3} \) in flares. Due to the low density the plasma is optically thin to the radiation it emits.

In this limit of low density the assumption of coronal equilibrium is applied. In coronal equilibrium the collisional excitation of an ionic species, from level \( i \) to \( j \), is balanced by the non-collisional process of spontaneous radiative decay, so that

\[
 n_i n_e C_{ij} = n_j A_{ji} = I_{ji},
\]

where \( n_i \) and \( n_j \) are the number densities in state \( i \) and \( j \) respectively, \( C_{ij} \) is the electron collisional excitation rate coefficient, \( A_{ji} \) the Einstein spontaneous emission rate coefficient and \( I_{ji} \) the photon emission rate per unit volume (\( \text{photons cm}^{-3} \text{s}^{-1} \)). Thus the photon emission rate is simply equal to the collisional excitation for this simple two level ion. From Equation 1.7 the ratio of the excited state to the ground state is

\[
 \frac{n_j}{n_i} = \frac{n_e C_{ij}}{A_{ji}},
\]
which in the low density conditions of the solar corona is always small compared with unity (Gabriel 1992).

For a simple two level ion the photon intensity \( I_i \) of a spectral line produced by an electron transition from an excited state \((i)\) to the ground state \((g)\) is given by

\[
I_i = \int_V n_g n_e C_{gi} dV \quad \text{(photons s}^{-1})
\]  

(1.9)

The quantity \( n_g \) can be expressed as the identity

\[
n_g = \left( \frac{n_g}{n_{ion}} \right) \left( \frac{n_{ion}}{n_z} \right) \left( \frac{n_z}{n_H} \right) \left( \frac{n_H}{n_e} \right) n_e,
\]

(1.10)

where \( n_g/n_{ion} \) is the fraction of ions in the ground state, assumed to be unity in the coronal approximation (Equation 1.8), \( n_{ion}/n_z \) is the fraction of ions in the particular ionization state for the transition under consideration, \( n_z/n_H \) is the relative abundance \( A_z \) of the ion relative to Hydrogen and \( n_H/n_e \) is the abundance of Hydrogen relative to free electrons \( \approx 0.8 \) in the solar corona.

Calculations of \( n_{ion}/n_z \) assuming ionization equilibrium have been carried out for numerous ion species (Arnaud & Rothenflug 1985, Arnaud & Raymond 1992). The time-scales to achieve ionization equilibrium have been tabulated for a variety of different astrophysical plasmas by Mewe (1987). For solar flare plasmas the effects of transient ionization, where there is significant departures from equilibrium, are only important for large temperature changes occurring on time-scales less than the relaxation time, \( t_{rel} \), of the plasma, which is given by,

\[
n_e t_{rel} \leq 10^{11} \rightarrow 10^{12}.
\]

(1.11)

In solar flares where the density is \( 10^{11} \rightarrow 10^{13} \) this relation gives relaxation times of a second or less; shorter than the integration times of the spectrometer (Chapter 3.3).

In the absence of any dominant non-thermal influences (e.g. a strong electric field) the transfer of energy and momentum to or from the free electrons is dominated by electron-electron collisions. This thermal distribution of electrons is the
Maxwellian distribution and is a function only of temperature. For a Maxwellian electron velocity distribution the collisional excitation coefficient can be expressed as,

\[ C_{gi} = \frac{8.63 \times 10^{-6} \gamma_{gi} e^{\nu/kT_e}}{\omega_g T_e^{-1/2}}, \]  

(1.12)

where \( T_e \) is the plasma temperature, \( \gamma_{gi} \) is the thermally averaged effective collision strength for that transition, which is tabulated for many ions (e.g. for He-like ions see Dubau (1994)) and \( \omega_g \) the statistical weight or degeneracy of the state (Mason & Monsignori Fossi 1994).

Substituting Equations 1.10 and 1.12 into Equation 1.9 gives,

\[ I_i = \frac{8.63 \times 10^{-6} \gamma_{gi} A_{2.8}}{\omega_g} \int_V \left( \frac{n_{ion}}{n_z} \right) T_e^{-1/2} e^{\nu/kT_e} n_e^2 dV. \]  

(1.13)

The temperature dependent terms are incorporated into the contribution function or emissivity,

\[ G(T) = \left( \frac{n_{ion}}{n_z} \right) T_e^{-1/2} e^{\nu/kT_e}. \]  

(1.14)

If the emitting plasma is assumed to be isothermal then the \( G(T) \) function can be removed from the integral and the intensity in the line becomes,

\[ I_i = \frac{8.63 \times 10^{-6} \gamma_{gi} A_{2.8} G(T)}{\omega_g} \int_V n_e^2 dV. \]  

(1.15)

The quantity \( \int_V n_e^2 dV \) is the emission measure of the emitting plasma. The line intensity is thus a function of electron temperature and emission measure.

### 1.3.2 Dielectronic Satellites

Dielectronic satellite lines arise as a result of a two step process. Figure 1.2 shows a simplified energy level diagram of a He-like ion together with the resonance levels of the type 1s2s2p which lie just below the first excited state 1s2p. The first step in the process is the capture of an electron from the continuum by the ion (denoted \( C_s \) in Figure 1.2). Some of the energy is used to simultaneously excite one of the bound electrons resulting in a doubly excited state. The excited ion
can either decay by the reverse process of auto-ionization \((A_a)\) or radiatively decay \((A_r)\) to the ground state \((1s^22s)\) of the Li-like ion. The excited bound electron (or spectator electron) acts to perturb the energy of the stabilizing transition, of the other excited electron, so that the emitted line appears as a satellite to the principal transition. The magnitude of the perturbation is dependent upon the energy level of the spectator electron (denoted \(nl\)). As the value of \(n\) increases the influence of the spectator electron decreases due to a diminishing shielding effect.

\[
\begin{align*}
\text{He-like} & \quad \text{Li-like} \\
\hline
1s^2 & \quad 1s^2 2s \\
\hline
1s2p & \quad 1s2s2p \\
\hline
\text{E}_r & \quad \text{E}_s
\end{align*}
\]

Figure 1.2: Energy level diagram for a He-like ion showing the adjacent satellite energy levels of the Li-like species (from Gabriel 1992).

The emission rate of dielectronic satellite lines is given by,

\[
\begin{align*}
I_s &= n_i n_e C_s, \quad \text{(1.16)} \\
C_s &= \frac{2.06 \times 10^{-16} e^{-E_s/kT_e}}{T_e^{3/2} \omega_g} F_2(s), \quad \text{(1.17)} \\
F_2(s) &= \omega_s A_a B_r, \quad \text{(1.18)} \\
B_r &= \frac{A_r}{A_a + A_r}, \quad \text{(1.19)}
\end{align*}
\]
CHAPTER 1. INTRODUCTION

(Gabriel & Paget 1972; Gabriel 1972; Bely-Dubau et al. 1982). $I_s$ and $C_s$ are the photon emission rate and collisional excitation rate of the satellite line and $B_r$ represents the branching ratio, the probability that doubly excited 3-electron state decays by radiation. Recalling Equation 1.7 and 1.12 we can write the ratio of the He-like resonance transition to the satellite transition as,

$$\frac{I_s}{I_r} = \frac{2.39 \times 10^{-11} F_2(s)}{T_e T_{gi}} \times e^{\left(\frac{E_s - E_r}{kT_e}\right)},$$  \hspace{1cm} (1.20)

$$\Rightarrow \frac{I_s}{I_r} \propto \frac{e^{\left(\frac{E_s - E_r}{kT_e}\right)}}{T_e},$$  \hspace{1cm} (1.21)

hence this ratio is a function of $T_e$ only and can be used to determine the temperature of the emitting plasma. When $A_a$ is small, the satellite is weak and the method does not work. It is therefore necessary to use an allowed auto-ionization, that is with a rate in the region of $10^{13}s^{-1}$. Because the two transitions are very close in energy, the difference $E_s - E_r$ is small compared with $kT_e$ hence the main sensitivity comes from the $T_e^{-1}$ term. Thus this method is excellent for determining temperatures when they are high enough to produce He-like ionic species (Gabriel 1992).

These diagnostic techniques are employed in the analysis of spectral data from the Bragg Crystal Spectrometer (BCS) and are utilized in the determination of the non-thermal broadening of emission lines in Chapters 4 and 5.

1.3.3 Line Broadening Mechanisms

Natural Width

An atom completely isolated from its neighbours and free from any disturbing forces will radiate a sharp spectral line, broadened only as a result of Heisenberg’s uncertainty principle, which states that,

$$\Delta E \Delta t \geq h$$  \hspace{1cm} (1.22)
where $\Delta E$ is the energy uncertainty, $\Delta t$ the uncertainty in time and $\hbar$ Planck’s constant. This broadening is known as the natural width of the line.

For an electron in the ground state, which cannot move to a lower energy, $\Delta t$ is large, hence $\Delta E$ is small and the line is very sharp. For an electron in an excited state $\Delta t = 1/A$, where $A$ represents all rates of leaving the state, therefore $\Delta E = A\hbar$. For example, the Lyman $\alpha$ transition at 1216\,Å, $A = 4 \times 10^8$\,s$^{-1}$, therefore $\Delta E = 2.65 \times 10^{-25}$\,J and the line is broadened by only $2 \times 10^{-4}$\,Å.

**Doppler Width**

The primary source of line broadening in spectral lines is due to the Doppler effect. Thermal, turbulent and unresolved wave motions produce random ionic motions along the line of sight, resulting in random frequency shifts of the emitted radiation. For a thermal plasma the one dimensional line of sight velocity distribution of ions is Maxwellian,

$$f(v)dv = \frac{1}{v_0 \sqrt{\pi}} e^{-\left(\frac{v}{v_0}\right)^2} dv$$

(1.23)

where

$$v_0 = \sqrt{\frac{2kT}{m}}$$

(1.24)

where $k$ is Boltzmann’s constant, $T$ the temperature of the plasma and $m$ the mass of the ion. The resulting emission line profile is given by,

$$I(\lambda) = I_0 e^{-\left(\frac{\Delta\lambda}{\Delta\lambda_0}\right)^2}$$

(Zirin 1988) where $\Delta\lambda$ is the wavelength distance from the centre of the line and $\Delta\lambda_0$ the thermal width of the line is given by,

$$\Delta\lambda_0 = \frac{w_0 \lambda}{c} = \frac{\lambda}{c} \sqrt{\frac{2kT}{m}}.$$  

(1.26)

The width is therefore dependent upon the square root of the temperature. Using the Lyman $\alpha$ emission line again as an example, the thermal width of the line at a temperature of $5 \times 10^4$\,K is 0.1\,Å, three orders of magnitude greater than the natural width.
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External Fields

If the emitting ion is subjected to an external field when the electron transition occurs, this can also give rise to line broadening. Collisional broadening results when the emitting ion experiences the electric field of a nearby particle when the electron transition occurs. The electric field perturbs the orbit of the electron altering its energy. Because the collisions are random, so are the perturbations and the line is broadened. Collisional broadening scales as the square of plasma density and can be important in the wings of spectral lines emitted in the solar atmosphere (Phillips 1992).

In the presence of a magnetic field the electron energy levels can become separated in energy. The orbiting electron can be thought of as a small electric current with an associated magnetic field. If this field is aligned with the external field the orbit has a lower energy and vice versa. The line appears as a doublet with the splitting of the energy levels proportional to the external magnetic field strength. The two components are also oppositely circularly polarized.

Although the Zeeman effect produces a splitting of the line, if the resulting components are close together and are already broadened by other mechanisms they may overlap to effectively produce another broadening mechanism.

In Chapters 4 and 5 of this thesis I examine the line width of the Ca XIX resonance line at 3.174Å. The width of the line is greater than that excepted from a thermal plasma due to Doppler broadening. The excess width is known as the non-thermal broadening of the emission line. In Chapters 4 and 5 I examine the behaviour of non-thermal broadenings in solar flares to determine the location of the source and mechanism for their generation.
1.4 Solar Flares

Solar flares are sudden and rapid releases of magnetic energy that occur in the solar atmosphere. The energy is explosively released into various forms, including accelerated particles, heating of plasma, bulk acceleration of plasma and enhanced radiation fields (Tandberg-Hanssen & Emslie 1988). Flares can have durations ranging from a few minutes to many hours and range in brightness over many orders of magnitude; up to $\approx 10^{23}$ ergs over the flare duration. Flares most commonly occur in areas of strong and complex magnetic fields, i.e. in active regions. This association with active regions causes the frequency of flare occurrence to vary with the solar cycle.

1.4.1 Classification

The strength of a solar flare is measured by the amount of radiation it emits. This is typically measured in two wavelength bands, Hα and Soft X-rays (SXR). The Hα wavelength flare classification is measured by the Importance scheme. This classifies flares according to their area (corrected for fore-shortening) and brightness. The brightness of flares can be bright (b), normal (n) and faint (f) (Table 1.1).

<table>
<thead>
<tr>
<th>Importance</th>
<th>Flare Area ($10^6 km^2$)</th>
<th>Flare Brightness</th>
</tr>
</thead>
<tbody>
<tr>
<td>S</td>
<td>&lt; 300</td>
<td>f, n, b</td>
</tr>
<tr>
<td>1</td>
<td>300 → 750</td>
<td>f, n, b</td>
</tr>
<tr>
<td>2</td>
<td>750 → 1850</td>
<td>f, n, b</td>
</tr>
<tr>
<td>3</td>
<td>1850 → 3650</td>
<td>f, n, b</td>
</tr>
<tr>
<td>4</td>
<td>&gt; 3650</td>
<td>f, n, b</td>
</tr>
</tbody>
</table>

Table 1.1: Hα flare classification scheme.
In SXRs flares are classed according to the GOES classification scheme. The 
*Geostationary Operational Environmental Satellite* (GOES) is a full Sun SXR flux 
monitor that operates in two wavelength ranges, 0.5 → 2Å and 1 →8Å. The 
GOES flare classification is dependent upon the flare’s peak X-ray flux in the 
1 → 8Å channel. The scale varies from $1 \times 10^{-8}$ to $1 \times 10^{-4} W m^{-2}$, with the letters 
A, B, M and X representing whole powers of ten (Table 1.2).

<table>
<thead>
<tr>
<th>GOES Class</th>
<th>Peak Flux ($W m^{-2}$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>A</td>
<td>$10^{-8}$</td>
</tr>
<tr>
<td>B</td>
<td>$10^{-7}$</td>
</tr>
<tr>
<td>C</td>
<td>$10^{-6}$</td>
</tr>
<tr>
<td>M</td>
<td>$10^{-5}$</td>
</tr>
<tr>
<td>X</td>
<td>$10^{-4}$</td>
</tr>
</tbody>
</table>

Table 1.2: GOES flare classification scheme.

1.4.2 Flare Physics

There are numerous models of solar flares that can explain the many observational 
characteristics of individual events. Each of these models, however, incorporate 
some basic features that are common to all flares. In this section I describe some 
of these common features, which are shown in Figure 1.3, a proposed model for a 
single loop flare (from Masuda 1994).

Energy Release

The energy release in solar flares is believed to be a result of magnetic reconnection. 
Magnetic reconnection was first proposed to be the flare energy release mechanism
Magnetic reconnection is defined as the change in connectivity of plasma elements in a magnetic field. It is typically visualized as the breaking of two adjacent oppositely directed field lines which subsequently join to produce two new field lines (Figure 1.4). This process occurs in regions where the diffusion term in the induction equation (Equation 1.2) dominates and results in a change in the magnetic field morphology, converting the magnetic energy into heat and kinetic energy.
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Random footpoint motions in the photosphere, where the plasma pressure dominates the magnetic pressure, stresses and shears the magnetic field, storing magnetic energy in the corona. This stored magnetic energy is released in the corona by magnetic reconnection as the field relaxes back to a potential state.

![Figure 1.4: Schematic diagram of magnetic reconnection. Field lines C→A and B→D are broken and join to form field lines B→A and C→D.](image)

**Energy Transfer**

The energy released by magnetic reconnection in the corona is transported down to the chromosphere. Historically the dominant transfer mechanism was believed to be energetic electrons only. The electric field setup between the electron depleted corona and the chromosphere would drive a return current that replenished the corona, providing a source of electrons. However, gamma ray observations of nuclear emission lines, emitted from excited nuclei after proton impact, have shown that the energy contained in accelerated protons is comparable with that of electrons (Miller *et al.* 1997, Mandzhavidez & Ramaty 1996).

There are two models for electron energy transport: *Thermal* and *Non-thermal*. In the thermal model the electron energy distribution is Maxwellian and results from the presence of a hot source (≈ 10⁶K) in the corona generated by the reconnection process. Heat conduction transports the electrons to the chromosphere.

In the non-thermal model the electrons are energized above their Maxwellian values. The process by which these electrons are accelerated is presently unresolved.
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and is a major topic of flare research (e.g. Miller et al. 1997). There are a number of observational aspects of solar flares that support the presence of non-thermal electrons as the dominant transfer mechanism. These include the very close timing of double footpoint HXR footpoints (Sakao 1994), microwave spectra which do not match that of a thermal source (Takakura 1967; Švestka 1976) and coincidental timing of microwave and HXR emission (Crannell et al. 1978, Dennis 1988).

When the electrons reach the chromosphere they produce two principal observable effects: Bremsstrahlung emission in HXR from the footpoints of the magnetic loop and chromospheric evaporation due to the explosive heating of the chromosphere.

1.4.3 Footpoint Emission

Energetic electrons lose their most of their energy through collisions with ambient electrons. These collisions heat the target atmosphere. When an energetic electron experiences a close encounter with a proton or ion the deceleration of the electron results in the emission of a photon and is termed Bremsstrahlung radiation. This photon can be comparable in energy with the incoming electron and hence result in a photon of HXR wavelengths. Bremsstrahlung radiation produced in these collisions is dependent upon the nature of the ‘target’ seen by the electrons, this can be either thin or thick. In the thin target regime no significant modification of the electron spectrum occurs, whereas in the thick target scenario the injected electrons are completely stopped, or more accurately thermalized in the Bremsstrahlung source.

In the majority of solar flare models the electrons are injected at the top of the loop and propagate along the magnetic loop, a low density environment. When they encounter the denser layers of the upper chromosphere they can interact with the ambient protons and produce thick-target HXR emission at the loop footpoints. The spectrum of the HXR emission can, in principle, be used to calculate the
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Footpoint emission is also observed in the Ultraviolet (UV) and Extreme Ultraviolet (EUV) range (10 \( \rightarrow \) 1030Å; Kane et al. 1980). The link between the UV and HXR flux was established in general by the relation in peak energies (Kane & Donnelly 1971), the simultaneity of individual peaks in HXR and O\( \text{v} \), Si\( \text{iv} \), O\( \text{v} \) and UV continuum (Woodgate et al. 1983, Cheng et al. 1988), the relation of energies emitted in HXR and O\( \text{v} \) and their similar rise and fall (Poland et al. 1984) and the spatial co-incidence of HXR and O\( \text{v} \) sources (Cheng et al. 1981; Duijvemann et al. 1982; Machado et al. 1982; Poland et al. 1982; Canfield et al. 1986). The EUV/HXR flux ratio exhibits a centre to limb variation which is interpreted as the EUV emission emanating from a trench or well in the optically thick chromosphere (Donnelly & Kane 1978). The impulsive EUV emission is believed to arise as a result of the conductive heating of the lower atmospheric layers as a result of the deposition of energy by the energetic electrons that also generate the HXR burst (Brown 1973; Emslie & Nagai 1984; Canfield et al. 1986). Although the temporal correlation between UV and HXR flux is striking (Figure 1.5) the spatial correlation is less grounded. HXR imaging from SMM on which the above references refer was rather poor, 32 arc-second resolution, and although all HXR kernels had accompanying UV emission the reverse was not always the case. Hence, the notion that EUV/UV emission can be used as a proxy for HXR emission and consequentially energetic electron deposition sites, should be used with caution.

1.4.4 Chromospheric Evaporation

Chromospheric evaporation was first proposed by Neupert (1968) to explain the different numbers of electrons producing microwave emission and Fe\( \text{xxv} \) line emission during flares observed by the third Orbiting Solar Observatory. The mismatch
in electron populations led Neupert to propose that the SXR material was chromospheric in origin.

The rate at which an optically thin plasma radiates energy, known as the radiative loss function, is a function of temperature and has been calculated by several authors (e.g. Pottasch 1965; Cox & Tucker 1969; Raymond et al. 1977; Rosner et al. 1978; Landini & Monsignori Fossi 1990; Landini & Landini 1999). If the rate at which energy is supplied to the chromosphere exceeds the rate at which it can be radiated away then the plasma is heated. The form of the radiative loss function from $10^5 \rightarrow 10^7 K$ (Figure 1.6) is such that an increase in temperature leads to a decrease in radiated energy. This results from the diminishing contributions of line emissions as more species are ionized (Figure 1.6). The heated plasma can therefore undergo a radiative instability and can be brought from chromospheric to coronal temperatures on very short time-scales, less than the hydrodynamic time-scale. This rapid heating causes a strong pressure gradient so that the hot dense
plasma quickly expands and fills the loop. This hot dense material emits SXR as it cools and because it is confined within the magnetic field, forms the SXR loops that are characteristic of a flare.

![Graph showing radiative loss function for an optically thin plasma.](image)

Figure 1.6: The radiative loss function for an optically thin plasma, from Landini & Landini (1999)

Blue shifts observed in SXR spectra are believed to be direct signatures of chromospheric evaporation in solar flares (e.g. Antonucci et al. 1982, 1984; Doschek et al. 1986). As the plasma expands up the loop the outward radial velocity produces a blue shift in the spectrum. This blue shift has been observed to exhibit a centre to limb variation implying the flows are indeed radial (Mariska, Doschek & Bentley 1993). The close temporal correlation of the HXR burst intensity to the blue-shifted emission (Bentley et al. 1994; Doschek, Mariska & Sakao 1996) is further evidence that the observed flows are a signature of chromospheric evaporation.

Conservation of momentum dictates that the up-flow of plasma into the corona should be balanced by an equal downward flow termed a chromospheric condensation (Fisher et al. 1985c). Chromospheric condensations consist of cool dense plasma and are observed as Hα red-shifts during the impulsive phase (Tang 1983,
Ichimoto & Kurokawa 1984). Such a momentum balance has been observed in flares (Canfield et al. 1990; Wülser et al. 1994). Wülser et al. (1994) also showed that the SXR blue-shift originated from the same location as the Hα red-shifts, hence adding further support to the evaporation theory.

Although there is much evidence to support the presence of chromospheric evaporation in solar flares there are some features that cannot be explained by the present model. For example, very bright loop tops are observed in some flares throughout the duration (Seely et al. 1994; Feldman et al. 1995; Feldman & Seely 1995) suggesting that the energy deposition occurs in a small volume at the top of the flaring loop and that hot plasma is confined to the flare loop top. Also, blue shifted emission is sometimes observed before the detection of any HXRs (Doschek 1990; Plunkett & Simnett 1994), although this could be due to inadequate sensitivity in HXR detectors. Numerical modeling of electron heated flares suggests that SXR emission should be dominated by up-flowing plasma (Mariska, Emslie & Li 1989; Antonucci et al. 1993), however observations of line spectra with Bragg crystal spectrometers also show the presence of a strong stationary component (Doschek et al. 1986).

1.4.5 Solar Flare Homology

Definition

Solar flare homology was first observed at Hα wavelengths by Smith and Smith (1963). Homologous flares are a series of flares almost identical in location and morphology that occur repeatedly in the same active region (Zirin 1988). The definition of flare homology given in the literature (e.g. Woodgate et al. 1984) states that:

a) members of a series must have the same main footpoints, allowing for horizontal motions of the underlying structures, as defined by Hα or EUV kernels.

They must also share the same general shape in the main phase, essentially
including similar loop structures, as defined by Hα ribbons or SXR images, or that,

b) radio bursts in a series must show similar dynamic spectra.

Implications

Homologous flares are a particularly interesting phenomena that pose some challenging questions about the conditions that lead to flaring. For example, are the pre-flare conditions destroyed by the first flare? If they are, then how are these conditions rebuilt? If they are not destroyed, then what mechanism is responsible for the termination of the first event and the subsequent triggering of the second (Martres et al. 1984)?

The presence of homologous flares puts heavy constraints on flare models. For example, the storage and energy release mechanisms, the origins of the flare energy, the lifetime of the magnetic configuration, the start process and the time-scales for these features. Thus a homologous flare series provides an interesting tool to study critical conditions required for the production of solar flares.

The majority of models that explain the occurrence of homologous flares have done so in a way which preserves the pre-flare conditions of the first flare (Woodgate et al. 1984; Shibata 1998; Choe and Cheng 2000). These models adopt a scenario in which a single magnetic bipole or bipolar arcade is continually sheared allowing successive reconnection events to occur within the same magnetic environment and hence produce very similar flares.

Figure 1.7 shows the schematic model of Woodgate et al. (1984). The magnetic configuration is a simple bipolar arcade within which is supported a filament. Horizontal photospheric flows with a differential velocity occur on either side of the photospheric neutral line, shearing the arcade structure. This shear concentrates the field over the neutral line and causes it to rise through magnetic buoyancy. The rising field promotes the ejection of the filament and the flare impulsive phase
begins as a consequence. After the flare the reformation of the filament recreates similar pre-flare conditions. The continuation of photospheric flows rebuilds the shear in new magnetic structure, providing the energy for the next member of the homologous series.

**Previous Observations**

Previous studies of homologous flares have centred around Hα observations. Martres (1989) classified three types of homologous flare according to time separation. The time separation classes are hours, days and rotations. Martres (1989) notes some important features of homologous flares occurring hours apart, namely that a) members of a flare series will most likely have a similar Hα importance, implying the importance is related to the flare site, and b) that the presence of a large spot is a necessary condition for a homologous flare series. A homologous flare series observed in EUV (Fe xxı and O v) and X-rays from the Solar Maximum Mission (SMM) has also been observed (Cheng and Pallavicini 1987). Their results indicated that homology is best recognized in high temperature emissions (Fe xxı and SXR).

Morita et al. (1998) studied a series of homologous flares observed with SXT that occurred on the 21st, 24th and 27th of February 1992. As the active region in which all the flares occurred rotated around from the east limb the three homologous flares were observed from different line of sight angles. Hence they were able to see essentially the same structure viewed from largely different angles. This enabled them to determine the 3D-structure of the flare. Morita et al. (1998) also studied magnetogram data at the times of the three flares and noted that there were no systematic changes such as footpoint shearing. Such an observation is in contradiction to the model of Woodgate et al. (1984) and suggests that in this instance a different mechanism, other than photospheric shearing, was the cause of the homology.

In Chapter 3 I study a series of homologous flares and show that these flares
Figure 1.7: Woodgate et al. (1984) homologous flare model. (1) the pre-flare state, (2) filament eruption and impulsive phase, (3) Re-establishment of similar pre-flare conditions. (a) gives the view from above the arcade, (b) show the radial extent of the arcade. The dotted line indicates the magnetic neutral line and double headed arrows indicate flow directions.
are better represented by a model invoking emerging magnetic flux rather than the continual shearing of a single or group of magnetic structures. The implications for homologous flares is then discussed.

1.4.6 Non-Thermal Broadening in Solar Flares

Non-thermal broadening is defined as the difference between the Doppler temperature ($T_d$) and the plasma temperature ($T_e$). In an ionized He-like species these temperatures are derived from the width of the main resonance line and the ratio of the main resonance line and the dielectronic satellite lines (Chapter 1.3.2) respectively. Because Doppler broadening is the only mechanism in coronal conditions capable of producing the large excess widths observed (Chapter 1.3.3), this temperature difference is often expressed as a non-thermal velocity ($V_{nt}$) where,

$$V_{nt} = \sqrt{\frac{2k(T_d - T_e)}{m_i}},$$

where $k$ is the Boltzmann constant and $m_i$ the mass of the ion under consideration.

The nature and location of the source of the non-thermal broadenings is still unknown as indeed is its role in the flare process. For example, are they a direct signature of the flare energy release process, or a hydrodynamic response of the solar atmosphere to the injection of flare energy?

Studies of $V_{nt}$ characteristics have been carried out by a number of spaced based instruments, including Skylab, P78-1 and the Solar Maximum Mission (SMM). Results from these studies (e.g. Doschek et al. 1986) showed that they can be approximated as Gaussian broadening with peak values ranging from $150\, km\, s^{-1}$ to $300\, km\, s^{-1}$. They are present before the peak of the hard X-ray (HXR) flux and diminish to between $0\, km\, s^{-1}$ and $50\, km\, s^{-1}$ by the time of SXR maximum. There is no correlation observed between maximum $V_{nt}$ and position on the disk. There is also evidence for the presence of $V_{nt}$ in C IV ($T_{\text{max}} = 1 \times 10^5\, K$) spectral lines at $\approx 50\, km\, s^{-1}$. 


In a large study of small flares Harra-Murnion et al. (1997) examined the variation of $V_{nt}$ with electron temperature, GOES classification, duration, rise time and source size of the SXR event. Their results showed that $V_{nt}$ is independent of flare size, complexity and intensity of HXR bursts, but there is a weak dependence on duration and rise time. The longer the rise time the lower the value of $V_{nt}$. There is also a trend of increasing $V_{nt}$ with electron temperature.

In order to determine the location of the source of $V_{nt}$, studies of occulted limb flares, in which the flare footpoints are obscured by the solar disk, have been undertaken (Khan et al. 1995; Mariska, Sakao & Bentley 1996; Mariska & McTiernan 1999). Khan et al. (1995) and Mariska & McTiernan (1999) showed that for partially occulted limb flares the measured $V_{nt}$ was similar to that observed for disk flares, indicating that the source of the $V_{nt}$ could not be the flare footpoints since these were occulted by the limb. Khan et al. (1995) concluded that the source of the $V_{nt}$ was at the loop top or that $V_{nt}$ is the same throughout the loop. These results are contradictory to those of Mariska, Sakao & Bentley (1996) who, using a smaller data set, showed a tendency for occulted flares to have a lower $V_{nt}$.

Doschek et al. (1986) and Mewe et al. (1985b) have suggested that the observed $V_{nt}$ can be explained by the fact that full Sun Bragg crystal spectrometers observe an integrated spectrum of plasma moving at a range of velocities. Fludra et al. (1989) showed a weak correlation between $V_{nt}$ and the measured blue shift velocity and Mariska, Doschek and Bentley (1993) showed a correlation between line width and the line centroid shift, but no centre to limb variation in $V_{nt}$ was found. These results indicate that chromospheric evaporation could account for at least some of the observed $V_{nt}$ particularly after the flare impulsive phase.

Recent work by Ding et al. (1999) using 2D spectra of a resolved flare loop has shown that the Hα profiles are more broadened at the loop top than anywhere else along the loop. However, although the Hα loops are believed to cool from SXR loops, this result should not be taken as true for SXR loops, since flare loops in Hα are often only seen when the SXR non-thermal broadenings have decayed.
Knowledge of how the values of $V_{nt}$ develop over time can also place stringent criteria on the location and generation of the $V_{nt}$. Alexander et al. (1998) studied the relationship between the peak times in $V_{nt}$ and the HXR flux and showed that the peak of the $V_{nt}$ occurred before the maximum in hard X-rays. Mariska & McTiernan (1999) in a similar study of a larger sample observed that, for the majority of the events they studied, the peak in $V_{nt}$ occurs after the first significant HXR peak, although the opposite behaviour was seen in a minority of events. The attainment of high values of $V_{nt}$ early in a flare, a similar result from both studies, is more indicative of plasma turbulence rather than hydrodynamic flows as the source of $V_{nt}$ (Alexander et al. 1998).

In this thesis I have conducted two studies in order to advance our knowledge on the location of the source of $V_{nt}$ (Chapter 4) and the understanding of the temporal behaviour of $V_{nt}$ (Chapter 5). In Chapter 4 I show that the location of the source of $V_{nt}$ is in the flare loops, not at the footpoints and in Chapter 5 I show that the time evolution of $V_{nt}$ is not common in all flares but is dependent upon the shape of the Hard X-ray burst.
Chapter 2

Solar Satellites and Instrumentation

There is presently a wealth of excellent dedicated solar instrumentation, both space borne and ground based, providing quality data for the solar community. Free from Earth’s obscuring atmosphere, satellite based instrumentation can potentially observe the Sun in any wavelength band at instrument determined resolution with uninterrupted coverage. Studies from the Solar Maximum Mission showed the importance of multi-wavelength observations in the understanding of solar flares. This thesis is based on multi-wavelength observations and has utilized data from many space based instruments. In this chapter I describe in detail the satellites and instrumentation that were utilized to provide the data and also outline some data analysis techniques.

2.1 Yohkoh

The SOLAR-A satellite (Ogawara et al. 1991) is a collaborative project between many institutions in Japan, the United States and the United Kingdom. It was launched by the Institute of Space and Astronautical Science (ISAS), Japan on 30th August 1991 from the Kagashima Space Center. Following the ISAS tradition
after launch it was renamed *Yohkoh*; the Japanese for Sunbeam. The primary scientific goals of *Yohkoh* are to study solar flares and other high energy phenomena in the X-ray and γ-ray regimes. To accomplish this *Yohkoh* carries a coordinated set of instruments comprising the Hard X-ray Telescope (HXT), the Soft X-ray Telescope (SXT), the Bragg Crystal Spectrometer (BCS), and the Wide Band Spectrometer (WBS). The WBS is further divided into the Hard X-ray Spectrometer (HXS), the Soft X-ray Spectrometer (SXS), the Gamma Ray Spectrometer (GRS) and the Radiation Belt Monitor (RBM).

The spacecraft was placed in a quasi circular orbit (eccentricity $\approx 0.02$; apogee 792km; perigee 517km) at an altitude of $\approx 600$km. The orbit is inclined at $31^\circ$ and has a period of 97 minutes. Due to the low inclination angle, the spacecraft passes into the Earth’s shadow for approximately 40 minutes during each orbit. This time period is known as ‘spacecraft night’. Ordinarily the spacecraft orbit is below the terrestrial radiation belts and the spacecraft functions normally. However the terrestrial radiation belts are anomalously low over the south Atlantic due to the offset of Earth’s magnetic axis with respect to the rotation axis. This region is known as the South Atlantic Anomaly (SAA). During spacecraft passage through the SAA the flux of energetic particles encountered by the spacecraft drastically increases and to protect the instruments the high voltage power supplies are switched off.

A schematic diagram of the *Yohkoh* spacecraft is shown in Figure 2.1. The dimensions of the spacecraft are $1000 \times 1000 \times 2000$mm. The two solar panels on each side measure $1500 \times 2000$mm each. The spacecraft consists of essentially seven panels. The centre panel houses the two large imaging telescopes, SXT and HXT, and the BCS and acts as an optical bench. The top panel facing the Sun holds the WBS detectors and has aperture windows for the other instruments.

Since *Yohkoh* takes high spatial resolution images as well as spectra, precise control of the spacecraft attitude is essential. *Yohkoh* is stabilized along all three axes, with the Z-axis pointed at the centre of the Sun with a stability of the order
of 1 arc-second $s^{-1}$. The Y-axis is directed toward celestial north. The attitude control system consists of momentum wheels, magnetic torques and control moment gyros as the actuators. Two Sun sensors and a star tracker, as well as geomagnetic sensors are used to determine the spacecraft pointing relative to the solar direction and the ecliptic plane respectively.

2.1.1 Soft X-ray Telescope

The Soft X-ray Telescope (Tsuneta et al. 1991) is a grazing incidence telescope capable of imaging the whole Sun in Soft X-rays (SX) with good spatial and temporal resolution. Plasma temperatures and densities can be obtained by selecting various metallic entrance filters that have different wavelength dependent transmission functions. Originally a white light telescope allowed accurate alignment of
SXR structures with optical features, however this failed after one year of operation in 1992.

The main scientific objectives of SXT are centred around flare studies. SXT images the SXR emitting plasma that is confined by magnetic fields, hence allowing us to observe the magnetic field topology.

In design concept the SXT is a very simple instrument. It has a fixed focus and comprises a shutter, dual filter wheels and two co-aligned imaging elements: a mirror for X-rays and a lens for visible light (Figure 2.2). A commandable door behind the lens is used to exclude visible light from the telescope when it is undesired. The same Charge Coupled Device (CCD) is used for both X-ray and visible light images. The angular pixel size of the CCD (2.45 arc-seconds) is approximately equal to the angular resolution of both the optical and X-ray telescopes. The CCD field of view is 42×42 arc-minutes, hence each exposure
images the whole Sun. However, tight telemetry budgets require the transmission of only parts of the CCD and/or the binning of the image. On chip pixel summation has three modes 1×1 (2.45′′; Full resolution), 2×2 (4.9′′; Half resolution) and 4×4 (9.8′′; Quarter resolution). The data is also compressed before being telemetered to the ground.

Through the implementation of various wavelength dependent transmission filters (Table 2.1.1) SXT is able to calculate the temperature and emission measure in each pixel of an image. The theory and technique behind this process is described in Chapter 3.1 along with the inherent caveats.

### 2.1.2 Hard X-ray Telescope

The Hard X-ray Telescope (HXT; Kosugi et al. 1991) is a Fourier imaging telescope which images simultaneously in four energy bands; 14 → 24keV (LO channel), 24 → 35keV (M1 channel), 35 → 57keV (M2 channel) and 57 → 93keV (HI channel). HXT is the first instrument capable of imaging solar flares at energies greater than 30keV. In flare mode the telescope makes observations every 0.5s and has a spatial resolution of 5 arc-seconds.
Since HXT images at such high energies with some spectral resolution its principal scientific goals are to study the role of energetic particles in solar flares, to try to ascertain the acceleration site and mechanism, propagation, confinement and deposition of these energetic particles during solar flares.

Figure 2.3: Schematic Diagram of HXT (from Kosugi et al. 1991).

Figure 2.3 shows a schematic diagram of the layout of HXT. HXT consists essentially of three parts: the collimator section (HXT-C), the detector assembly (HXT-S) and the electronics (HXT-E). The collimator consists of a metering tube 1400mm in length with X-ray grid plates at both ends. Each grid plate is an assembly of 64 subcollimator grids made of 0.5mm thick tungsten. In addition the front grid holds the filters and lens for the optical aspect system.

HXT-S comprises 64 identical non-position sensitive detectors. These are arranged in pairs on the detector assembly in an arrangement that optimizes image
reconstruction. Each pair of detectors measures a pair of complex Fourier components of the HXR source map. To obtain the solar flare HXR image it is necessary to perform, in principle, a two-dimensional inverse Fourier transform (image reconstruction) of the spatially modulated photon count data (Sakao 1994). However, HXT has fewer detector pairs than are needed and the angular transmission functions are not sinusoidal. Therefore a direct inverse transform is not feasible without producing significant noise and spurious signals in the recovered source map. Actual methods that perform the inversion were developed for the reconstruction of images from incomplete sets of measurements and were modified for the specifics of HXT. The two most favoured algorithms for HXT image reconstruction are the Maximum Entropy Method (MEM, Gull & Daniell 1978; Willingdale 1981; Sakao 1994; Sato et al. 1999) and Fixons (Metcalf et al. 1996). The MEM reconstruction procedure that was used throughout this thesis is explained in detail in Chapter 3.2.

HXT can also act as a wide band spectrometer. Data from all 64 detectors in one energy channel can be summed to provide a HXR flare spectrum with data points centred at 18, 28, 43 and 73keV. This spectrum can then be fitted with either a power law or thermal spectrum to determine the photon spectrum. This can then, in principle, be inverted to give the electron population that generates the HXR emission.

2.1.3 Bragg Crystal Spectrometer

The Bragg Crystal Spectrometer (BCS; Culhane et al. 1991; Lang et al. 1993) employs four bent Germanium crystals to view the whole Sun in four discrete wavelength bands covering the resonance line complexes of S\textsc{xv}, Ca\textsc{xix}, Fe\textsc{xxv} and Fe\textsc{xxvi}. The BCS has a sensitivity increase of approximately tenfold over previous similar instruments. This enhanced sensitivity provides greater temporal resolution and allows measurements to be taken at earlier stages during a flare.

The principal scientific objective of the BCS is the study of hot plasma in solar
flares. Since BCS observes individual spectral lines it is able to make measurements of line of sight bulk plasma velocities and non-thermal velocities evident from spectral line shifts and line broadening respectively, both of which can be observed during the impulsive phase of solar flares. Plasma diagnostics are achievable through the comparison of specific line strengths (Chapter 1.3.2) that can be used to determine the plasma temperature and consequently the emission measure.

Figure 2.4: Schematic diagram of (a) a flat scanning crystal spectrometer and (b) a fixed bent crystal spectrometer (from Culhane et al. 1991).

Conventional Bragg spectrometers (Figure 2.4a) scan in wavelength by rotating a flat crystal so the range of incident angles (\( \theta \)) translate to a range of wavelengths (\( \lambda \)) as given by Bragg's law,

\[
n\lambda = 2d \sin \theta,
\]

where \( d \) is a fixed property of the crystal lattice. For BCS each spectrometer crystal is curved with a fixed radius so that a beam of parallel solar X-rays is incident at a range of Bragg angles (Figure 2.4b). The crystals on BCS are concave, rather than convex as shown in the figure, however the effect is the same. The diffracted radiation is then detected by a one dimensional position sensitive proportional counter. Due to the presence of X-ray imaging telescopes on board Yohkoh, the unlikely occurrence of two flares at once and a tight mass and volume budget the
BCS has no multi-grid collimator and therefore views the whole Sun. However, during a flare the intensity increases to well above the quiet corona and combined active region fluxes, hence the emission observed by BCS can be considered to be of flare origin only.

Figure 2.5: Schematic diagram of the BCS (from Culhane et al. 1991).

The four bent Germanium crystals are mounted in two structures as shown in Figure 2.5 with the crystal dispersion axis aligned approximately perpendicular to the solar equator. The wavelength ranges selected are shown in Figure 2.6. The wavelength ranges were selected so they provide diagnostic information in the way of line ratios, yet do not cover too large a range, in order to maintain high sensitivity. The size of the wavelength range is inversely proportional to the sensitivity. The minimum spectral coverage must also account for the variation of source position on the solar disk. Since Bragg angle translates to wavelength
Figure 2.6: Example Bragg crystal spectra of solar flares. (a) Fe XXVI obtained with the HINOTORI spacecraft. (b) Fe XXV and (c) Ca XIX obtained with the bent crystal spectrometer on SMM, and (d) S XV obtained with the flat crystal spectrometer on SMM. The solid lines represent the wavelength ranges selected for the four spectrometers on BCS (from Culhane et al. 1991).

range a change in latitude of flare location on the disk will shift the position of the spectrum in the detector.
2.2 Solar and Heliospheric Observatory

The Solar and Heliospheric Observatory (SoHO; Domingo, Fleck and Poland 1995) is a co-operative mission between the European Space Agency (ESA) and the United States National Aeronautics and Space Administration (NASA). SoHO's scientific objectives are to study the Sun from deep within its core to the outer reaches of its atmosphere and the solar wind. To achieve these goals SoHO carries a payload of 12 instruments that are divided into three categories according to their own principal scientific objectives:

a) Helioseismology

b) Solar Atmosphere Remote Sensing

c) Solar Wind 'in Situ'

The instruments dedicated to solar atmospheric science are listed in Table 2.2 and those that were utilized during this thesis will be described in more detail in the following sub-sections.

SoHO was launched from the Kennedy Space Center on the 3rd December 1995 by an Atlas II-AS rocket. It was inserted into a halo orbit around the Earth-Sun L1 Lagrangian Point; the point on the Earth-Sun line, under the gravitational influence of both, where the spacecraft will remain approximately at rest relative to them. The halo orbit has a period of 6 months and semi-diameters of ≈200,000km within the ecliptic on Earth-Sun direction, ≈650,000km perpendicular to the Earth-Sun line and ≈200,00km perpendicular to the ecliptic (Figure 2.7). The benefits of placing a solar observing spacecraft at the L1 Lagrangian point are threefold. Firstly, free from occultations by the earth, Sun observations are continuous. Secondly, the spacecraft is outside the Earth's magnetosphere away from the energetic particles trapped inside, thus providing accurate 'in situ' measurements of the solar wind. Finally, it provides smooth changes in the Sun-spacecraft velocity throughout its orbit, which are required for accurate heliosiesmology measurements.
Acronym | Instrument | Measurements |
---|---|---|
SUMER | Solar Ultraviolet Measurements of Emitted Radiation | Plasma flow in chromosphere through to corona |
CDS | Coronal Diagnostic Spectrometer | Plasma diagnostics in transition region and corona |
EIT | Extreme Ultra-violet Imaging Telescope | Evolution of chromospheric and coronal structures |
UVCS | Ultra-Violet Coronograph Spectrometer | Plasma diagnostics and velocities in outer-corona |
LASCO | Large Angle and Spectrometric Coronagraph | Coronagraph observations |

Table 2.2: The SOHO Scientific Instruments dedicated to atmospheric science.

The integrated spacecraft is approximately $4.3 \times 2.7 \times 3.7m^3$ and after solar array deployment this extends to a total span of $9.5m$. It is composed of essentially two parts, a service module comprising the service equipment and the propulsion module, and the payload module, onto which the individual scientific instruments are secured (Figure 2.8). The total weight of the spacecraft is $1861kg$ of which $655kg$ is comprised of payload instruments.

The spacecraft attitude was designed to be controlled by gyros, sun-sensors, star trackers and reactor wheels. Unfortunately on 21\textsuperscript{st} December 1998 the final gyro on the spacecraft failed. However, new software developed by the mission team enables the attitude to control and the safe unloading of the momentum wheels to be accomplished using only the star tracker, sun sensors and hydrazine thrusters.
2.2.1 Coronal Diagnostic Spectrometer

The Coronal Diagnostic Spectrometer (CDS, Harrison et al. 1995) is one of two Extreme-Ultraviolet (EUV) imaging spectrometers on board SoHO designed to study the corona and the origins of the solar wind. Critical to the understanding of these phenomena are the determination of plasma temperature, density, flow velocity and abundance at high spatial and temporal resolution. To achieve this CDS views the Sun in the wavelength range 150 → 800Å containing many prime emission lines from ions with formation temperatures in the range $10^4 \rightarrow 10^6 K$.

Unfortunately such a large wavelength range cannot be covered by a single instrument due to the vanishingly small reflective efficiencies of materials below 300Å. Consequently the CDS is a double spectrometer consisting of the Grazing Incidence Spectrometer (GIS) and the Normal Incidence Spectrometer (NIS). This setup has the advantages of a wide wavelength range for the GIS, combined with the
stigmatic imaging capabilities of the NIS. The field of view (FOV) of the telescope is 4 arc-minutes square and can be re-pointed through an angle of ±0.75° in order to cover the entire solar disk.

The optical elements of CDS are shown in Figure 2.9. A grazing incidence Wolter-Schwarztchild type 2 telescope feeds both the NIS and GIS. The spectrometers share common entrance slits where two segments of the focused telescope beam are extracted, one for each spectrometer.

**Grazing Incidence Spectrometer**

The GIS comprises a spherical grating set at grazing incidence with four one-dimensional microchannel plate detectors placed around the Rowland circle. The
resulting spectra are one-dimensional plots of intensity versus wavelength. Thus for GIS operation pin hole slits are used. To build images the slits can be moved in one arc-second increments in the plane perpendicular to the plane of dispersion and the scan mirror incremented in 2 arc-second steps in the plane of dispersion. In the early phases of the mission the CDS science team imposed a rigid limit on the maximum count rate for the GIS that limited its observations to quiet Sun and small non-flaring active regions. Subsequently the GIS has not been utilized in this thesis.

**Normal Incidence Spectrometer**

In the NIS the beam is reflected off two toroidal gratings and focused onto the Viewfinder Detector Subsystem (VDS). The VDS is composed of a microchannel-plate intensifier to detect EUV photons which are then convert to visible radiation which is collected by a CCD. By imposing a small out-of-plane tilt on the gratings two spectral bands are simultaneously displayed on the same CCD but displaced
in the direction perpendicular to the spectral dispersion so the images appear stacked above one-another. The image in the upper half of the detector covers the wavelength range 310 → 380Å at 0.08Å resolution (NIS1) and the lower image the wavelength range 520 → 630Å at 0.14Å resolution (NIS2). The image formed on the detector is essentially that of the entrance silt. To build larger images the scan mirror is scanned through integer multiples of 2 arc-seconds.

In practice telemetry limits are the defining factors governing temporal resolution, therefore the entire wavelength range is rarely used. Instead spectral windows of the order of 1.2 → 5.6Å are selected around emission lines of interest which are subsequently transmitted to the ground. Other factors contributing to the temporal resolution by means of the amount of data that needs to be transmitted are spatial size of observing region, spatial resolution, number of selected lines, spectral window size and exposure times. When designing a CDS observing programme it is sometimes necessary to limit one or more of the above in order to achieve greater temporal resolution.

### 2.2.2 Extreme-Ultraviolet Imaging Telescope

The Extreme-Ultraviolet Imaging Telescope (EIT; Delaboudiniére et al. 1995) was designed to provide full Sun images out to a distance of 1.5 \( R_\odot \) at a range of temperatures from chromospheric to coronal. The primary scientific objectives of EIT are to study the dynamics of coronal structures over a range of sizes, times and temperatures in order to further understand the dominant mechanisms that generate and heat the corona and accelerate the solar wind. Full disk EIT images also allows the connection to be made between low coronal structures and high coronal structures, seen with LASCO and UVCS the coronagraphs on board SoHO (Table 2.2).

To this end the EIT views the Sun in four narrow wavelength ranges. Each wavelength range is centred around a prime emission line of a particular ionic
species. These ions and wavelength ranges are listed in Table 2.3 along with the peak temperature of formation for that ion and the solar features that give rise to emission at that temperature.

Imaging the Sun in selected narrow wavelength ranges is accomplished by employing multilayer normal incidence EUV optics (Spiller, 1994). By dividing the telescope mirrors into quadrants it is possible to isolate different temperature ranges by 'tuning' the coating on each quadrant to the desired wavelength range (Table 2.3). The various multilayers are fabricated from alternating layers of Molybdenum and Silicon. Wavelength ‘selection’ is achieved by interference effects arising within the coating.

Figure 2.10 shows a schematic diagram of the EIT. The rotating mask allows only one quadrant of the telescope, coated in a single multi-layer, to be illuminated at any one time. Since each quadrant of the optics forms an image on the same CCD the image registration is achieved automatically. The CCD is a 1024×1024 pixel array giving a spatial resolution of 5.2 arc-seconds (pixel size 2.6 arc-seconds). The filter wheel near the focal plane has two filters to block long wavelength solar emission and an open position.

<table>
<thead>
<tr>
<th>Wavelength (Å)</th>
<th>Ion</th>
<th>Peak Temp. (K)</th>
<th>Atmospheric Region</th>
</tr>
</thead>
<tbody>
<tr>
<td>304</td>
<td>He II</td>
<td>$8.0 \times 10^4$</td>
<td>Chromospheric network; coronal holes</td>
</tr>
<tr>
<td>171</td>
<td>Fe IX/X</td>
<td>$1.3 \times 10^6$</td>
<td>Corona/transition region boundary</td>
</tr>
<tr>
<td>195</td>
<td>Fe XII</td>
<td>$1.6 \times 10^6$</td>
<td>Quiet corona</td>
</tr>
<tr>
<td>284</td>
<td>Fe XV</td>
<td>$2.0 \times 10^6$</td>
<td>Active regions</td>
</tr>
</tbody>
</table>

Table 2.3: The EIT Wavelength Bandpasses (from Delaboudinière et al. 1995).
2.2.3 The Solar Oscillations Investigation - Michelson Doppler Imager

The Solar Oscillations Investigation - Michelson Doppler Imager (MDI: Scherrer, et al. 1995) is primarily a helioseimological instrument designed to probe the solar interior by detection of photospheric manifestations of global solar oscillations. To achieve this MDI images the whole Sun in five 94mÅ bandpasses centred around the Ni I 6768Å spectral line. MDI is thus capable of calculating line of sight velocity, line of sight magnetic field strengths, continuum intensity around Ni I 6768 and Ni I 6768 line intensity.

The line of sight magnetograms are constructed by measuring the Doppler shift separately in right and left circularly polarized light. The difference between these two is a measure of the Zeeman splitting (Chapter 1.3.3) of the Ni I line and is roughly proportional to the magnetic flux density, i.e. the line of sight component of the magnetic field averaged over the resolution element. The error in the flux density for each pixel is approximately ±20 Gauss.
MDI can operate in two modes; high-resolution and normal resolution. In normal resolution the telescope views the full Sun (34 arc-minute square field of view) at a resolution of 4 arc-seconds. In high resolution mode the image is magnified by a factor of 3.2 to provide a resolution of 1.25 arc-seconds over a field of view of 11 arc-minutes square. The position of the high resolution field is centred at 160 arc-seconds north of the equator to allow observations of active regions and quiet Sun.
2.3 Transition Region and Coronal Explorer

The Transition Region and Coronal Explorer (TRACE; Handy et al. 1999; Schrijver et al. 1999) is a normal incidence imaging EUV telescope that views the Sun through three narrow-band EUV filters and several UV filters. TRACE views the solar photosphere, transition region and corona with unprecedented spatial resolution and temporal continuity.

TRACE, a NASA small explorer mission (SMEX), was launched on 2nd April 1998 from Vandenberg Air Force Base on a Pegasus XL launch vehicle into a Sun-synchronous polar orbit, allowing uninterrupted viewing of the Sun for 9 months of the year. During the remaining 3 months the spacecraft is in 'eclipse season' where part of the orbit is occulted by the earth.

The TRACE observatory is a single instrument spacecraft; a normal incidence imaging EUV telescope (Figure 2.12). The telescope is a Cassegrain design with a 30cm aperture and a $8.5 \times 8.5$ arc-minute field of view. TRACE uses normal incidence multi-layer optics to view the Sun in selected narrow wavelength ranges.

![Figure 2.12: Schematic Diagram of the TRACE instrument (from Handy et al. 1999).](image)
Table 2.4: The TRACE temperature response.

<table>
<thead>
<tr>
<th>Wavelength (Å)</th>
<th>Emission</th>
<th>Temperature (K)</th>
</tr>
</thead>
<tbody>
<tr>
<td>171</td>
<td>Fe IX/X</td>
<td>0.2 → 2.0 × 10^6</td>
</tr>
<tr>
<td>195</td>
<td>Fe XII</td>
<td>0.5 → 2.0 × 10^6</td>
</tr>
<tr>
<td>284</td>
<td>Fe XV</td>
<td>1.3 → 4.0 × 10^6</td>
</tr>
<tr>
<td>1216</td>
<td>H I Lyα</td>
<td>1.0 → 3.0 × 10^4</td>
</tr>
<tr>
<td>1550</td>
<td>C IV</td>
<td>0.6 → 2.5 × 10^5</td>
</tr>
<tr>
<td>1600</td>
<td>UV cont, C I, Fe II</td>
<td>0.4 → 1.0 × 10^4</td>
</tr>
<tr>
<td>1700</td>
<td>Continuum</td>
<td>0.4 → 1.0 × 10^4</td>
</tr>
<tr>
<td>5000</td>
<td>White Light</td>
<td>4.0 → 6.4 × 10^3</td>
</tr>
</tbody>
</table>

(Table 2.4). The TRACE optical path is composed of a set of entrance filters, multilayer coated primary and secondary mirrors, two filter wheels behind the primary mirror and a lumogen coated CCD. The telescope is divided into quadrants with each quadrant optimized for different wavelength bands. One quadrant each for 171Å, 195Å and 284Å and the other quadrant for the UV ranges. The angular resolution of the instrument, 1 arc-second, is defined by the pixel size of the CCD (21μm, 0.5 arc-seconds).

Figure 2.13 shows the signal generated in the CCD through the various TRACE EUV filters as a function of temperature. The peak in the 195Å filter response function at ≈ 20 × 10^6 K allows the detection of hot plasma. Plasma emission at ≈ 20 × 10^6 K with the TRACE 195Å filter is dominated by Fe XXIV line emission (Feldman et al. 1999; Warren et al. 1999). This is only visible during flares when the amount of ≈ 20 × 10^6 K plasma becomes significant.
Figure 2.13: Signal at the CCD from each of the TRACE EUV pass bands as a function of source temperature. The dotted, dashed and solid line refer to different primary entrance filters. The emission measure is $10^{14}\text{cm}^{-3}$ (from Handy et al. 1999).
2.4 Burst and Transient Source Experiment

The Burst and Transient Source Experiment (BATSE; Fishman 1989) is one of a suite of instruments on board the Compton Gamma Ray Observatory (CGRO) designed for the study of γ-ray bursts. BATSE is a whole sky HXR flux monitor operating in the range 10keV→100MeV. Because BATSE views the whole sky HXR from solar flares are frequently observed.

BATSE consists of eight detector modules that are mounted on the faces of an octahedron formed by the three major axes of the CGRO satellite. Each detector module (Figure 2.14) comprises two NaI(Tl) scintillation detectors: a large area detector (LAD) and a spectroscopy detector (SD).

The LADs, optimized for sensitivity and directional response, have a sensitive area of 2025cm² and operate in the energy range 25keV→1.9MeV with an energy resolution of ≈ 27% at 88keV and increases with energy. The scintillation light
from the NaI crystal is detected by three five-inch diameter photo-multiplier tubes (PMT). The signals from the three tubes are then summed. The detector is shielded on the front by a 7mm thick plastic anti-coincidence shield to reduce the charged particle background and at the rear by a thin lead and tin shield to reduce scattered radiation. The angular response of the detectors below 300keV is approximately a cosine function, thus allowing the combination of fluxes in different detectors to be used to locate the HXR source location. The temporal resolution can be as low as 64ms.

The SDs, optimized for energy range and resolution, have a sensitive area of 127cm$^2$ and operate in the range 10keV→100MeV with a maximum of 256 channels. Scintillation light is collected by a single PMT. The shielding is similar to that for the LAD with the addition of a beryllium window on the front face to provide high efficiency down to 10keV.
Chapter 3

Data Methods

In this chapter I describe the methods and techniques of data analysis for the principle instruments used in this thesis. The analysis of data from all the solar instrumentation used in this thesis is aided by a software package called SolarSoft. The programs therein perform many of the basic tasks required to extract data in a form that can be easily analysed from the raw data telemetered to the ground from the respective satellite. This chapter deals in principal with examining the methods and techniques required to perform this extraction and comments on the resulting reliability and accuracy of the data and of the techniques themselves. A physical description of all the instruments was given in Chapter 2.

3.1 Yohkoh SXT

Data from a particular orbit from the the SXT instrument, and indeed from every Yohkoh instrument, are stored as binary files at the Institute of Space and Astronautical Sciences (ISAS) and at several mirror sites around the world including MSSL. Processing raw SXT data files into images involves the following steps some of which are self explanatory others are described in more detail in the following paragraphs.

1) subtraction of the dark current, flat field and pinhole leak,
2) apply the vignette correction,

3) flag the saturated pixels and remove cosmic ray spikes,

4) apply roll correction, align the images and exposure normalise.

In 1992 the white light filters at the front of the telescope before the mirror assembly, were struck a micro-meteorite. This produced a pinhole in the filter which allows undesired light to enter. Because the CCD is also sensitive to white light this visible light contamination shows up on X-ray images. This effect is compensated for by taking an image at the start of every orbit whilst the satellite views the Sun through the Earth’s atmosphere. Earth’s atmosphere completely attenuates the X-ray flux leaving only a white light image of the pin-hole leak plus dark current. An image of this type plus the pre-flight data of the CCD flat field is then subtracted from each X-ray exposure. Vignetting is the reduction in effective area with increasing off axis angle. This is due primarily to the reduced collecting area of the mirror assembly and secondly due to the obscuration of the optical axis by the telescope construction axis and baffles (Fuller, Lemen & Acton 1994). The change in effective area was evaluated preflight.

Saturated pixels cannot be used for plasma diagnostics and must therefore be flagged in order to prevent spurious temperature measurements. The removal of cosmic rays is accomplished by looking for pixels that are considerable brighter than their surrounding eight neighbours and also those in the previous and subsequent frames. Since cosmic ray impacts are random, the chances of two striking the same pixel in succession is remote. Hence bright solar features of the scale of a single pixel that last longer than 4 seconds (the time to make three exposures) are not mistaken for cosmic rays.

After these corrections have been applied the SXT images are ready for examination. The intensity scale in the images is measured in units of Data Numbers (DN), a measure of the number of photons detected in each pixel (see Section 3.1.1 for definition).
Plasma diagnostics are achieved through the use of different analysis filters. These filters are of different thickness and elemental composition and therefore have different transmission functions, \( t_i(\lambda) \), as a function of wavelength, where \( i \) denotes a specific filter. The transmission functions of the filters are shown in figure 3.1. Figure 3.1 indicates that there are essentially three thin filters (Al 1265 Å, Al/Mg/Mn (Dagwood filter), and Mg 2.52 \( \mu \)m) and two thick filters (Al 11.6 \( \mu \)m and Be 119 \( \mu \)m) and that the thick filters have a larger response at higher temperatures. Therefore merely by comparing images in a thin filter (e.g. Al 1265 Å) and a thick filter (e.g. Be119\( \mu \)m) it is possible to determine which features are the hottest. However to obtain a quantitative measure of the temperatures and emission measures it is necessary to take ratios of images taken through different analysis filters. The technique behind the derivation is as follows.

3.1.1 SXT Temperature Diagnostics

The overall sensitivity of the instrument \((n_i(\lambda))\) can be expressed as a product of the telescope aperture \((A)\), the mirror reflectivities \((r(\lambda))\), the filter transmission functions \((t_i(\lambda))\) and the CCD efficiency \((e(\lambda))\).

\[
n_i(\lambda) = A \times r(\lambda) \times t_i(\lambda) \times e(\lambda) \tag{3.1}
\]

these functions were calculated pre launch and therefore the value of \( n_i(\lambda) \) for each analysis filter is known.

The total energy input onto a single CCD pixel \((E_i)\) can be written,

\[
E_i = \frac{\sigma}{4\pi D^2} \int_\lambda \int_{0}^\infty P[\lambda, T(l)] n_i(\lambda) n_e(l)^2 \, d\lambda \, dl \tag{3.2}
\]

where \( \sigma \) is the area on the Sun covered by one pixel, \( D \) is the Earth-Sun distance, \( l \) the line of sight depth, \( n_e(l) \) the electron density and \( T(l) \) the temperature of the plasma. The function \( P[\lambda, T(l)] \) represents the emissivity of the plasma as a function of wavelength and temperature. This function can be calculated theoretically from the synthetic codes (Mewe et al. 1985a, 1986).
Figure 3.1: Effective area of the SXT; (a) no analysis filter, (b) 1265Å Al filter, (c) composite filter containing 2930Å Al, 2670Å Mg, 562Å Mn and 190Å C termed Dagwood filter, (d) 2.52µm Mg, (e) 11.6µm Al and (f) 119µm Be. (b)-(f) compare the no filter case (dashed line) with the filtered case. (a) is plotted on a logarithmic scale to better illustrate the regions of small effective area (from Tsuneta et al. 1991).
Figure 3.2: (a) The total SXT signal as a function of temperature for the open filter and each of the analysis filters (assuming an emission measure $10^{44} \text{ cm}^{-3}$). (b) The ratios of the SXT response functions. The effective observation time has been multiplied for some filters where indicated (from Tsuneta et al. 1991).

With no a priori knowledge of the line of sight dependencies we must assume the plasma is isothermal, therefore $T(l) \equiv T$. Combining the two wavelength dependent functions gives,

$$F_i(T) = \int P[\lambda, T] n_i(\lambda) \, d\lambda. \quad (3.3)$$

The Volume emission measure, $VEM$, is given by,

$$VEM = \sigma \int_{0}^{\infty} n_e(l) \, dl. \quad (3.4)$$

The emission measure (EM) is simply $EM = VEM/\sigma$. We can therefore rewrite equation 3.2 using equations 3.3 and 3.4 as,

$$E_i = \frac{1}{4\pi D^2} \times F_i(T) \times VEM. \quad (3.5)$$

Photons striking the CCD are converted into electron-hole pairs. The number of electron-hole pairs, $N_i$, is given by,

$$N_i = \frac{E_i}{k_e}, \quad (3.6)$$
where $k_e$ is a conversion factor and is approximately $3.6eV$. A Data Number (DN) is defined as,

$$DN_i = \frac{N_i \times \delta t}{100} \quad (3.7)$$

where $\delta t$ is the exposure time of the image. Combining equations 3.5, 3.6 and 3.7 and setting,

$$f_i(T) = \frac{1}{400\pi D^2} \times F_i(T), \quad (3.8)$$

we obtain the expression describing the intensity in an SXT image as a function of temperature and emission measure,

$$DN_i = f_i(T) \times VEM \times \delta t. \quad (3.9)$$

The signal dependence on temperature for each SXT filter is shown in figure 3.2a. Therefore in order to obtain the temperature we take ratios of intensity, $R_{ij}$, from two different SXT filters, labeled $i$ and $j$, to give,

$$R_{ij} \equiv \frac{DN_i \delta t_j}{DN_j \delta t_i} = \frac{f_i(T)}{f_j(T)}. \quad (3.10)$$

This ratio is only a function of temperature. Hence the temperature of the plasma can be calculated. The ratio dependence on temperature for several SXT filter pairs is shown in figure 3.2b.

In order to then compute the emission measure the calculated temperature is used in equation 3.9 to extract the $VEM$.

### 3.1.2 Error Sources in SXT temperature Diagnostics

There are several sources of error in SXT plasma diagnostics, which result both from the telescope design and from the method with which they are calculated. The principal error sources are described briefly below.

1) **Statistical errors in the count rate**: Errors in $DN_i$ due to Poisson noise in the data propagate through the above derivation to give an associated error in $T$ and $VEM$.
2) **Scattered light:** Light scattered of mirror micro-roughness and telescope elements results in a redistribution of counts in the image.

3) **Spacecraft jitter:** The resulting slightly misaligned images, at sub-pixel level can result in erroneous ratio measurements.

4) **Source variations:** The source temperature may change during the finite time delay (2 seconds in flare mode) between taking images in different filters.

The complete derivation of the temperature and emission measure errors resulting from statistical noise can be found in Hara (1992a) and Kano & Tsuneta (1995). The resulting formulae express the standard deviations of temperature, \( \sigma_T \) and emission measure, \( \sigma_{EM} \), in terms of measurable quantities,

\[
\frac{\sigma_T}{T} = \left| \frac{d \ln R_{ij}(T)}{d \ln T} \right|^{-1} \sqrt{\frac{3}{DN_i} + \frac{3}{DN_j}},
\]

\[
\frac{\sigma_{EM}}{EM} = \left| \frac{d \ln R_{ij}(T)}{d \ln T} \right|^{-1} \sqrt{\left[ \frac{d \ln f_i(T)}{d \ln T} \right]^2 \frac{3}{DN_i} + \left[ \frac{d \ln f_j(T)}{d \ln T} \right]^2 \frac{3}{DN_j}},
\]

It is these statistical errors that are quoted throughout this thesis on measurements made with SXT.

The other sources of errors in SXT temperature maps, listed above, have proved more difficult to quantify. Spacecraft jitter and image misalignment were shown to produce spurious temperatures by Siarkowski et al. (1996). They showed that a misalignment of as little as 0.12 arc seconds can create temperature deviations of as much as 1.5MK. However the spacecraft jitter should in theory be random. Therefore if a hot or cold source is present in several images, we can discount spacecraft jitter. Compensation for source variations and indeed improving the counting statistics can be achieved by summing together several SXT images from each filter and then performing ratios. Temperature measurements are more reliable during periods in which the source variation is small.

Scattering is caused by micro-roughness of the grazing incidence X-ray mirror and is described by the power-law wing of the point spread function (PSF) that
was calculated by Hara et al. (1994). Therefore in principle the scattered light contribution to an image can be removed to a first approximation by deconvolving the image with the known PSF. However, a quantification of the effects of scattered light performed by Foley (1998) showed that the level of scattered light was not as straightforward. It was not constant in azimuthal angle and highly vignetted. The levels of scattered $I_{stat}$ could be represented by,

$$I_{stat} = \alpha(\theta) \eta(\lambda r)^{-2}$$  \hfill (3.13)

where $\alpha(\theta)$ is a function describing the azimuthal variation in scattered light, $\eta$ is the correction for vignetting, $r$ if the radial distance from the source and $\lambda$ the wavelength of the scattering X-rays. This function can be calculated for individual flares only if a saturated full frame image is available.

### 3.2 Yohkoh HXT

Due to the vanishingly small grazing incidence reflectivities of current materials at HXR wavelengths it is not possible to build a simple telescope similar to SXT. HXT therefore is a Fourier imaging telescope. The two sets of grids modulate incoming parallel radiation, with a period, $\lambda$, equal to $\lambda = p/D$, where $p$ is the wire spacing (pitch spacing) and $D$ the grid spacing. By setting pitch spacing to twice the wire thickness gives a transmission function that is triangular (Figure 3.3).

Let $(x, y)$ be the sky plane orthogonal to the collimator axis, then the transmission function of each cosine sub-collimator, $M_C(kr)$, is given by the Fourier expansion in the natural polar co-ordinate system $(r, \theta)$,

$$M_C(kr) = \frac{2}{\pi^2} \left\{ \frac{\pi^2}{8} + \cos(kr) + \frac{1}{9} \cos(3kr) + \frac{1}{25} \cos(5kr) + \ldots \right\}, \quad (3.14)$$

$$r = x \cos \theta + y \sin \theta. \quad (3.15)$$

Each cosine sub-collimator has a sine collimator partner with the same pitch and position angle but whose phase has been shifted by $\pi/2$, therefore the transmission
of this subcollimator can be written as,

\[ M_S(kr) = M_C(kr - \frac{\pi}{2}) \]  

(3.16)

where the subscript \( S \) denotes a sine sub-collimator and \( C \) a cosine. A cosine sub-collimator can be changed to a sine by shifting its position perpendicular to the wires by \( \pi/4 \).

If \( B(x, y) \) the the source distribution in the sky plane, then the count rate data from the sine-cosine sub-collimators is given by,

\[ b_C(k, \theta) = A \int_{F_OV} B(x, y) M_C(kr) \, dx \, dy, \]  

(3.17)

\[ b_S(k, \theta) = A \int_{F_OV} B(x, y) M_S(kr) \, dx \, dy, \]  

(3.18)

where the integral is over the field of view of the collimator and \( A \) the the effective area. \( b_C(k, \theta) \) is a cosine type spatial Fourier component of \( B(x, y) \) corresponding to wavenumber \( k \) and the angle \( \theta \), plus DC component. Thus by observing a distribution \( B(x, y) \) through a pair of sine-cosine collimators, having the same
wavenumber and position angle, we uniquely determine a spatial Fourier component of \( B(x, y) \) corresponding to \( k \) and \( \theta \). Thus by adopting the \((U, V)\) co-ordinates for the Fourier plane which is conjugate to \((x, y)\), it can be said that a pair of cosine and sine collimators measures a complex Fourier component \( b(U, V) = b_C + ib_S \) where, \( U = k \cos \theta \) and \( V = k \sin \theta \). Consequently if \( k \) and \( \theta \) are allowed to vary as two independent variables, or if we have many sine-cosine collimator pairs with different values of \( k \) and \( \theta \) which allows us to sample many Fourier components we can reconstruct \( B(x, y) \) through an inverse Fourier transform.

### 3.2.1 Image Reconstruction Theory

As was mentioned in Chapter 2.1.2, because the transmission functions are not purely trigonometric and there are insufficient Fourier component pairs, direct inversion produces significant noise and spurious signals in the recovered image. One image reconstruction process used by members of the community and in this thesis is the Maximum Entropy Method (MEM). MEM produces the most uniform distribution of intensity that best represents the data. The principal behind the most uniform distribution is to try to ensure that all the structure in the images is real and not due to noise.

The entropy, \( S \), in the image is given by,

\[
S = -\sum_{i,j=1}^{N} \left( \frac{B_{ij}}{B_0} \right) \ln \left( \frac{B_{ij}}{B_0} \right),
\]

where \( B_{ij} \) is the MEM image of the HXR source and \( B_0 \) the average HXR photon count rate per pixel, the sum is over all image pixels. The solution that best satisfies equation 3.19 is a grey map, i.e. a uniform counts distribution. Therefore in order to constrain the solution to one that represents the data, a measure of the goodness of the fit is required. Due to the good statistical quality of the data the \( \chi^2 \) statistic is used,

\[
\chi^2 = \frac{1}{64} \sum_{K=1}^{64} \frac{(b'_K - b_K)^2}{\sigma_K^2} \approx 1
\]

\( \text{where } b'_K \text{ and } b_K \text{ are the measured and estimated photon count rates, respectively, and } \sigma_K \text{ is the standard deviation of the count rate.} \)
where $\sigma_K$ the the standard deviation of the original data set and the sum is over all sub-collimators. $b_K$ are the observations and $b'_K$ the estimates of the Fourier components from the current solution $B_{ij}$ and are given by the generalized form of equations 3.17 and 3.18,

$$b_K = \sum_{i,j} B_{ij} M_K(k, r)$$

$$P_{ij,K} = A M_K(k, r)$$

$$\Rightarrow b_K = \sum_{x,y} P_{ij,K} B_{ij} + n_K$$

where $P_{ij,K}$ is the response matrix for the $K^{th}$ sub-collimator and $n_K$ represents uncertainties in $b_K$ due to Poisson noise and systematic errors.

The process of reconstructing an image then becomes equivalent to maximising, $\bar{S}$, which is given by,

$$\bar{S} = S - \frac{\lambda}{2} \chi^2,$$

where $\lambda$ is a parameter that weights observations ($\chi^2$) against the entropy. Increasing the value of $\lambda$ adds more weight to the observations, hence the $\chi^2$ value decreases and synthesised images contain more structure.

Following Sakao (1994) taking partial derivatives of equation 3.24 with respect to $B_{ij}$ to be zero, we obtain,

$$B_{ij} = B^0 \exp \left[ -\frac{S}{N^2} + \lambda B^0 \sum_{K=1}^{64} \left\{ \frac{1}{\sigma_K^2} \left( b_K - \sum_{i,j=1}^{N} P_{ij,K} B_{ij} \right) P_{ij,K} \right\} \right]$$

the iterative solution of which is,

$$B_{ij}^{(l+1)} = (1-\gamma) B_{ij}^{(l)} + \gamma B^0 \exp \left[ -\frac{S^{(l)}}{N^2} + \lambda B^0 \sum_{K=1}^{64} \left\{ \frac{1}{\sigma_K^2} \left( b_K - \sum_{i,j=1}^{N} P_{ij,K} B_{ij}^{(l)} \right) P_{ij,K} \right\} \right]$$

where $\gamma$ is a weighting function termed iteration gain. Large values of $\gamma$ make the iteration more stable but the convergence slower. Usually $\gamma \approx 0.02 \rightarrow 0.1$ gives good convergence within a reasonable number of iterations.

From a starting point, $l = 0$, of a flat intensity distribution, or gray map, and successively iterating equation 3.26 with increasing values of $l$ we obtain the
maximum entropy image, $B_{ij}$, at a given value of $\lambda$, at convergence. The condition for convergence is given by,

$$\frac{\sqrt{\sum_{i,j=1}^{N}(B_{ij}^{l+1} - B_{ij}^{l})}}{\sqrt{\sum_{i,j=1}^{N}(B_{ij}^{l})}} \approx 0.01. \quad (3.27)$$

Once the solution has converged with start value of $\lambda$ ($\approx 0.2$), $\lambda$ is incremented by a small amount ($\approx 0.1$) and the iteration restarted with the new value of $\lambda$. The solution of the previous convergence is used as a starting point in the new iteration sequence. The value of $\lambda$ is successively increased until it reaches the user defined value, which is usually set to $\approx 50$.

### 3.2.2 Image Reconstruction Practice

For the whole image synthesis process that involves obtaining HXR images from raw data the reconstruction described above is the final step a larger process. The overall process is as follows,

1) Background subtract the data

2) calculate the image synthesis field of view

3) run the MEM reconstruction.

Background subtraction of the data is performed by selecting the count rate from a lightcurve, preferably just before the flare or just after, and subtracting this count rate from the data.

Before attempting a MEM image reconstruction we must first determine the location of the flare on the Sun and hence determine the synthesis aperture. The synthesis aperture of HXT is 126 arc-seconds and is determined by the grid and pitch spacings. The image synthesis is then performed assuming that no counts come from outside the synthesis aperture window. There are essentially two ways to determine the flare location, firstly to retrieve the location from another imaging
telescope, e.g. SXT or an Hα instrument, or secondly to construct a dirty map from the HXT data. A dirty map is a direct two dimensional inverse Fourier transform of the HXT data, and is termed dirty due to the spurious structure and repetitive patterns that result from the incomplete UV plane coverage and fundamental period of the Fourier transform respectively (Sakao 1994). Although this dirty map contains this 'unreal' structure the brightest peak is likely to be the flare itself and from this we can determine the flare location.

In this thesis the flare location was determined using the pointing information from SXT. This method is computationally much faster than constructing dirty maps, and gives ≈ 1 arc-second co-alignment between the resulting synthesised HXT images with SXT images (Masuda 1994).

Before running the MEM procedure we must determine the interval period over the HXR burst from which we wish to construct an image. Approximately 200 counts is considered a minimum to generate a reasonably good image. More counts give a better quality image, with a corresponding loss of temporal resolution.

### 3.3 Yohkoh BCS

The BCS operates in specific wavelength ranges that cover the principal emission lines of He-like S XV, Ca xix and Fe xxv and H-like Fe xxvi. For the He-like ions these wavelength ranges cover the principal resonance line $1s2p \rightarrow 1s^2$ and the nearby dielectronic satellite lines. Hence the spectra can be used for temperature diagnostics of the emitting plasma (Chapter 1.3.2).

When describing the theory of dielectronic recombination and temperature sensitivity in Chapter 1.3.2 I used the specific example of the $1s2p + e \rightarrow 1s2s2p \rightarrow 1s^22s$ transition where the spectator electron is in a 2s level. However, the spectator electron can be at any energy level depending on the collisional energy of the initial exciting electron. Hence conventionally the excited state is denoted $1s2pnl$ where $n$, the principle quantum number ranges from 0 → ∞. The presence of a
spectator electron shields the $2p$ electron from the positive charge of nucleus, hence the energy of the resonance transition is lowered. The lower the value of $n$, the greater the shielding effect. Satellite transitions with $n = 2$ are well separated from the resonance line, satellites with $n = 3$ can be resolved from the resonance line, however, satellites with $n \geq 4$ cannot (Figure 3.4). The number of individual satellite lines increases drastically with $n$. For $n = 2$ there are 22 individual transitions, for $n = 3$ there are 148. These individual satellite transitions are very close in energy, much less than the thermal width of the line and cannot be resolved (Figure 3.4).

Figure 3.4: The relative contributions of the satellite lines with different $n$ values (from Bely-Dubau, Gabriel & Volonté 1979).

Figure 3.5 shows a sample BCS Ca XIX spectrum from the decay phase of an M1.7 solar flare, the resonance line, $n = 3$ satellites are labeled. Comparison of this spectrum with the theoretical spectra shown in figure 3.4 illustrates that simply
fitting two Gaussians to the resonance line and the $n = 3$ satellites would not be accurate as it would overestimate the strength of the resonance line by neglecting the $n \geq 4$ satellite lines.

Therefore to extract the intensities of the main resonance line and the satellite lines from a raw spectrum, an iterative fitting procedure is used. This procedure involves the repeated synthesis of theoretical spectra, including contributions of satellites with $n \leq 10$, covering the spectral region to be fitted. The starting theoretical spectra, composed of Gaussian line profiles, is produced from intelligent guesses of the parameters to be fitted, together with independently determined atomic data (Bely-Dubau et al. 1979a,b, 1982a,b; Lemen et al. 1984; Sampson & Clark 1979; Vainshtein & Safronova 1978, 1985). The generated theoretical spectrum is first convolved with the instrument response function and then compared
with the raw data. The differences are then carried forward to the next iteration as modifications to the fit parameters. This process continues until the minimization of $\chi^2$ between the observed spectrum and the fitted theoretical spectrum. The computer programme that performs the above procedure was written by D. Zarro and J. Lemen in the IDL language and is incorporated in the SolarSoft data analysis package.

The BCS instrument response composes of both a Gaussian and a Lorentzian component. The Gaussian component describes the position sensitivity of the 1D proportional counters and the Lorentzian describes the crystal rocking curve. Bragg’s law states that a photon incident on the crystal, with the correct wavelength at the required angle will experience specular reflection. However imperfections in the crystal lattice implies that the reflection occurs over a small range of angles. This is known as the crystal rocking curve which was measured before launch. The FWHM of the Gaussian component of the CaXIX channel is $2.9 \times 10^{-4}$Å and the rocking curve FWHM is $3.9 \times 10^{-4}$Å. In comparison the thermal width of the line at $15 MK$ is $1.4 \times 10^{-3}$Å.

One of the parameters of the fitted theoretical spectrum is the emission measure. This is determined through the application of Equation 1.15 with independent knowledge of the $G(T)$ function for the ionic species. $G(T)$ functions for SxV, CaXIX and FeXXV are shown in Figure 3.6 and are taken from Arnaud and Rothenflug (1985), for CaXIX and SxV, and Arnaud & Raymond (1992). for FeXXV.

The errors in BCS data are governed by counting statistics. However the errors on the derived plasma parameters incorporate these and several extra uncertainties, namely, errors in the atomic data, abundances and the fit. Abundances and atomic data are required to calculate the line intensities and derive the plasma temperature (Chapters 1.3.1 and 1.3.2 respectively). Contributions from all these sources of errors are included in the error values given with BCS data throughout this thesis.

Line narrowing in the BCS data is an effect which artificially narrows and
distorts the line shapes in the BCS data, this effect is most prominent at high count rates however it is always in effect (Trow, Bento & Smith 1994). Line narrowing is caused by the low mobility of the ions in the proportional counters. When a photon enters the proportional counters it ionizes the gas, producing an electron cloud that is registered as a pulse on the anode. At low count rates the resulting positive ion cloud has time to disperse and be quenched. At high count rates the ion cloud can remain and subsequent electron clouds produced by photon entry are attracted toward it. The largest ion clouds form at regions of high count rate, i.e. the position of the main resonance line. Hence the main resonance line appears much thinner and neighbouring lines are skewed in that direction.

Figure 3.6: $G(T)$ functions for S XV, Ca XIX and Fe XXV
CHAPTER 3. DATA METHODS

3.4 SoHO CDS-NIS

CDS data is provided in the form of a FITS file, composing the exposure data in binary format, the exposure and spacecraft state information in ASCII format. In order to reach a point at which we can fit emission lines the data must first be calibrated. This involves several steps, as outlined below.

1) Removal of the VDS flat field, burn-in, dark current and cosmic rays.
2) Conversion to physical units.
3) Corrections to remove the tilt and slant of the image on the detector.

The design of the VDS (Chapter 2.2.1) is such that bright emission lines will always fall on the same location at the detector. Those pixels that correspond to the position of the bright lines will receive the largest dose over time, consequently the loss of sensitivity will be greatest in those areas. This effect is known as the burn-in. By using the 90 arc-second slit the burn-in can be directly observed, quantified and subsequently removed from thin slit observations.

The slant is a small angular offset of the wavelength dispersed slit image on the CCD and results from the slight angular offset of the toroidal gratings. By exposing and reading out the entire CCD this slant has been measured and can be removed. The tilt is a small angular tilt of the spectral lines with respect to the dispersion direction (Figure 3.7). This tilt is also a function of wavelength (or equivalently spectral bin) and can be evaluated from a full CCD exposure.

The NIS is a stigmatic imaging telescope, meaning that in order to construct a two dimensional image of, for example an active region, the entrance slit must be placed at one edge and then successively stepped across the region, in steps the same size as the width of the slit, taking an exposure at each location, until the 2D image is complete. Each slit exposure produces a spectrum of a 1D slice of the region. Each pixel in the final 2D image therefore has spectral information across the wavelength range of NIS at the instrument resolution. In order to produce an
image formed by a given ion, for example O\textsubscript{v}, it is necessary to fit a Gaussian to the observed emission line at each pixel. The intensity of the pixel is then the integral under the Gaussian. For the strong isolated emission lines the procedure is straightforward, however, for blended lines it is necessary to fit more than one Gaussian.
Chapter 4

Multi-Wavelength Study of a Homologous Flare Series

4.1 Introduction

Homologous flares are an interesting component of the solar flare phenomenon. By studying homologous flares we have the potential to gain insight into the flare trigger mechanism and the time-scales involved for the build up of flare energy in the coronal magnetic field (Chapter 1.4.5). Under the assumption that homologous flares result from the continual stressing of an arcade we can study the series of flares to check for similar circumstances that occur before for each one, for example critical shear angle or velocity. If common circumstance can be found then it is possible that we have identified the flare trigger. The time interval between the flares is also a useful quantity and determines the maximum time for the corona to store the energy that was released during the flare. Both the trigger mechanism and time-scale for energy storage are critical quantities in understanding the phenomena of solar flares.

In this chapter I have studied a series of flares that occurred in an active region on the west limb of the Sun on 17th September 1997. Two of these flares were
observed to be homologous. I have used multi-wavelength data to perform a de­
tailed study of the flares in this series in order to extend the understanding of the
conditions and mechanisms that lead to the formation of homologous flares.

In the following sections I briefly describe the observations and present a de­
tailed study of the individual flares in the series. A schematic diagram of the
evolution of the active region inferred from the observations is then presented il­
ustrating the nature of the observed homology. Finally I discuss how these results
impact on the understanding of flare homology.

4.2 Observations

The flares studied in this chapter all originated in active region NOAA 8084 on the
west solar limb on 17th September 1997. This active region was observed by both
Yohkoh and SoHO at all times from the beginning of the first flare to the end of
the last.

This work is centred around images taken with SXT, HXT, EIT and CDS. These
instruments are described in detail in Chapter 2. During the flares SXT and HXT
operated in flare mode, taking full resolution images (2.5 arc-seconds) in all filters
and 0.5 second temporal resolution data, respectively. During non-flare periods
SXT takes half resolution images in only the two thin analysis filters (Al1265Å and
Dagwood). HXT takes 2 second resolution data only in the LO energy channel. EIT
was running a CME watch programme which involves half resolution (5 arc-second)
full Sun images in Fe xii every 12 minutes for 6 hours, followed by full resolution
images in all four filters (Chapter 2.2.2). This cycle is repeated continuously for
many days. CDS was running a relatively fast active region scan study which
had a temporal resolution of 16 minutes and spatial resolution of 4x1.68 arc­
seconds. The study incorporated the lines of Fe xvi($\lambda = 361\AA$), Si x($\lambda = 347\AA$),
Si x($\lambda = 356\AA$), Mg ix($\lambda = 368\AA$), O v($\lambda = 629\AA$) and He I ($\lambda = 584\AA$) hence
covering a temperature range of approximately $2 \times 10^4 K$ to $2 \times 10^6 K$. The field of
view was 240 arc-seconds square. No radio data was examined because full coverage of all flares was not available. Data from all the instruments was calibrated using the techniques discussed in Chapter 3). The HXT image reconstruction technique used throughout this Chapter and this thesis is the Maximum Entropy Method (Chapter 3.2)

Figure 4.1 shows a GOES SXR light curve for the observing period. The flares that are studied in this chapter, Flare A, Flare B and Flare C, are indicated on Figure 4.1. Their start times, maximum times and SXR classes are listed in Table 4.1.

![GOES Light Curve of 17 September 1997](image)

Figure 4.1: GOES Light Curve of 17 September 1997, showing the three flares under study and the observing times of Yohkoh. Shaded areas mark the times of Yohkoh night. Regions between dot-dashed lines indicate satellite passage through the South Atlantic Anomaly (SAA). Flares marked with an asterisk did not occur in the active region under study. The EIT observations were continuous throughout the observing period. CDS observed the active region from 12:41UT until 15:55UT.
### Table 4.1: The timing and GOES classification of the studied flares.

<table>
<thead>
<tr>
<th>Flare Number</th>
<th>X-Ray Class</th>
<th>Start</th>
<th>Max</th>
</tr>
</thead>
<tbody>
<tr>
<td>Flare A</td>
<td>M1.7</td>
<td>11:35</td>
<td>11:43</td>
</tr>
<tr>
<td>Flare B</td>
<td>M1.0</td>
<td>17:45</td>
<td>18:03</td>
</tr>
<tr>
<td>Flare C</td>
<td>C1.2</td>
<td>13:57</td>
<td>14:10</td>
</tr>
</tbody>
</table>

#### 4.3 Individual Flare Properties.

##### 4.3.1 Flare A

The first flare in the series, Flare A, was classified as a GOES M1.7 event. Enhanced SXR emission, measured by GOES, began at 11:35UT and reached maximum emission at 11:43UT. The decay from maximum was initially fast, falling to a flux level of C1.6 by 12:10UT. The decrease in intensity was then more gradual, taking another hour to fall to pre-flare levels.

Images from SXT, taken through the Dagwood filter, during the pre-flare, impulsive and decay phases of Flare A are shown in Figure 4.2. The pre-flare image (Figure 4.2a) shows an amalgamation of SXR structures located above latitude N20. The impulsive phase image (Figure 4.2b) is dominated by a single bright SXR source, which is spatially coincident with the HXR source. The HXT image is a LO channel image constructed from a 10 second integration at 11:42UT, the time at which the HXRs reached a maximum of 35 counts/sec/subcollimator. This source is labeled Loop 3 and is located at the footpoint of the southern most part of the pre-flare structures. Although this bright patch of emission does not strictly resemble a loop, it is interpreted as an unresolved loop. The lightcurve of the flare (Figure 4.1) shows nothing out of the ordinary that would suggest an alternative explanation. Figure 4.2c shows an EIT Fe XII image taken during the impulsive phase. This image shows that the principal EUV source is also a small...
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Figure 4.2: Yohkoh and EIT images of Flare A. (a) SXT pre-flare, (b) SXT impulsive phase image, logarithmically scaled, with HXT LO channel contours, (c) EIT Fe XII impulsive phase image, (d) SXT decay phase image. Heliocentric grid lines are shown as dotted lines at 10° intervals. Longitude and latitude labels are shown in (a).
unresolved patch of emission. Figure 4.3 shows more clearly the relative location of the EUV and SXR sources. The error in alignment from the two instruments is the sum of the squares of the individual image uncertainties, i.e. a pixel, therefore around 2 arc-seconds. However there is the possibility of larger systematic errors that cannot be accounted for. Although the alignment of the EUV and SXR sources in Figure 4.3 is not exact, the centroids of emission are separated by about 10 arc-seconds, which is larger than the pointing errors. This was an M1.7 class flare from which we expect an enhancement of EUV emission. Figure 4.4 shows a lightcurve comparison of the principal sources and shows, even though the cadence of EIT is low, that the two sources evolve in a similar fashion. The similar spatial structure and size of the emission also suggests the EUV emission is a result of the flare. Flare theory (Chapter 1.4.3) implies that the emission is most likely due to footpoint emission, however EUV emission in flares can also come from loops that have cooled from the SXR loops that formed during the main phase. The inadequate spatial resolution of EIT does not allow an unambiguous determination of the nature of the source.

The decay phase image (Figure 4.2d) clearly illustrates the new loop morphology of the active region that has arisen as a consequence of the flare. The temperature and emission measure of the pre-flare structure (Loop 1) has been increased by the flare (Figure 4.5) and two new loops have appeared; one of similar size adjacent to Loop 1 (Loop 2), and another which spans the whole active region (Loop 4). The principal flare loop (Loop 3) which dominated the impulsive phase has faded.

In the corona the magnetic field pressure is greater than the plasma pressure, hence the magnetic field confines the plasma and the plasma structures can therefore be used as a tracer of the field structure (Chapter 1.2). By examining the plasma structure in SXRs we can determine an approximate magnetic configuration but not directivity of the field lines or the field strengths.

The active region studied here was located on the limb, thus it is possible to
determine to a high degree of certainty, using SXR observations, what magnetic loops are involved in the flare. Accompanying observations in HXR and EUV of the footpoints provide further clarification of the locations of individual loops. This technique to determine the schematic magnetic field structure and consequently a schematic reconnection scenario has been applied previously with success by a number of other authors (e.g. Woodgate et al. 1981; Inda-Koide et al. 1995; Nishio et al. 1997; Aschwanden et al. 1999) and its application also led to the detection of loop top Hard X-ray sources (Masuda et al. 1994).

The flare reconnection scenario that best describes the observations of Flare A, at X-ray and EUV wavelengths, is a quadrupolar reconnection scenario (Sweet 1958, 1969; Baum and Bratenahl 1980; Machado et al. 1988; Longcope 1996; Melrose...
1997; Aschwanden 1999). Figure 4.6 gives a schematic drawing of a quadrupolar reconnection scenario. The magnetic configuration is of the plus-minus-plus-minus type. A separatrix is a surface that bounds areas of similar connectivity. The intersection of two separatrices surfaces defines the separator, a single field line joining the magnetic null points at P and Q. Reconnection at the separator mapped out by the merging pre-flare loops, drives the flare creating the principal flare loop along the separator. Comparison of this model with the observations identifies Loops 1 and 2 as the merging pre-flare loops, Loop 3 as the principal flare loop in between their respective pre-flare loop footpoints. In the model of Machado et al. (1983) the high altitude loop (Loop 4) is formed by the magnetic raking of coronal plasma by the field lines that are ejected upwards from the reconnection site. A possible mechanism for the heating of the pre-flare loops is the injection and subsequent thermalisation of energetic particles from the reconnection site (Machado et al. 1983). This heating mechanism is discussed further in the analysis of Flare B.
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Figure 4.5: The time evolution of the temperature (asterisks) and emission measure (crosses) of the pre-flare loop, loop 1, showing an increase in both parameters as a result of the flare. The methods to derive temperatures and emission measures from SXT data are described in Chapter 3.1.

The occurrence of flares in a quadrupolar magnetic configuration is a well-studied topic. The resulting reconnection scenario has also previously been successfully applied to disk observations of solar flares combined with magnetograms to explain such flare characteristics as Hα ribbons (Mandrini et al. 1991, 1993; Démoulin et al. 1993; van Driel-Gesztelyi et al. 1994), chromospheric flare brightenings (Mandrini et al. 1995) and X-ray loop morphology (Machado et al. 1983).

Since this flare occurred on the limb it has been possible to identify the flare loops involved in the flare, loops 1 to 4. However projection effects at the limb could be important and it may be possible that these loops are in fact largely separated in longitude and not in fact connected. However the close comparisons of the observations with the model suggests that this flare can be explained by loop interaction.
4.3.2 Flare B

The second homologous flare in this series, Flare B, occurred at 17:45UT. The GOES lightcurve for this event (Figure 4.1) shows a fast rise to maximum emission, less than 10 minutes, in both energy channels, followed by a much slower decline than displayed by Flare A, taking more than 3 hours to return to pre-flare flux levels.

Figure 4.7 shows pre-flare, impulsive and decay phase images of Flare B from SXT, taken through the Dagwood filter, and EIT Fe XV. In the pre-flare image the two separate loops are not clearly discernible. Similarly to Flare A the impulsive phase of Flare B (Figure 4.7b) is dominated by a single SXR source that is located beneath the pre-flare structure. This bright SXR source is again spatially coincident with the HXR source. The HXT image is a LO channel image constructed from a 10 second integration at 17:52UT, the time at which the HXRs reached a maximum of 12 counts/sec/subcollimator. A full resolution EIT observation (Figure 4.7c) of the early impulsive phase, taken in Fe XV, shows this emission clearly resolved as two sources. This emission is interpreted as originating from the flare loop footpoints,
based on the correlation of HXR and EUV emission which has been proven by many previous studies. The origin of this correlation is outlined in Chapter 1.4.3. This implies that the central source is indeed an unresolved loop. Allowing for solar rotation, the main SXR sources of Flares A and B appear co-spatial to within the resolution of SXT.

There is also a smaller HXR source in the impulsive phase images located at the footpoint of the northern flare loop (Loop 1). This suggests the presence of energetic particles in the loop which produce the HXR via thick target Bremsstrahlung at the footpoint. This also accounts for the heating and filling of the loop by the process of chromospheric evaporation (Chapter 1.4.4).

The decay phase image (Figure 4.7d) shows a very similar loop morphology to that of Flare A in that two small loops connecting the intermediate dipole to the outer poles lie beneath a larger structure which connects the two outer poles. The flare loops are labeled similarly to those in the images of Flare A to highlight the comparisons.

A comparison of the loop morphology and development of Flare A and Flare B shows that the two events are remarkably similar and can be classed as homologous from the literature definition (Chapter 1.4.5). To understand the conditions that lead to the creation of two homologous flares a detailed study of the events that occurred between the two homologous flares was undertaken. The GOES plot (Figure 4.1) shows many small events superimposed on the decay phase of Flare A. The flare marked with an asterisk did not occur in the active region under investigation. A C1.2 class flare did occur in this active region, labeled Flare C and is investigated in the next sub-section.

4.3.3 Flare C

The interim flare between the two homologous flares A and B, was a C-class flare that occurred at 14:10UT (Figure 4.1). The impulsive phase of Flare C occurred
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PRE-FLARE (SXT): 16:46UT

IMPLESSIVE (SXT): 17:48UT

Figure 4.7: Yohkoh and EIT observations of the Flare B. (a) SXT pre-flare, (c) SXT impulsive phase, logarithmically scaled, with a HXR image of approximately the same time contoured, (c) EIT Fe xvi impulsive phase image, (d) SXT decay phase image. Heliocentric grid lines are shown as dotted lines at 10° intervals. Longitude and latitude labels are shown in (a).
during Yohkoh night so neither SXR nor HXR images are available. Fortunately, SoHO-CDS was observing this active region and imaged the impulsive phase of this flare. The process for making CDS-NIS images is outlined in Chapter 3.4. Figure 4.8a shows an SXT Dagwood filter pre-flare image in which the flare loops of the preceding M-class flare (Flare A), can be seen in addition to a new loop. This new loop was continually brightening in SXRs until the cessation of Yohkoh observations at 13:46UT.

The flare occurs due to the interaction between the emerging loop and the pre-existing loop (Loop 1) and is manifested by an impulsive brightening in all CDS lines (He I, O V, Mg IX, Si X and Fe XVI). These brightenings occur co-spatially with the interaction point of the two SXR loops as seen with SXT. Brightenings at the footpoints of these two interacting loops are also visible in O V and He I. These impulsive features seen in CDS are contoured on the SXR pre-flare image in Figure 4.8b. Post-flare SXR images show the presence of two loops formed at the site of the pre-flare loops and a bright source above these loops. These newly formed loops connect the footpoints of the pre-existing loops to the footpoints of the recently emergent loop.

For this flare the observational evidence supports a different reconnection scenario to that of the two homologous flares, namely that of an emerging loop interacting and subsequently reconnecting with a pre-existing loop. This emerging flux model for solar flares is described by Heyvaerts, Priest & Rust (1977). In their model flux emerging from the photosphere interacts with the overlying field. During the pre-flare heating phase (Figure 4.9a) continuous reconnection occurs in a current sheet that forms between the new and old flux. When the emerging flux reaches a critical height, depending on the merging speed and field strength, the impulsive phase begins. Electrons accelerated at the reconnection site propagate along the field lines to the chromosphere where they generate the chromospheric flare and produce chromospheric evaporation (Chapter 1.4).
Figure 4.8: Yohkoh-SXT and CDS images of Flare C. (a) SXT pre-flare, (b) SXT pre-flare with CDS He I and Fe XVI impulsive phase images contoured, (c) and (d) SXT decay phase. Heliocentric grid lines are shown as dotted lines at 10° intervals. Longitude and latitude labels are shown in (d).
Due to the absence of magnetogram observations of this flare series the presence of emerging flux during Flare C cannot be directly observed. Its presence is assumed due to the close association of the observations with the predictions of the Heyvaerts, Priest & Rust (1977) model. This loop may indeed not be emerging merely expanding or heating. However, the important point is that it interacts with the existing loop to produce a flare and rearrange the loop topology in the active region.

4.4 Active Region Evolution

By combining the schematic reconnection scenarios of the three major flares in the active region a suggested evolution of the active region has been constructed. This is shown schematically in Figure 4.10. The magnetic configuration has been inferred in the absence of magnetograms by the connectivity of the magnetic loops. In the corona $\beta \ll 1$ hence the plasma is confined by the magnetic field and can consequently be used as a tracer for magnetic field lines (Chapter 1.2).

Before Flare A the active region had an inferred magnetic configuration of the plus-minus-plus-minus type (Figure 4.10a) with Loop 1 connecting footpoints W and X. Flare A occurred within this initial configuration through reconnection
between Loops 1 and 2, creating the Loops 3 and 4 by the mechanism described in the preceding section (Figure 4.10b). Later, the emergence of a new loop below Loop 1 (Figure 4.10c) and subsequently Flare C regenerates a similar quadrupolar magnetic configuration on a smaller scale (Figure 4.10d) to that present before Flare A (Figure 4.10a). This new quadrupolar magnetic configuration is then the site of Flare B (Figure 4.10e).

Examination of photospheric magnetograms of this active region before it approached the limb revealed that the region was complex. Figure 4.11 shows an MDI
Figure 4.11: An MDI photospheric magnetogram of AR8084 from 20:00UT September 15\textsuperscript{th}, showing the existence of a quadrupolar magnetic configuration.

The magnetogram from 20:00UT on the 15\textsuperscript{th} of September when the Active Region was within 60° of disk center, which is the longitudinal limit of a magnetogram's effectiveness. This magnetogram shows that the region does indeed possess a quadrupolar magnetic configuration in an orientation that is consistent with the observations of the SXR loops. Further examination of a series of magnetograms up to the 15\textsuperscript{th} of September shows magnetic flux emergence. Both these characteristics of the photospheric magnetograms support our interpretation of the flares within the active region.
4.5 Conclusions

I have presented multi-wavelength observations of two similar flares (A and B) that occurred approximately 6.3 hours apart. They occurred in the same active region, displayed the same essential four part loop structure during the main phase, share some common footpoints and within the resolution of SXT, appear to share the same principal flare footpoints. They were also of similar GOES SXR magnitude. Therefore according to the literature definition (e.g. Woodgate et al. 1984) these flares could be classed as homologous.

In this example it is shown that the pre-flare condition for both homologous flares (A and B) is a quadrupolar magnetic configuration. Using multi-wavelength data I have made a careful study of the events occurring between the homologous flares and propose that an interim flare (Flare C) caused by the reconnection of emerging flux and a pre-existing loop regenerated a quadrupolar configuration on a smaller scale that subsequently becomes the site of the second homologous flare (Flare B). Therefore in reference to the question posed in Chapter 1.4.5, i.e. are the pre-flare conditions rebuilt after the first flare or is there a mechanism responsible for stopping the first event and triggering the second, these results support the hypothesis that the pre-flare conditions appear to be destroyed but are rebuilt by the emergence of new flux.

Previous models that have addressed homologous flares have involved the continual stressing or shearing of a single, or group of, magnetic structures (Woodgate et al. 1984; Shibata 1998; Choe & Cheng 2000). In these models the homology results from successive reconnection episodes occurring in the same sheared arcade. The continual shearing of the arcade by footpoint motions inputs magnetic energy into the arcade that is released by the flares. The importance of shear in producing flares in a quadrupolar region has also been shown (Karpen et al. 1995, 1998).

However, although these flares fit the literature definitions of homology they appear to conform better to the scenario posed in the previous section rather than
the continual shearing of a single or group of magnetic structures. Hence it is believed that a new type of homology has been discovered.

With the continual advancements in instrument sensitivity and image resolution, at all wavelengths, we can begin to see considerably more detail and fine structure in flare over a wider range in magnitude. As image resolution increases and individual flare events are seen simultaneously over larger wavelength ranges, the number of reported homologous flare observations is decreasing. It appears that the more detail we see in individual flares the differences between them become increasingly apparent.

In the light of this new type of homology we might reassess our perception of homologous flares. Do truly homologous flares exist i.e those that occur in a single magnetic structure? These are the types of homologous flares from which it is possible to gain insight into the time-scales over which energy storage in the magnetic field can occur and decipher 3D structure. Or is it that as we begin to see flares at continually greater spatial resolution, dynamic and spectral ranges will more coincidental sequences of events be revealed that could lead to apparent homology?
Chapter 5

Location of the Source of Non-Thermal X-ray Emission Line Broadening

Since the discovery of non-thermally broadened soft X-ray line profiles, there has been considerable research implemented in order to understand their origins. During solar flares the magnitude of the broadenings is greatly enhanced over active region background levels. It has been hypothesized that these non-thermal X-ray line broadenings could be related to the initial flare energy release mechanism. Hence they are a particularly important facet of solar flares. By determining the location of the source of the non-thermal broadenings it will be possible to eliminate some probable generation mechanisms and possibly determine their role in the flare process.

The aim of this chapter is to determine the location of the non-thermal broadenings seen in the \textit{Yohkoh} BCS Ca\textsc{xix} channel. To accomplish this goal I use multi-wavelength data from the SXT, BCS and HXT instruments on \textit{Yohkoh} in conjunction with TRACE. In Section 5.2 I briefly describe instrumental observations and setups and in Section 5.3 describe the basic properties of the studied flare.
In Section 5.4 I describe the method used to locate source of the $V_{nt}$ and present the results which eliminate the footpoints as the source of $V_{nt}$. In Section 5.5 I discuss the implications of the results and probable locations for the source of $V_{nt}$.

5.1 Introduction

The accurate location of the source of $V_{nt}$ is not known. Previous studies on the location of the source of $V_{nt}$ (see Chapter 1.4.6) have identified several regions in a flare where the source could be located. These regions are described below in terms of how they are created during a flare and the role they play, what mechanism could possibly generate $V_{nt}$ and why it has been postulated as a possible source. These regions are indicated on Figure 5.1, which shows a schematic diagram of a proposed unified flare model (Shibata 1999).

(1) The Reconnection Site. The site of magnetic field line reconnection and possible particle acceleration. This region was proposed as a possible $V_{nt}$ source after the work of Alexander et al. (1998) that showed that the peak of $V_{nt}$ was generally before the HXR peak, or that the levels of $V_{nt}$ were decreasing from when first measured. This implied that the $V_{nt}$ were more indicative of plasma turbulence rather than hydrodynamic flows. Plasma turbulence at the reconnection site was hence a possibility. However, at the reconnection site the density must be small to allow for efficient particle acceleration (e.g. Miller et al. 1994) and will have very small spatial scales. Thus the emission measure is likely to be low. New evidence for the presence of inflows to the reconnection region has recently been presented (Yokoyama et al. 2000), however the reconnection region itself remains unresolved. Hence this region is an unlikely source of the observed $V_{nt}$. 
(2) **The above the loop source.** This region was also this can be divided into two components: the HXR source and the SXR source. The nature of the HXR above the loop top source is still a topic of considerable debate (Fletcher 1999; Somov 1999). The two most popular explanations for the creation of HXRs above the loops are from a super-hot source ($\approx 100MK$) generated by the fast shock (Masuda 1994; Tsuneta et al. 1997; Tsuneta & Naito 1998) or from thin-thick target Bremsstrahlung from trapped electrons (Wheatland & Melrose 1995; Fletcher & Martens 1998; Metcalf & Alexander 1999). This above the loop source is believed to be confined by two slow mode shocks that extend down from the reconnection region and confine the energetic electrons by acting as magnetic mirrors (Tsuneta & Naito 1998) or simply by a converging field geometry (Fletcher & Martens 1998). The hot SXR plasma that has also been observed to originate from above the loop source.
top, is believed to be heated by the slow-mode shocks (Tsuneta 1996; Tsuneta et al. 1997).

Evidence for the $V_{nt}$ source in this region is twofold. Firstly the timing arguments of Alexander et al. (1998) and Mariska & McTiernan (1999) that $V_{nt}$ is high before the HXRs reach peak flux. Secondly the observations of limb flares (Kahn et al. 1985; Mariska & McTiernan 1999) that indicated the loop tops as the source. Tsuneta (1994, 1995) argued that turbulence in this region is could result from the impact of the reconnection jet with the closed loops leading to particle acceleration.

(3) **The SXR loop top.** Observations of limb flares (Kahn et al. 1985; Mariska & McTiernan 1999) have indicated the loop tops, not the footpoints are the source of $V_{nt}$ and unusually bright loop tops have been reported by Doschek and Feldman (1996) using SXT. The presence of turbulence and intensity enhancements can be explained by the model of Jakimiec et al. (1998) which invokes a turbulent loop-top kernel within which the magnetic field is tangled and transient current sheets occur.

(4) **Evaporating chromospheric plasma.** This is a direct consequence of the flare electron deposition at the footpoints, initiated when the energy deposition rate greatly exceeds the rate at which the energy can be conducted and radiated away. The plasma expands explosively and is driven up the magnetic loop into the corona by strong induced pressure gradients (Chapter 1.4.4). Turbulence within this evaporation flow could generate the observed $V_{nt}$.

(5) **Flare loop footpoints.** Visible at HXR, chromospheric and transition region wavelengths, these result from the deposition of flare electrons in the chromosphere at the base of the magnetic loop, where they produce HXRs via thick target Bremsstrahlung and supply heat to the chromosphere (Chapter 1.4.3).

A study of occulted and non-occulted flares by Mariska et al. (1994) showed
that non-occulted flares generally had a higher value of $V_{nt}$. Therefore the footpoints were invoked as a possible source. $V_{nt}$ is thought to result as a consequence of the thermalisation of the electron beam and the heating of the plasma.

The regions described above are created by a variety of different physical processes associated with a flare. By eliminating regions that are not responsible for producing the observed non-thermal broadenings and ultimately locating the region(s) of the flare that are responsible, we can eliminate and identify possible mechanisms responsible for the generation of $V_{nt}$.

5.2 Observations

This study utilises principally a combination of data from SXT, HXT, BCS and TRACE. These instruments are described in detail in Chapter 2. From the BCS only the Ca xix channel was used in this study. The active region in which the flare occurred was not the only one on the disk. Consequently the S xv channel, which is responsive to lower temperatures (Chapter 2.1.3), was contaminated with several sources. Hence the S xv data was unusable during the early stages of the flare that were of interest in this study. Data in the higher energy channels, Fe xxv and Fe xxvi, had insufficient counts during the early stages to produce reliable fits to the data. Yohkoh flare mode was triggered at the start of the flare and all instruments began standard flare mode observations (see Chapter 4.2).

For this flare study TRACE was observing with all EUV filters (see Table 2.4) and with the Lyman $\alpha$ (L$\alpha$) filter. Observations of the flaring active region began at 18:00UT and continued for several hours. Full resolution images were taken with a field of view of 6.4 arc-seconds square, with temporal resolutions of $\approx 120s$ and $\approx 20s$ and exposure times of 6s and 0.7s, in the EUV and UV bandpasses respectively.
5.3 Flare Properties and Evolution

The flare was a GOES M1.7 class event which occurred in NOAA Active Region 8592 (N20E40) on the 22nd June 1999 and was observed as part of the first Max Millennium co-ordinated observing campaign. The GOES start time was approximately 18:15UT with a maximum at 18:29UT in the 1→8 Å channel.

In the pre-flare stages Big Bear Solar Observatory (BBSO) Hα images show the ejection of a filament beginning at 18:10UT. After 18:10UT the filament can be seen in emission in all TRACE filters, indicating that the structure is multi-thermal. A Coronal Mass Ejection (CME) was also observed by the Large Angle Spectroscopic
Coronograph (LASCO) ejected from the North-East solar quadrant. The calculated onset time of the CME, from its velocity profile, was 18:00UT to within an hour. The error estimates on CME launch time are large for several reasons. Firstly, the height of the CME is determined by manually selecting the leading edge of the CME in a ‘difference’ image. The lack of any definition for the ‘CME edge’ when selecting, contribute significant errors to the height determination. Secondly, with only a few data points in the height time profile, five in this case, determination of the CME acceleration, via a second order polynomial fit, has limited accuracy. Finally, the CME is first detected at 3 solar radii, therefore no information of CME propagation and possible acceleration below this height is available.

During the filament eruption SXT was observing a different active region in which a GOES C2.6 class event was decaying and within which a small brightening occurred. An SXT temperature analysis of these events shown in figure 5.3 reveals that there is significant emission at a temperature of $8 \times 10^6 K$ therefore it is not possible to reliably discount Ca xix emission originating from these events. Therefore the BCS data before 18:20UT, the start of the studied M-class event, has two possible sources and hence cannot be utilized fully. SXT began observing AR8592 at 18:20:30UT when flare mode was triggered.

The flare displays two footpoint regions that outline the base of the flare arcade and are visible in HXR and TRACE images (Figure 5.2). The footpoints are located on either side of a simple photospheric magnetic neutral line (Figure 5.2a). In TRACE $\text{La}$ images these footpoints are observed to move apart, away from the neutral line, as the flare progresses. Both footpoint areas are elongated in shape with the longer axis parallel with the neutral line (Figure 5.2b). The south footpoint area exhibits stronger emission at both $\text{L}\alpha$ and HXR wavelengths, indicating greater energy deposition. The magnetic field strength at both footpoints was calculated by aligning the TRACE $\text{L}\alpha$ and MDI magnetogram and summing magnetogram pixels that corresponded to the footpoint area. The magnetic field strength at the south footpoint was weaker than the north footpoint, in agreement
with the results of Sakao (1994). He found that greater electron deposition occurs at the magnetically weaker footpoint due to reduced mirror forces experienced by the electrons propagating along the magnetic field.

HXT images of the impulsive phase show two footpoints, one on either side of the photospheric neutral line. These footpoints are visible in the LO (Figure 5.2e), M1 (Figure 5.2f) and M2 channels. The HXT images were constructed using the Maximum Entropy Method (Chapter 3.2). The HI channel does not contain enough counts to construct an image (Figure 5.4). Images in the low channel also show a third HXR source located between the two sources seen at higher energies. Figure 5.2e shows the LO channel HXR image in comparison to a gradual phase SXT image which illustrates the position of the flare loops. It can be seen that the HXR source is above the SXR loops, slightly offset to the south from the loop apex. This source is not present in the higher energy channels so no diagnostics are possible. From TRACE images (Figures 5.2b and c) we can clearly see that this 3rd HXR source is not co-spatial with a footpoint. Therefore it is likely that this is a coronal HXR source. The relative co-alignment between SXT and HXT is believed to be
less than 1 arc-second (Masuda). The lack of counts in the higher energy channels suggests this loop source is thermal. There are also too few images to create reliable lightcurve comparisons with the footpoint sources. Lightcurve comparisons of loop top and footpoint HXR sources are useful to determine if the two sources vary co-temporally. If they do, then it might be assumed that they are generated by the same electron population. The double footpoint and loop top source are only visible from 18:21→18:23UT, the time at which the HXR flux in the LO, M1 and M2 channels was high (Figure 5.4).

An arcade of flare loops connecting the two footpoints is visible first (18:24UT) at high temperatures (SXR and Fe xxiv, Figure 5.2d) and then later (18:35UT)
at cooler temperatures (Fe\textsc{i}x and Fe\textsc{xii}, Figure 5.2c). These cool flare loops continue rising, visible in TRACE Fe\textsc{xii} and Fe\textsc{i}x lines, until at least 21:00UT. The orientation of the flare loop arcade and footpoints is such that the top of the loop system is observed against the dark background of the disk. Hence the signal from the top of the loop arcade is not contaminated by emission from the footpoints.

The combined multi-wavelength images of this flare suggest it was of the type described by Carmichael (1964), Sturrock (1966), Hirayama (1974) and Kopp and Pneuman (1976) (hereafter CSHKP) and subsequently by other authors (e.g Cargill & Priest 1983; Forbes \textit{et al.} 1989). In this model reconnection begins low in the corona after the ejection of mass, and proceeds upwards creating a SXR loop arcade that increases in height with footpoints that move apart with time. The individual hot SXR loops then cool through radiation and conduction to form the EUV and eventually H\alpha loops (Švestka 1987; Antiochos & Sturrock 1976, 1978; Cargill \textit{et al.} 1995). The model is shown schematically in Figure 5.5.

![Schematic diagram of the CSHKP reconnection model](image)

**Figure 5.5:** Schematic diagram of the CSHKP reconnection model (from Kopp & Pneuman 1976). The rising neutral point reconnects successively higher field lines, resulting in rising SXR loops and footpoints that move away from the neutral line.
5.4 Location of the Source of $V_{nt}$

In this work I examine the Ca XIX resonance line and the $n \geq 3$ satellite lines. Plasma parameters are calculated from a best fit theoretical spectrum to the observed data. Two sample BCS spectra are shown in Figure 5.6 from the time of $V_{nt}$ maximum and the maximum in BCS Ca XIX counts. The plots show that the observed data are well represented by a single component fit at both times. A single component fit implies that the spectra were fitted using just one value of $T_e$ and $E_M$, implying a single isothermal plasma. Two component fits to the data were attempted and although these yielded similar $\chi^2$ values, the calculated plasma parameters were unphysical: blue shift velocities were zero and $T_e \approx 300 MK$. Two component fits implies that the data were fitted using two values of $T_e$ and $E_M$, implying that the flare consisted of two distinct plasma populations distinguishable due to their different velocities along the line of sight. This is the case for flares that show a strong evaporating component. The BCS Ca XIX lightcurve and the derived plasma parameters are shown in Figure 5.7. The total count rate in the Ca XIX channel has been corrected by a dead-time factor, which varies during the flare, and has a maximum value of two. This value is large and results from the large count rate in the S XV channel of $\approx 9000$ counts per second. Since the S XV and Ca XIX channels share the same detector the dead-time correction factor is calculated from the combined counts in the two channels.

Figure 5.8 shows the relative timings of the non-thermal broadenings and the M1 channel HXR flux. The time of $V_{nt}$ peak can be well determined for this event and is seen to occur after the first small HXR peak ($\approx 18:20$ UT) and before the maximum in HXRs.

Figure 5.7e shows that the wavelength shift of the main resonance line changes complementary to the $V_{nt}$ throughout the flare. The spectral resolution of the Ca XIX detector is $0.53 m \AA$, therefore the measured centroid shift is approximately two pixels ($\approx 100 km s^{-1}$). Assuming that these line shifts are representative of bulk
plasma motions associated with chromospheric evaporation, then the $100\,\text{km}\,\text{s}^{-1}$ shift places a lower limit on the up-flow velocity. In Chapter 2.1.3 it was mentioned that due to the orientation of the BCS crystal line shifts can result from a latitudinal movement of the source. Bragg's law states that $n\lambda = 2d\sin \theta$ where in the case of the Calcium channel on BCS $2d = 4.00\,\text{Å}$. This gives,

$$d\lambda = (2d) \cos \theta \, \text{d}\theta \approx 2.02 \times 10^{-5} \phi,$$

where $\phi$ is the angle subtended at the Earth in arc-seconds. Therefore in order to account for the observed line shifts due to this effect the source must move approximately 51 arc-seconds. Such source shifts are not evident in movies of the flare. Hence we can assume that the wavelength shifts are due to plasma velocities. However, a latitudinal distribution of emission combined with this effect can result in an increased line width, and it is possible that this accounts for the increased line widths of background emission (Sterling 1997). However during a flare the emission source becomes more compact with an associated increase in line width, therefore another broadening mechanism is required.
Figure 5.7: The results from the spectral fitting of the BCS Ca XIX data.
Figure 5.8: The relative timings of the non-thermal broadening and HXR flux.

The temperature responses of the SXT filters and the BCS Ca xix channel are shown in Figure 5.9. This plot indicates that although SXT is more sensitive to lower temperature plasma (i.e. \( T_e < 10\,MK \)) it also detects plasma at Ca xix temperatures. To locate the dominant source of Ca xix emission images taken by SXT through the two thickest filters, Al12 and Be119, are used.

SXT was used to determine the spatially resolved distribution of temperature and emission measure within the flare area. When generating SXT temperature and emission measure images the SXT images are summed over the same time range as the BCS data to improve the counting statistics. This increases the accuracy of the calculated plasma parameters and makes the values derived from both instruments directly comparable. The method for calculating temperatures and emission measures from SXT images is described in detail in Chapter 3.1. The errors quoted here are those derived from Poisson noise in the data. Chapter 3.1 described how to compensate for scattered light in SXT images. For this flare no full frame SXT images were available to calculate an accurate scatter function, however I estimated the scattered light using only the PSF function. Figure 5.10
Figure 5.9: The responses of BCS Ca xix channel and SXT filters as a function of temperature.

shows an image illustrating the percentage levels of scattered light in the Be119μm image at 18:21UT on which is contoured the DN image. The levels of scattered light are below 15% for the area of interest above an between the two bright footpoints. Following Tsuneta (1997) in a similar piece of analysis on the 13th Jan 1992 flare, assuming that scattering affects the thick aluminum and beryllium filters in much the same way and that the filter ratio has the same error as the scattering ration this leads to temperature errors of less than 10%.

Figure 5.11 shows temperature, emission measure and intensity images calculated from SXT from times around the $V_{nt}$ maximum. Pixels in the SXT temperature maps that had a temperature within one standard deviation of the BCS temperature at that time were averaged to produce the derived SXT parameters. These pixels are enclosed by the contours in the SXT intensity image. The derived SXT temperatures are thus similar to that from BCS. If the emission measures from
SXT are similar to that derived from BCS then we can assume that the two instruments are observing essentially the same plasma. At 18:20:47 UT (Figure 5.11a) the emission measures differ by less than a factor of 2 within the errors and at 18:21:29 UT (Figure 5.11b) the values are the same to within the errors. Thus early in the flare, around the time of $V_{nt}$ maximum SXT and BCS are observing approximately the same plasma.

Figure 5.12 shows a plot of the SXT emission measure and temperature derived this way, along with the BCS temperature and emission measure, as a function of time. The plot shows that early in the flare the derived emission measures differ
Figure 5.11: This figure shows an SXT intensity, EM and temperature image of the flare from 18:20:47 → 18:21:29UT, the time at which $V_{nt}$ peaked. The pixels bounded by contours have a temperature within one standard deviation of the BCS temperature for the same time interval.
only slightly and change co-temporally. However, as the flare progresses the differ­
ces become more significant. The BCS emission measures are also systematically
higher than the SXT emission measures. The systematic differences imply that not
all the hot component is seen by SXT, some of it is superimposed along the line
of sight with cooler plasma (cf. Doschek 1999). The divergence indicates that this
effect becomes stronger in the later stages of the flare. Cooler plasma along the
line of sight will dominate SXT's response, resulting in lower derived temperatures.
Hence pixels in which the hot component may be present along with cooler plasma
will not show a high temperature in SXT and will not be within one standard
deviation of the BCS temperature. These pixels will not therefore contribute to
the SXT derived temperature and emission measure of the hot component but will
be included in the BCS derivation, hence the discrepancy. Therefore, only before
18:23:00UT can we reliably say that SXT and BCS are observing essentially the
same plasma.

The hot pixels enclosed by the contours in Figure 5.11a and b are spread over
the whole flare area. However, the dominant emission will come from those hot
pixels where the emission measure is also high. These high emission measure, high
temperature pixels are located in an area extended along the flare loops above the
footpoint emission.

Wülser et al. (1994) used a similar method of comparing SXT and BCS temper­
atures and emission measures in the early phase of a solar flare. In their analysis
the derived plasma parameters from BCS and SXT differed by similar amounts
to the values calculated here and led them to conclude that BCS and SXT were
observing the same plasma. This enabled them to determine that the location of
up-flowing SXR plasma, detected by BCS, was within a SXR flare loop connecting
two HXR footpoints.

To verify the result that the BCS is detecting only the flare loop plasma and
no additional sources, the lightcurves of the BCS Cα x i x with the SXT pixels
(DN/s; Be119μm filter) that have a temperature in the range of \( \log T_e = 7.1 \rightarrow 7.3 \)
Figure 5.12: The variation of temperature and emission measure with time for the SXR flare loops, from SXT (thin solid line) and BCS (thick dashed line).

(the range over which BCS CaXIX is most sensitive; Figure 5.9), are compared. The results are shown in Figure 5.13. The fact that these two lightcurves behave similarly (in the early stages of the flare) suggests that both instruments are observing the same plasma and that neither instrument is detecting an additional source. Also shown in Figure 5.13 is the lightcurve of the SXT cool component ($\log T_e = 6.6 \to 6.9$), the accumulation of this lower temperature plasma inhibits the ability of SXT to detect the hot component, hence the discrepancies between the hot component and BCS lightcurves after 18:24UT.

It is now well established that during solar flares the TRACE 195 Å image becomes considerably contaminated by emission from Fe XXIV ($\lambda=192$ Å, Feldman et al. 1999; Warren et al. 1999; Warren 2000). Fe XXIV and Ca XIX ionization fraction functions both peak at the same temperature (Arnaud and Rothenflug 1985; Arnaud & Raymond 1992), therefore emission in these two spectral lines should originate from approximately the same plasma sources. By comparing TRACE 195
CHAPTER 5. Vnt SOURCE LOCATION

Figure 5.13: Comparison of the BCS CaXIX lightcurve (solid line) with the lightcurve of the SXT hot (dashed) and cool (dot-dashed) component.

Å images with TRACE 171 Å (Fe IX) and 284 Å (Fe XV) we can determine which features seen in the 195 Å image are due to Fe XII and which are from Fe XXIV. This is accomplished simply by assuming that any feature that appears in Fe XII and not at either Fe XV or Fe IX must come from Fe XXIV. In Figure 5.14 TRACE difference images in each EUV filter are shown along with a temperature map from SXT from the time of maximum Vnt. The difference images were constructed by subtracting a pre-flare image from 18:00UT. This illustrates that the hot SXT component sources are co-spatial with the Fe XXIV emission in the TRACE 195 Å images. With the greater spatial resolution of TRACE it is possible to determine more accurately where, within the flare area, this Fe XXIV emission comes from. It is within and above the flare loops, offset to the south from the loop apex. This is co-spatial with the location of the loop top HXR source. Figure 5.15 shows that the HXR loop top source is co-spatial with the hottest area of the SXT temperature map of 18:21UT. Figure 5.2e compares the SXT temperature map of 18:21UT with contours of the SXR loops that form in the main phase (18:25UT), and shows that
Figure 5.14: TRACE difference images in all three filters and an SXT temperature image. The feature that appears in the Fe xii filter and not the other two is judged to be an Fe XXIV feature. Note that this is co-spatial with the hot component in the SXT temperature image.

although the hot area extends over a region that covers these loops the hot area is located mainly above them. Consideration must also be taken for the fact that this is a two ribbon flare, thus the flare loops increase with size as the reconnection occurs at increasing heights in the corona.

I have shown using BCS and SXT temperature and emission measure comparisons and TRACE Fe XXIV observations, that the dominant source of BCS Cα xix emission at the peak value of $V_{nt}$ is located within and above the SXR flare loops. Figure 5.12 shows the variation of temperature and emission measure of the flare loops (i.e. above the footpoints). The emission measure can be seen to increase steadily as the flare progresses. Figure 5.16a and b show an SXT intensity, emission measure and temperature map at later times in the flare. These figures illustrate
that the area occupied by the hot plasma is expanding. Note that after 18:24 UT the BCS derived emission measure becomes significantly greater than SXT, this is because the increased amounts of cooler plasma component \((\log T = 6.6 \rightarrow 6.9)\) masks SXT’s ability to detect the hot component (cf. Figure 5.13) as discussed by Doschek (1999). Figure 5.17 shows selected SXT images during the rise phase of the SXR emission. The images show that the loops are becoming brighter. I showed in Section 3 that this flare is a two-ribbon (CSHKP) flare, this model relies on chromospheric evaporation to create the hot SXR loops. Figure 5.12 shows that the emission measure in the loops is increasing, therefore according to the model the loops are filling with SXR plasma. Unfortunately no proper motions are apparent in the SXT or TRACE images. In the TRACE images this likely results from the relatively long exposure times (6 seconds) and inadequate temporal resolution (120 seconds). In the SXT images the spatial resolution (5 arc-seconds) may be too low to observe the evaporation flows. Direct observations of chromospheric evaporation flows with SXT have been observed in very few events (Savy 1997).

Chromospheric evaporation has become the generally accepted theory to explain the increase in emission measure during a flare. The evaporation of chromospheric
CHAPTER 5. $V_{nt}$ SOURCE LOCATION

(a) Comparison of BCS and SXT derived plasma parameters
18:22:56UT → 18:23:26UT

<table>
<thead>
<tr>
<th>Parameter</th>
<th>BCS</th>
<th>SXT</th>
</tr>
</thead>
<tbody>
<tr>
<td>$E_M$</td>
<td>$2.84 \pm 0.21 \times 10^{18}$ cm$^{-3}$</td>
<td>$2.11 \pm 0.067 \times 10^{18}$ cm$^{-3}$</td>
</tr>
<tr>
<td>$T_e$</td>
<td>$16.2 \pm 0.52$ MK</td>
<td>$15.6 \pm 1.6$ MK</td>
</tr>
</tbody>
</table>

(b) Comparison of BCS and SXT derived plasma parameters
18:26:26UT → 18:26:56UT

<table>
<thead>
<tr>
<th>Parameter</th>
<th>BCS</th>
<th>SXT</th>
</tr>
</thead>
<tbody>
<tr>
<td>$E_M$</td>
<td>$6.43 \pm 0.32 \times 10^{18}$ cm$^{-3}$</td>
<td>$3.67 \pm 0.095 \times 10^{18}$ cm$^{-3}$</td>
</tr>
<tr>
<td>$T_e$</td>
<td>$15.7 \pm 0.31$ MK</td>
<td>$15.2 \pm 1.5$ MK</td>
</tr>
</tbody>
</table>

Figure 5.16: SXT intensity, EM and temperature image of the flare from (a) 18:22:56 → 18:23:26UT and (b) 18:26:26 → 18:26:56UT. The pixels bounded by contours have a temperature within one standard deviation of the BCS temperature for the same time interval. In (a) the BCS and SXT emission measure and temperature are still similar. In (b) the BCS and SXT emission measures now differ significantly.
plasma occurs as a natural physical response to the release of energy in the corona and the subsequent transport of energy down the field lines to the chromosphere. The plasma is then explosively heated to SXR temperatures and is forced up the flare loop (Chapter 1.4.4). This process has been successfully modeled by many authors (Fisher, Canfield and McClymont 1985a, b, c; Mariska, Emslie and Li 1989; Yokoyama and Shibata 1998) and agrees well with the observations of SXR flares (Hori et al. 1997, 1998; Yokoyama and Shibata 1998). The CSHKP model of two-ribbon solar flares, which is believed to well describe this flare, relies on chromospheric evaporation to fill SXR loops.

It has been shown that the energy spectrum of the incident electron beam onto the chromosphere affects the efficiency of converting electron energy into chromospheric evaporation. The steeper the spectrum the more input electron energy will be used to drive chromospheric evaporation rather than be radiated or conducted away from the energy deposition site (Mariska, Emslie and Li 1989; Antonucci et al. 1993; McDonald et al. 1999). Figure 5.18b shows the values of the photon spectral index, $\gamma$, at various times during the flare calculated from the HXT LO and M1 channels. The photon spectral index is related to the electron energy spectral index ($\delta$), assuming thick target HXR production, by; $\gamma = \delta - 1$ (Tandberg-Hanssen & Emslie 1988). The photon spectrum during the early phases of the flare 18:20:10 → 18:21:00UT is very steep, hence chromospheric evaporation will occur at this time. However, it is noted that the value of the spectral index may be an over estimation due to contamination of the HXT LO channel by a thermal source. Figure 5.18c shows spectral indices from the same time calculated using the M1 and M2 channels and shows that the values are similar.

The footpoint separation is $\approx 19.5 \times 10^3 km$ from TRACE L\(\alpha\) images. Assuming semicircular loops this gives a half loop length of $\approx 15 \times 10^3 km$. The evaporating plasma is assumed to travel up the loop at speeds ranging from a lower limit set by BCS observations of the line centroid shift ($100 km s^{-1}$) to an upper limit of the
sound speed in the SXR loops. The sound speed, \( c_s \), is given by;

\[
c_s = \left( \frac{\gamma r P}{\rho} \right)^{1/2}
\]

where the pressure, \( P = nkT \), the density, \( \rho = nm_i \), \( n \) is the number density, \( k \) the Boltzmann constant, \( \gamma_r \) the ratio of specific heats and \( m_i \) the ion mass. This expression reduces to,

\[
c_s = 152T^{1/2} \text{km}s^{-1}.
\]

The temperature of the flare loops is calculated from SXT observations to be \( \approx 1.2 \times 10^7 K \) which gives a sound speed of \( \approx 530 \text{km}s^{-1} \). Therefore the time taken for evaporating SXR plasma to reach the loop apex is \( \approx 25 \rightarrow 150 \text{s} \). During the early stages of evaporation the pressure in the pre-flare loop will be at its lowest level and the pressure in the evaporating plasma will be high. Therefore the up-\( \text{flow velocity, driven by the pressure gradient, will be at its highest value, close to the sound speed. Hence after the start of evaporation, it will take only } \approx 25 \text{s for plasma to reach the flare loop apex. Therefore the small HXR burst at 18:20UT (Figure 5.8) is consistent with the presence of evaporating plasma near the loop apex at 18:20:30UT and thus consistent with evaporating plasma as the source of } V_{nt}.

5.5 Conclusions

Using the spatially resolved images of a two ribbon solar flare, from SXT, HXT and TRACE and the spectral information from BCS, I have determined the location of the non-thermal soft X-ray broadenings. Using SXT and BCS temperature and emission measure comparisons and TRACE Fe XXIV observations the location of the dominant Ca XIX emission has been shown to be an extended region within and above the SXR flare loops.

The conclusion therefore is that the source of the non-thermal line broadenings is not at the flare footpoints but a region within and above the SXR flare loops.
This conclusion is in agreement with that of Khan et al. (1995) and Mariska & McTiernan (1999) who, using observations of partially occulted flares, also found that the source of the $V_{nt}$ was not located at the flare footpoints but within the flare loops.

Having excluded the flare footpoints as a possible source of $V_{nt}$, there remain three alternatives: an above the loop SXR source, the SXR loop top, evaporating plasma or a combination. I do not include the reconnection site for the reasons discussed in Section 5.1. The evidence pertaining to each region is outlined below.

The spatial coincidence of the hottest SXR region and the HXR loop top source, both of which are above the flare loops, suggests the source of $V_{nt}$ is also above the SXR loops, a region postulated by Tsuneta (1996, 1997) to be heated by the slow shocks extending from the reconnection region. Tsuneta et al. (1997) calculated a
temperature of \( \approx 20M\text{K} \) for this hot above the loop top region, using SXT, for an M2 class single loop impulsive flare. This temperature is not dissimilar to the temperature derived in this flare. This hot above the loop region could form before the start of the detectable HXR burst and hence account for the increased levels of \( V_{nt} \) before the start of the HXR burst, a feature common to many flares (Alexander et al. 1998). Line of sight effects could also be significant in this flare because of its location (N20E40). If the hot SXT component was located above the SXR loops, as has been observed in other flares (Tsuneta 1996; 1997), then seen in projection it could appear to be partially located within the SXR loops.

In support of evaporating plasma as the source of \( V_{nt} \) is the fact that the Ca xix source location extends over a large area that incorporates the tops of the flare
loops. Also the observations of the Ca xix line centroid shifts in the BCS data, suggesting bulk plasma motions, which change complementary to the values of $V_{nt}$ imply a possible causal relationship. Results of recent numerical simulations of chromospheric evaporation in solar flares (Hori et al. 1997, 1998; Yokoyama & Shibata 1998) suggest that in the early phases the hot plasma is located in the uppermost flare loops and is evaporating from the chromosphere. Hence the location of the Ca xix plasma derived in this chapter is in agreement with these simulations. I have shown that the approximate travel time of evaporating plasma from the flare footpoints to the loop top is $\approx 25$ s. The evaporating plasma is driven by the electrons producing the hard X-ray burst, therefore the elevated levels of $V_{nt}$ after the first small HXR peak are consistent with association to evaporating plasma. The presence of HXRs before they are detected by HXT must not be ruled out, since flux levels may have been below the instrument detection threshold. In this case evaporation would begin earlier and further strengthen the association of $V_{nt}$ with evaporating plasma.

The finding of Sakao (1994) that greater electron deposition and brighter HXR emission are associated with weaker magnetic field strengths has already been referred to in section 3. It is thus interesting to note that the TRACE Fe xxiv observation in Figure 5.14 shows that the hot SXR source is not in fact located exactly at the loop top but is slightly offset, along the loop axis, towards the south footpoint which has weaker associated magnetic field and stronger HXR emission. This image was obtained at 18:21:24UT, very early in the HXR event when evaporating plasma might still be in the process of propagating up from the footpoints. Therefore this hot source could be evaporating plasma and is hence consistent with evaporating plasma as the source of $V_{nt}$. This asymmetry is also apparent in the SXT intensity images in the early stages of the flare (Figure 5.17b).

I have presented arguments that the source of the $V_{nt}$ is associated with either an above the loop top source or evaporating plasma. Both sources are consistent with the observations and I believe that neither can be excluded with the present
These results highlight the need for high spatial resolution observations of line profiles in solar flares, enabling us to obtain SXR spectra for isolated regions in the flare simultaneously, thus allowing the unambiguous determination of the source location of $V_{nt}$. Line profiles would be required over the whole flare area with a time resolution down to a few seconds in order to separate the pre-flare, impulsive and decay phases of the flare. Observations of transition region and chromospheric line profiles will also help in understanding the origins of $V_{nt}$. The Solar-B/EUV Imaging Spectrometer (EIS) scheduled for launch in 2005 will possess this capability. In the near future images from the High Energy Solar Spectroscopic Imager (HESSI) may indicate the location of hot thermal plasma in flares. These could also be used in comparison with BCS data to help locate the source of the non-thermal broadenings.
Chapter 6

Relative Timings of Hard X-ray Flux and Non-thermal Broadening of X-ray Emission Lines

In this Chapter I continue the study of non-thermal line broadenings in solar flares by investigating their temporal behaviour. By incorporating the results from the previous chapter, that the $V_{nt}$ may be due to evaporating plasma or plasma above the loop top that is associated with the initial energy release, I attempt to distinguish which mechanism is dominant through comparisons with the time evolution of the Hard X-ray (HXR) emission.

I study 59 solar limb flares using the Bragg Crystal Spectrometer (BCS) and the Burst and Transient Source Experiment (BATSE) to investigate the relative timings between the Hard X-Ray (HXR) emission and the observed non-thermal broadenings of X-ray emission lines ($V_{nt}$). In Section 6.3 I present several observational results that imply a causal relationship between the HXR flux and $V_{nt}$. I also show that the temporal evolution of $V_{nt}$ in flares is dependent upon the
characteristics of the HXR burst. The implications are discussed with a view to understanding the mechanism of $V_{nt}$ generation.

6.1 Introduction

Previous studies of the timing relationships between $V_{nt}$ and HXRs (Alexander et al. 1998; Mariska & McTiernan 1999) showed that the $V_{nt}$ peaks early in the HXR burst often before the HXR maximum flux but after the first significant HXR peak. Such behaviour is consistent with the $V_{nt}$ being related to the initial energy release process rather than a hydrodynamic response of the electron deposition (Alexander et al. 1998). Alexander et al. (1998) also note that the decay phase of $V_{nt}$ displays no strong signature of subsequent individual HXR bursts.

Attempts to locate the source of the $V_{nt}$ within a flare structure (Khan et al. 1995; Mariska & McTiernan 1999; Chapter 4) have eliminated the flare footpoints as possible source locations and indicated that the source is within or above the flare loops. However, these studies could not differentiate between the source of $V_{nt}$ as either evaporating chromospheric plasma or plasma that is related to the initial flare energy release, since both possibilities were consistent with the results.

In the model for $V_{nt}$ generation in a turbulent evaporation flow we assume that the magnitude of $V_{nt}$ is directly linked to the amount of electron energy converted to evaporation. Therefore large inputs of electron energy to the chromosphere, signaled by a HXR peak should be followed by a $V_{nt}$ peak. However if the $V_{nt}$ is generated by plasma associated with the initial energy release, then the $V_{nt}$ peak would be expected to precede or be coincident with the HXR peaks. The model of Tsuneta (1994, 1995) is an example of such a model. Turbulent plasma above the loop top is generated by the collision of the reconnection jet. The time varying turbulence subsequently accelerates the electrons that produce the HXR burst. It should therefore be possible, in the simplest scenario, to distinguish between these two models based purely on the timing relationships between $V_{nt}$ and HXR flux.
6.2 Observations and Instrumentation

This study uses data taken with the Burst And Transient Source Experiment (BATSE) on board the Compton Gamma Ray Observatory (CGRO) and the Bragg Crystal Spectrometer (BCS) on board Yohkoh. These instruments are described in detail in Chapter 2.

To obtain the dataset used in this analysis the BATSE solar flare data base was automatically searched from Yohkoh launch (30th August 1991) to CGRO decommissioning (4th June 2000), for all events that exceeded a peak count rate of 500 counts per second in BATSE's low energy channel (25 → 50keV). Flares with a count rate greater than this threshold appeared to have a correspondingly good signal in the BCS Ca xix channel. The resulting flares were then included into a preliminary data set if they;

a) Occurred at an angle greater than 60° from Sun centre. This criterion was set in an attempt to exclude any blue-shifted plasma components in the line profiles.

b) Occurred during Yohkoh day and not during an SAA passage by the satellite.

For all the flares matching the above criteria the corresponding spectral data from the BCS Ca xix channel were fitted using the standard BCS software. The fitting procedure generates theoretical spectra which are fitted to the data via \( \chi^2 \) minimization (Chapter 3.3). The resulting fitted data were then examined manually for flares that showed a well defined peak in \( V_{nt} \). Any events in which a blue-shifted component was present were discarded. Due to the orientation of the Bragg crystal in the BCS, the wavelength dispersion axis is aligned with solar longitude. Therefore emission lines from sources at different latitudes appear at different wavelengths in the BCS detector. If two sources were determined to be
present during the early stages of a flare, then it was discarded. This procedure yielded 59 events for further analysis.

The reason for discarding flares with a blue-shifted or secondary component is the difficulty they represent when spectral fitting. Figure 6.1 shows a BCS Ca xix spectrum from the flare of 16th December 1991, that shows a distinct blue-shifted component (Phillips 1996). Note that the $n = 3$ satellite lines of the blue-shifted component are superimposed on the main resonance line of the primary component. It is therefore not possible to uniquely determine the temperature of both components and we must make certain assumptions during the fitting process, i.e. that both components have the same temperature. The evaporative flows that these blue-shifts represent have been shown to be approximately radial (Mariska, Doschek & Bentley 1993), hence observed on the limb these flows are perpendicular to the line of sight. Therefore the evaporative component does not appear shifted from the primary component. Therefore although the limb flares show no evidence

![BCS Ca xix Spectrum](image)

Figure 6.1: An example BCS Ca xix spectrum from the 16th December 1991 flare that shows a distinct blue-shifted component. The spectrum has been fitted with two components. The primary/stationary component is shown with the solid line and the secondary/evaporating component with the dashed line.
for bulk flows they are still believed to be present, just undetectable. We have no reason to believe that limb flares are any different from disk flares.

6.3 Analysis

In the absence of any other a priori information on the plasma state, when analysing BCS data, it must be assumed that the plasma is isothermal. From the flare spectrum we then obtain $T_e$, $T_d$ and consequently $V_{nt}$ of the whole flare. From all the flares studied the following parameters were calculated:

a) time and magnitude of the main $V_{nt}$ peak,

b) time and magnitude of the main HXR peak,

c) the start time and rise time of the HXR emission.

The start time of the HXR emission was defined as the time at which the flux exceeded $3\sigma$ above the pre-flare background level and the rise time was defined as the time elapsed from the HXR start time, to time of maximum emission.

Figure 6.2a shows a plot of the HXR rise time ($t_{rise}$) versus the time delay between the HXR flux peak and the $V_{nt}$ peak ($t_{delay}$). The errors shown on the delay time measurements originate principally from the determination of the peak time of $V_{nt}$. The errors on the rise time measurements are no bigger than the plotting symbols. A negative $t_{delay}$ implies that the $V_{nt}$ peak occurred before the HXR peak. From the whole data set, 22 flares had a positive time delay and 37 a negative time delay. The greater number of negative delay flares is in agreement with the results of Alexander et al. (1998); however, this may result from selection effects. Some flares that were originally examined for inclusion into the final data set had short HXR rise times but no BCS spectra were obtainable before the HXR peak, hence these events were not included in the dataset because no discernible $V_{nt}$ peak was observed.
Figure 6.2: (a) The relationship between the HXR rise time and the delay time between the HXR and $V_{nt}$ maximum, (b) the same as (a) but on a smaller scale, (c) The average time delay for flares with a particular number of subsidiary peaks. A negative delay time implies that the $V_{nt}$ peak occurred before the HXR peak.

In the introduction I described how the turbulent evaporation flow model should show $V_{nt}$ peaks after HXR peaks, whereas models invoking turbulent plasma associated with the initial energy release should display $V_{nt}$ peaks before HXR peaks. In this sample both cases are observed, suggesting that there may be two classes of flares that are consistent with both scenarios.

Figure 6.2a shows that $t_{\text{rise}}$ and $t_{\text{delay}}$ are related. A longer HXR rise time implies a greater time delay between the $V_{nt}$ and HXR peaks. A correlation analysis on $t_{\text{rise}}$ and $t_{\text{delay}}$ gave a Spearmans co-efficient of $\approx 1 \times 10^{-9}$, indicating a significant correlation, i.e. greater than 95

$$r_{\text{rank}} = 1 - \frac{6\Sigma D^2}{N(N^2 - 1)},$$

(6.1)
where, $D$, is the difference between $X$ and $Y$ values and $N$ the number of pairs of values. In the models of $V_{nt}$ generation outlined in the introduction the HXR flux and $V_{nt}$ are proposed to correlate in time with a possible time lag, but neither model predicts the variation of $V_{nt}$ peak timing with the HXR rise time. Therefore other factors must also be present that influence the timing of the $V_{nt}$ peak. I will discuss these in Section 6.4. Figure 6.2b shows how the average $t_{\text{delay}}$ varies for flares with a particular number of subsidiary HXR peaks that occur before the main HXR peak. Subsidiary HXR peaks are small transient increases in HXR emission that occur before the maximum in HXR emission. The subsidiary peaks for the 14th June 1999 flare are in Figure 6.3 by vertical dotted lines. The greater the number of subsidiary peaks the earlier the $V_{nt}$ peak occurs before the HXR peak.

To further investigate the relationship between $t_{\text{rise}}$ and $t_{\text{delay}}$ the sample of flares was separated into two categories: impulsive rise flares and gradual rise flares. An impulsive rise flare has a HXR profile that has a sharp rise to maximum emission in a single smooth peak, an example is shown in Figure 6.3a. A gradual rise flare has a slow rise to maximum with one or more subsidiary peaks (Figure 6.3b). Historically flares with gradual HXR emission are more commonly related to long duration events (LDEs) and two ribbon flares, whereas flares that display impulsive HXR lightcurves are often related to compact flares (Bai & Sturrock 1989).

Histograms of the measured parameters for these flares are shown in Fig 6.4 for both impulsive rise flares and gradual rise flares. The total number of gradual rise flares is 35 and impulsive rise flares 24. The histograms show a clear tendency for impulsive rise flares to have a shorter delay time than gradual rise flares. The shorter rise time of the impulsive rise flares is also evident. No distinction between gradual rise flares and impulsive rise flares is evident from maximum HXR flux, however there is a possible tendency for impulsive rise flares to exhibit a larger maximum $V_{nt}$. In impulsive rise flares the $V_{nt}$ peak occurs after the main HXR peak in 66% of events. In gradual rise flares the $V_{nt}$ peak occurs after the main HXR peak in 25% of events. These results indicate that there are systematic
CHAPTER 6. HXR AND $V_{nt}$ RELATIVE TIMINGS

Figure 6.3: Graphs showing $V_{nt}$ (with error bars) and HXR flux for two flares in the dataset. (a) an example of an impulsive rise HXR burst and (b) an example of a gradual rise HXR burst.

differences in the behaviour of $V_{nt}$ in impulsive and gradual rise flares.

Figure 6.5 shows $V_{nt}$, HXR and BCS count rates for two flares in the data set. These two examples clearly show that when a large HXR pulse occurs $V_{nt}$ also increases. This type of behaviour occurred in only $\approx 20\%$ of the studied flares of which 75% were gradual rise flares. Note that in the examples in Figure 6.5 the strong HXR peak and corresponding $V_{nt}$ peak occur at times when the BCS count rate is increasing but still at low values.

Figure 6.6 shows $V_{nt}$ and HXR flux for two flares in the data set. These two examples show that there is a possible relationship between the decay time of $V_{nt}$ and the decay time of the HXR flux. The longer the HXR burst lasts after attaining maximum flux, the longer the enhanced levels of $V_{nt}$ are present. This relationship does not hold for all flares (e.g. Fig 6.5b) due to the effects of line narrowing. Line narrowing is an instrumental effect that artificially narrows the main resonance line in BCS data. This effect is most significant when the count rate is high (Chapter 3.3), however the $V_{nt}$ peaks when the count rate is low. Thus
Figure 6.4: Histograms of maximum HXR flux (a and e), maximum $V_{nt}$ (b and f), delay time (e and g) and rise time (d and h) for impulsive rise flares (a,b,c, and d) and gradual rise flares (e,f,g and h). The point at $t_{delay} = -1200$ is not plotted on the graph for gradual rise flares for reasons of clarity.

Line narrowing may affect the magnitude of the measured $V_{nt}$ however, it is unlikely to influence the time of the $V_{nt}$ peak.

6.4 Discussion

The main observational results from this study are listed below.

a) i) There exists a relationship between the HXR rise time ($t_{rise}$) and the time delay between the $V_{nt}$ and HXR peak ($t_{delay}$).

   ii) $t_{delay}$ is also related to the number of subsidiary HXR peaks that occur before the main HXR peak.

b) Approximately 20% of flares studied showed secondary peaks in $V_{nt}$ that were
always associated with HXR peaks. Secondary peaks occur after the HXR peak.

c) A possible relationship exists between the decay times of $V_{nt}$ and HXR flux.

d) There are systematic differences in the $V_{nt}$ behaviour in impulsive rise and gradual rise flares. Gradual rise flares show a tendency for $V_{nt}$ to peak before the HXR peak, whereas the opposite behaviour is observed in impulsive rise flares.

The HXR flux is a proxy for the deposition of energy in the form of energetic electrons. $V_{nt}$ a measure of the turbulence in the plasma and therefore stored energy. Therefore a relationship between HXR flux and $V_{nt}$ implies that the energy deposition by energetic electrons is a cause or effect of the plasma turbulence. The relationship between $t_{rise}$ and $t_{delay}$ supports the argument for a link between the HXR flux and $V_{nt}$ because it implies that the HXR flux before the HXR peak is influential in determining the time of the $V_{nt}$ peak. In addition to this general
CHAPTER 6. HXR AND $V_{nt}$ RELATIVE TIMINGS

Figure 6.6: Graphs showing $V_{nt}$ (with error bars) and HXR flux. These are two examples that illustrate a possible relationship between the decay time of $V_{nt}$ and the decay time of HXR flux. Note the time scales on each plot are not the same.

trend there are examples of direct responses of $V_{nt}$ to a large HXR pulse, and flares that suggest that the longer HXR flux persists the longer elevated levels of $V_{nt}$ are observed. These examples also support the existence of a causal relationship between the HXR flux and $V_{nt}$.

I stated in the introduction that the two mechanisms for $V_{nt}$ generation, i.e. turbulent evaporation flow and loop top turbulence associated with the initial energy release, could be separated based on the timing relationships between $V_{nt}$ and HXR peaks. The turbulent evaporation model will produce $V_{nt}$ peaks that follow HXR peaks whereas the loop top turbulent model would produce $V_{nt}$ peaks that precede HXR peaks. The differences in $V_{nt}$ behaviour in gradual rise and impulsive rise flares may suggest that different mechanisms for $V_{nt}$ generation are dominant in the different classes. Turbulence at the loop top associated with the initial energy release (e.g. Tsuneta 1994, 1995) appears the dominant mechanism for $V_{nt}$ generation in gradual rise flares because in these flares there is an overall tendency for the $V_{nt}$ to peak before the HXR peak. Whereas for impulsive rise flares, the
$V_{nt}$ generally peaks after the HXR burst, hence the turbulent evaporation model appears more appropriate.

If a turbulent evaporation flow is the predominant mechanism in impulsive rise flares then this may have some interesting implications. In Figure 6.6b the $V_{nt}$ is already high before the start of the measured HXR burst. There could be several reasons for this including, (a) the HXR flux was present but below the threshold level of detection of BATSE, (b) the electron population causing the evaporation had a high energy cutoff less than the low energy limit of the BATSE channel (25keV).

An alternative explanation for the different behaviour in gradual rise and impulsive rise flares, that may also explain the relationship between $t_{\text{rise}}$ and $t_{\text{delay}}$, may result from the inhomogeneity of the flare plasma. When calculating plasma parameters from the BCS spectrum we must assume the plasma is isothermal. However in both models there is a turbulent component (i.e. the evaporating plasma or the loop top plasma associated with the energy release) and a thermal component (i.e. stationary plasma in the flare loops), which may have different temperatures.

When the measured $V_{nt}$ becomes dominated by emission from the thermal component it will begin to decrease, forming an observed peak. The time at which this occurs will depend on the evolution of the flare. For both the evaporation model and the loop top model the amount of turbulent plasma should correlate with the HXR flux with a positive or negative time lag. However the thermal component increases continually up to at least the end of the HXR burst. For flares that evolve slowly, i.e those with a long HXR rise time ($t_{\text{rise}}$), the thermal component can become stronger than the turbulent component early in the flare, hence the time delay between the $V_{nt}$ and HXR peak ($t_{\text{delay}}$) is also large. This effect can therefore describe the observed relationship between $t_{\text{rise}}$ and $t_{\text{delay}}$.

In summary, the different behaviour of $V_{nt}$ in impulsive rise flare and gradual rise flares might imply that different mechanisms of $V_{nt}$ generation are dominant or that the differences may arise as a result of a multi-thermal flare plasma and
the corresponding difficulties of calculating the true value of $V_{nt}$. With spatially resolved line profile observations, although we could not completely separate the evaporating plasma from stationary plasma or loop top turbulent plasma we may be able to define regions where we expect either component to dominate. By comparing the temporal evolution of $V_{nt}$ in these different areas to see if the peak times vary, we will be able to distinguish between non-isothermal effects and the presence of two $V_{nt}$ generation mechanisms. In the simplest case if the peak times do vary this would indicate that non-isothermal effects are strong. Spatially resolved line profile measurements will be available from the EUV imaging spectrometer (EIS) one of a suite of instruments to be launched on the Solar-B satellite in 2005.
Chapter 7

Conclusions & Future Research

This thesis has examined two facets of solar flares: homology and non-thermal SXR line broadening. Although these two topics are not directly related, individually they are both important parts of the solar flare phenomenon. In this chapter I present the conclusions deduced from the work I have undertaken on each topic along with suggestions for future research.

7.1 Solar Flare Homology

7.1.1 Summary

In Chapter 3 I studied a series of three flares that occurred on 17th September 1997. The first and last flare in this series were shown to be homologous. A multi-wavelength study of the active region showed that the apparent homology was produced by the emergence of new magnetic flux. The first flare of the homologous pair occurred in a quadrupolar magnetic configuration. The subsequent emergence of new flux at a chance location and the second flare in the series, regenerated similar pre-flare conditions to that which existed before the first flare. The third flare therefore occurred in a similar magnetic configuration and thus appeared homologous to the first. It was therefore the emergence of magnetic flux that led
to the apparent homology and not the continual shearing of a single or group of magnetic structures.

This type of homology has not previously been observed and differs considerably from previous models that explain homology by the continual shearing of a single or group of magnetic structures. Homologous flares that occur as a result of continual shearing can provide information on such aspects as the flare trigger mechanism and energy storage time-scales, whereas this information is lost for flares that result from a multi-step process as described above.

The implications for studies of homologous flares are therefore that if this type of model is prevalent then information on the time-scales of energy build up cannot be extracted. However, homologous flares of this type provide information on the magnetic evolution of the active region.

7.1.2 Future Work

With the continuing advancements in instrument technology we are able to observe the Sun at increasingly high spatial resolution, spectral range and resolution and sensitivity. Such improvements allow the identification of fine structure in solar flares. These revelations have shown us that at progressively smaller spatial scales the differences between flares become increasingly more apparent. As a result the number of homologous flare sightings is decreasing.

However, high resolution multi-wavelength observations of a potential flare site with good temporal resolution and continuity, allow us to observe the many facets of activity at all levels of the solar atmosphere prior to the onset of a flare. Presently the TRACE satellite obtains active region images at unprecedented spatial resolution and excellent temporal continuity. Combined studies of flares with Yohkoh and TRACE will help provide information on the trigger mechanism for individual flares, which is the key to understanding the energy release process in solar flares.
The next generation of solar instrumentation will further enhance the quality of observational data. The Solar Terrestrial Relations Observatory (STEREO) mission scheduled for launch in 2004 consists of two identical spacecraft. Both spacecraft will orbit at 1AU, one will precede Earth’s orbit by 60°, the other will follow 60° behind. Combined with imaging from spacecraft at earth orbit this will provide an excellent 3-Dimensional view of solar phenomena. 3D high spatial resolution observations of flaring active regions, with excellent temporal continuity, would increase our knowledge of the environments in which flares occur. Chromospheric observations and photospheric magnetograms may show footpoint motions that produce shear. Coupled to coronal observations of the loops may lead to the identification of individual events and circumstances that trigger the flare, for example the determination of a possible value of critical shear, loop interaction distances or emerging flux heights.

7.2 X-ray Non-Thermal Line Broadening

7.2.1 Summary

In Chapters 4 and 5 I have performed two studies on the characteristics of non-thermal X-ray emission line broadening in solar flares. The combined results of these two studies are that:

a) The source of $V_{nt}$ is located above or within the flare loops, not at the flare footpoints.

b) The HXR flux and $V_{nt}$ are causally related.

c) The timing of $V_{nt}$ maxima are dependent upon the form of the HXR burst. In impulsive rise flares the $V_{nt}$ peak generally occurs after the main HXR peak, whereas in gradual rise flares the main $V_{nt}$ peak has a tendency to occur before the main HXR peak.
The findings on the source location of $V_{nt}$ have shown that the mechanism for their generation is either turbulence associated with chromospheric evaporation or turbulence in plasma above the loop top that is associated with the initial energy release. The finding that $V_{nt}$ and HXR flux are causally related supports this conclusion.

The result of different $V_{nt}$ peak timings in gradual and impulsive rise flares has potentially some very interesting implications. It suggests that different mechanisms for $V_{nt}$ generation may be prevalent in the different flare classes. However, it was also shown that the timings may be affected by the inhomogeneous nature of the flare plasma.

### 7.2.2 Future Work

A logical extension to the study of $V_{nt}$ source location would be to perform a similar analysis to that carried out in Chapter 4 on several solar limb flares. Limb flares are free from the line of sight effects that could not be accounted for in the analysis of the 22$^{nd}$ June flare. This may help determine if the source of $V_{nt}$ is above or within the flare loops. Simultaneous analysis of BCS S x v , Ca x ix and Fe xxv of flares occurring in the only active region on the disk, will show how the location of the sources varies as a function of temperature. The need for the flare to occur in the only active region on the disk, or at the very least the largest by a significant amount, is to obtain spectra in the BCS S x v channel that are free from contaminating sources.

Chapter 6 showed how the relative timing of the $V_{nt}$ peak was different in impulsive and gradual rise values; classified based on the HXR time profiles. It would be productive to determine if these two types of flares also have systematically different SXR morphology or reconnection scenarios that may account for the different behaviour of $V_{nt}$.

By studying the complete temporal evolution of $V_{nt}$ during a solar flare and
comparing this with the HXR lightcurve can also provide information on the nature of the $V_{nt}$ source. This was first accomplished recently by Harra et al. (2001) for a single flare which showed that the $V_{nt}$ begins to rise before the HXR burst begins. This behaviour is more indicative of plasma turbulence associated with reconnection than with hydrodynamic flows. It is necessary to apply similar analysis to a wide variety of flares, including gradual and impulsive rise flares to find out if this $V_{nt}$ behaviour is common for all flare types.

The High Energy Solar Spectroscopic Imager (HESSI) due for launch in March 2001 will provide spatially resolved images from $3keV \rightarrow 20MeV$ with greater sensitivity, temporal, spatial and spectral resolution than presently available. HESSI's imaging capabilities over such a large spectral range will allow the determination of the location of hot plasma in the flare loops. A comparison of emission measures between BCS and HESSI, similar to those carried out in Chapter 4, may help to locate to a higher level of certainty the location of the source of the non-thermal broadenings.

The improved sensitivity of HESSI will allow the detection of HXRIs at earlier times in a flare. By combining these improved HXR lightcurves with BCS data on the rise of $V_{nt}$, we may be able to determine whether the start of the HXR burst signifies the beginning of the increase in $V_{nt}$.

The current instrumentation to study SXR line profiles (i.e. BCS) lacks spatial resolution, hence present techniques to infer the location must use indirect methods. To unambiguously determine the location requires spatial resolution of the line profiles. Spatially resolved line profile observations will be available with the launch of the EUV Imaging Spectrometer (EIS) on the Japanese Solar-B mission in 2005. Spatially resolved observations of line profiles in a solar flare will provide a leap forward in the understanding of $V_{nt}$. The variation of $V_{nt}$ with location in the loop might allow the distinction between the generation of $V_{nt}$ as a consequence of the energy release or from turbulence in evaporation up-flows. The former may possibly result in the $V_{nt}$ being very much localized around the loop top, whereas
turbulence in chromospheric up-flows may be more conducive to a more uniform distribution of $V_{nt}$.

Spatially resolved line profiles may also help to determine whether or not the $V_{nt}$ timing is influenced by the non-isothermal nature of the flare. By taking resolved observations it may be possible to separate the regions of significantly different temperatures and it will be possible to compare the $V_{nt}$ evolution at different parts of the flare to see if the peaks all occur co-temporally.
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