X–ray and Radio Diagnostics of Accelerated Electrons in Solar Active Regions

A dissertation submitted to the University of Dublin for the degree of Doctor of Philosophy

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Trinity College Dublin, March 2015
Declaration

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Summary

Solar eruptive events occur when magnetic energy stored in the atmosphere of the Sun is released as a mixture of rapidly ejected plasma (coronal mass ejections) and radiation across the electromagnetic spectrum (solar flares). Energy released in this way can reach as high as $10^{32}$ ergs, making these events the most powerful explosions in the solar system. Despite centuries of study since their first observation in 1859, understanding of the exact nature of the initiation of these events is still incomplete. In this thesis, I outline work which was performed in order to build upon the understanding of the initiation of solar flares, primarily through analysis of nonthermal X–ray and radio emissions. In the first two parts of this work, open questions on the motion and size of nonthermal X–ray sources in solar flare loops are addressed, and in the final part a unique observation of magnetic reconnection in the high corona and the associated radio emissions are presented and discussed.

The nonthermal distribution of accelerated electrons believed to produce solar hard X–ray emissions is known to evolve dramatically over the course of a solar flare, from a steep to a flat power–law spectrum as the acceleration process becomes more efficient. In the context of the commonly–used collisional thick–target model, this should produce nonthermal X–ray sources which move down the flare loop as the spectrum steepens. However, this has never been observed, constituting a problem with the standard model. In the first part of this thesis, the analysis of an early impulsive event, which exhibits low–energy nonthermal emission early in the flare, is presented. It is found that this downward motion is indeed present, and through modelling
is shown to be consistent with the flattening of the injected electron spectrum, revealed through X–ray spectroscopy.

The second piece of work outlined in this thesis includes a potential solution to another outstanding problem in X–ray solar physics, namely of extended X–ray source sizes. Previously, X–ray sources were found to be far larger than the sizes predicted using the collisional thick–target model. This problem is addressed by incorporating a unique model of the ionisation fraction of the plasma encountered by the accelerated electron distribution in a model flare. We include a local peak ionisation, which is expected to be produced by the beam itself. This locally ionised and heated chromosphere successfully produces a vertically extended X–ray source, which accounts for a large portion of the difference between observed and previously modelled X–ray emission.

The final part of this thesis outlines a new observation of magnetic reconnection in progress in an active region which also exhibits a radio noise storm. The collapse of the observed X–point includes inflows and outflows of 1–5 km/s and 30–100 km/s, respectively, implying a reconnection rate of 0.05, consistent with previous observations of reconnection on the solar limb. We demonstrate that the magnetic geometry suggested by extreme ultra–violet imaging and potential magnetic field extrapolations suggests the presence of a 3D separator connecting magnetic null points above the active region. In this context, the radio noise storm evolution during the collapse is then explained by the concept of long–term gradual acceleration followed by flare–related reconnection, which produces a rapid brightening of the radio emission.
For Mam and Dad.
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List of Publications

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7.9 Schematic diagram of the Swedish Solar Tower. The ~1 m primary fused silica lens is located at the top of the tower in order to avoid the refraction caused by convective air flows near the ground. The lens is mounted on a turret which tracks the Sun throughout the day. The light then reflects off the two shown 1.4 m flat mirrors to be passed to the mirror setup labelled as c, which passes the light to an optical bench which includes the desired beam splitters and cameras. a and b show the Schupmann corrector system, which is intended to cancel out the effects of the 1 m lens. (Scharmer et al., 2003).
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Introduction

Solar flares are the most powerful explosions known to occur in the solar system, regularly releasing as much as $10^{32}$ ergs of energy in the form of intense radiation at all wavelengths from low–frequency radio waves to gamma rays. Despite occurring up to many tens of times a day, the varied and complex physical processes which are involved in the phenomenon of solar flares are still being actively studied to this day. In particular, nonthermally accelerated particles are believed to play a key role in flare initiation, and yet there remains no clear consensus on how they propagate through the solar atmosphere. An important diagnostic for these accelerated particles is the analysis of hard X-ray and plasma radio emission. In this chapter, we place observations of the active Sun in these wavelengths, and others, in a historical context, and the general understanding of the physics driving them is outlined.
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Figure 1.1: A hand–drawn image of the first and one of the largest solar flares ever observed. This event was detected by Richard Carrington during a routine observation of a large sunspot group (shown as black umbrae and shaded penum-brae). The bright ribbons are indicated by the outlined white areas within the group, located at A and B upon first observation, and C and D after 5 minutes (Carrington [1859]).

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The first solar flare to be imaged was the Carrington event, which occurred on 1 September, 1859 ([Carrington 1859], [Hodgson 1859]). Sketched by Carrington himself as shown in Figure 1.1, this event was referred to as “exceedingly rare”. Indeed, it has since been learned that this event was unique in that it was one of the most powerful solar flares ever recorded, and the most geo–effective. The ejected material related to this event is believed to have impacted the Earth, causing aurorae as far south as Cuba, and generated currents in telegraph wires, in
1.1 Solar Flares

Figure 1.2: Early observations of X-ray solar flares, produced by the Hard X-ray Imaging Spectrometer (HXIS) onboard the Solar Maximum Mission (SMM). Shown are three flares (a–c), with contours of soft (3.5–5.5 keV) X-ray images on the top row, and hard (16–30 keV) X-ray images on the bottom row. These images reveal the footpoint–looptop spatial structure that still describes many modern flare observations. The upper images show hot, thermally–emitting plasma at the apparent flare looptop, while the bottom HXR images show pairs of nonthermal footpoints beneath. The dashed line in the April and May events correspond to the magnetic neutral line, and in the November event correspond to the location of the filament which was present before the impulsive phase of the flare (Duijveman et al., 1982).

In some cases causing fires, and allowing telegraphs to be used without power (Committee On The Societal & Economic Impacts Of Severe Space Weather Events, 2008). In addition, the event produced a massive amount of highly energetic solar protons at Earth, a number which has not been exceeded since (Shea et al.).
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This event sparked 150 years of investigation into these powerful, apparently rare solar storms. Ground–based observations continued to provide high–detail images of visible–light sunspots and flares, but in order to probe the high–energy aspects of common solar flares, numerous balloon– and space–based X–ray imagers and spectrometers were developed. Solar hard X–rays (HXRs) were first recorded by a balloon flight operated by Peterson & Winckler (1959). The first space–based X–ray instruments dedicated to solar flares were launched onboard the Solar Maximum Mission (SMM; Strong & Schmelz, 1999) on 4 February 1980. The HXR Burst Spectrometer (HXRBS Orwig et al., 1980), Gamma–ray Spectrometer (GRS Forrest et al., 1980), and the HXR Imaging Spectrometer (HXIS van Beek et al., 1980) onboard the SMM were designed to observe solar HXRs at energy ranges of 20–255 keV, 300–9000 keV and 3.5–30 keV, respectively. These combined spatial and spectral observations revealed the characteristic ‘footpoint–looptop’ structure of solar flares (see for example Figure 1.2, also observed by Ohki et al. (1983)), and important evidence for the nonthermal acceleration of particles (Brown & Loran, 1985).

The HXIS was later supplemented by the Soft X–ray Telescope (SXT: Tsuneta et al., 1991) and the Hard X–ray Telescope onboard the Japanese YOHKOH spacecraft, known as Solar–A prior to launch. The SXT was used in the study important aspects of solar flares, such as soft X–ray (SXR) jets (Shibata et al., 1992), however the HXT was used to uncover a key in understanding solar flare acceleration processes – an ‘above–the–looptop’ HXR source (Masuda et al., 1994, shown in Figure 1.3). This new type of source – thereafter determined to be quite common (Petrosian et al., 2002) – was interpreted to be nonthermal thin target
1.1 Solar Flares

Figure 1.3: The now–famous observation of an ‘above–the–looptop’ HXR source, observed by S. Masuda on 13 January 1992 using the YOHKOH HXT. The upper panels show contours of HXR emission for increasing energy from left to right. The soft X–rays show the usual loop–shaped structure, while a clear HXR source is visible above the looptop. The bottom images are produced using the SXT, demonstrating there is no apparent thermal emission present to explain the source. This source was therefore interpreted as nonthermal emission originating from the acceleration site itself (Masuda et al., 1994).

emission; HXRs produced by an accelerated population of electrons immediately interacting with their surrounding plasma. This event along with other observations of plasmoid motions were combined to form a comprehensive model of solar flares, relying on theories of a process known as magnetic reconnection (Shibata, 1996; see Section 2.1.5 for a detailed discussion of this process).
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Figure 1.4: Exploded view of the Ramaty High Energy Spectral Solar Imager (RHESSI) instrument. Shown are the dual-grid rotation–modulated collimator system as well as the hyperpure germanium detectors. RHESSI is described in detail in Section 3.1.

As we move to the modern era of solar HXR observations, we find it is dominated by a single instrument: the Ramaty High–Energy Solar Spectroscopic Imager (RHESSI, Lin et al. 2002). RHESSI, shown in Figure 1.4 and described in much more detail in Section 3.1, was launched in February 2002, and is capable of producing ~3 arcsecond (~2 Mm) resolution images of solar HXRs from ~3 keV to 17 MeV, at a uniquely high spectral resolution of ~1 keV for energies below 100 keV, increasing to 5 keV at 5 MeV. Given these strengths, RHESSI has contributed immensely to the understanding of the high–energy aspects of solar flares, which is outlined in the remainder of this Section.

These fifteen decades of solar observing have provided a wealth of information on the overall behaviour of solar flares. For example, as shown in Figure 1.5, the amount of energy released in solar flares covers over nine orders of magnitude. These data were produced from various studies dating as far back as 1980, us-
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Figure 1.5: A statistical overview of solar flare energy emission ranging from nanoflares on the left through microflares to large flares on the right. Each colour of plot corresponds to data gathered by a different solar instrument over a different duration, listed in the legend. As shown, studies frequently reveal a power law in the distribution of flare frequency with emitted energy (Hannah et al., 2011).

![Diagram of solar flare energy distribution](image)

TRA: C. [Hannah et al., 2011]
EIT: 1 hr (Jul-96) [B2002]
TRACE: 1 hr (Feb-98) [A2000]
SXT: 5 dys (Aug-92) [S1995]
RHESSI: 5 yrs (2002-2007) [H2008]

Figure 1.5: A statistical overview of solar flare energy emission ranging from nanoflares on the left through microflares to large flares on the right. Each colour of plot corresponds to data gathered by a different solar instrument over a different duration, listed in the legend. As shown, studies frequently reveal a power law in the distribution of flare frequency with emitted energy (Hannah et al., 2011).

ing many of the instruments outlined above (Aschwanden et al., 2000; Benz & Krucker, 2002; Crosby et al., 1993; Hannah et al., 2008; Lin et al., 2001; Parnell & Jupp, 2000; Shimizu, 1995). Studies such as these, reviewed in Hannah et al. (2011), demonstrate that the distribution of flare frequency with size tends to follow a power–law form (Akabane, 1956), suggesting the presence of a cascading mechanism of energy release in solar flares. Statistical studies have also revealed a number of general behaviours in solar flares, such as adherence to the Neupert effect, which states that the HXR flux is proportional to the temporal derivative of the SXR flux (Neupert, 1968).
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Table 1.1: Energy partition in two solar eruptive events. [Emslie et al., 2005]

Additionally, several studies have been performed with the goal of determining the energy partitioning of solar eruptive events (SEEs). SEEs are generally defined to include the solar flare, which constitutes the thermal and nonthermal electromagnetic radiation produced by an event, as well as the coronal mass ejection (CME) released. The resulting energy budget for two SEEs was produced by [Emslie et al., 2004, 2005], and is shown in Table 1.1. These results showed that the CME energy in both cases constitutes the dominant part of the energy budget, far more than the accelerated particles, which has placed constraints on flare models developed since.

1.1.1 Morphology

Solar flares occur in areas of the solar atmosphere characterised by a substantially increased magnetic field strength, known as active regions, however there are exceptions [Dodson & Hedeman, 1970; Martin, 1980]. This immediately highlights the significance of the solar magnetic field in all solar flare models, and provides a
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Figure 1.6: Active regions on the Sun. Shown to the left is a line–of–sight HMI magnetogram, and to the right an AIA 193 Å image, produced on 31 July 2012 (See Section 3.2 for a description of these instruments). As shown, pairs of strong regions of oppositely–aligned magnetic field in the magnetogram correspond to hot coronal loop–shaped structures in the EUV image.

backdrop for the spatial structure of a ‘standard’ solar flare. These active regions are commonly associated with bright extreme ultra–violet (EUV) loop structures (such as those shown in the full–disk image in Figure 1.6). This emission is created by de–excitation of highly ionised ∼0.1–10 MK coronal plasma, which is described as ‘frozen in’ to the active region magnetic field (see Section 2.1.3 for detail). As such, EUV emission is a useful starting point in the discussion of the spatial structure of active regions and solar flares.

Along with visible light, EUV emission constitutes the majority of radiated energy in a solar flare (Woods et al. 2006). EUV images have been produced at high resolution with many instruments such as the EUV Imaging Telescope (EIT; Delaboudinière et al. 1995) onboard the Solar and Heliospheric Observatory (SOHO), the Transition Region and Coronal Explorer (Handy et al. 1999)
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Figure 1.7: A reconnecting solar coronal loop, as seen in AIA 131 Å. This event occurred on the eastern solar limb, and exhibited inflows and outflows which were interpreted as an observation of magnetic reconnection (Su et al., 2013).

TRACE;), and the Atmospheric Imaging Assembly (AIA; Lemen et al., 2012) onboard the Solar Dynamics Observatory. Here we outline two characteristic structures in terms of solar flare plasma. Firstly, coronal loops are a common observation with these instruments, and are generally interpreted to trace out the magnetic field lines which connect positive and negative magnetic poles in the solar photosphere (e.g., Gosain, 2012). The flows exhibited within and between these coronal loops are often presented as evidence for magnetic reconnection when direct observations of nonthermal emission are unavailable (Narukage & Shibata, 2006; Savage et al., 2012a; Yokoyama et al., 2001). One such example is shown in Figure 1.7. In this event, an apparent X-point was produced by a hot coronal loop on the eastern limb, exhibiting clear inflow and outflow as predicted by 2D magnetic reconnection models (Su et al., 2013).

Another common observation seen in EUV is the coronal arcade. Arcades can be seen as the 3D extension of the limb–observed coronal loop, as they are essentially hot, thermally emitting plasma along the multiple sets of field lines which connect two or more regions of opposite magnetic polarity. This equivalence
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Figure 1.8: Two views of a solar flare arcade. Top: AIA 193 Å images, taken from the point of view of the Earth appear to be on the western solar limb. The loop is shown to expand with time (from left to right) as the event progresses. Bottom: STEREO 195 Å images at the same time intervals, from a location 121° counter-clockwise, as shown in the inset of the first image. As shown, from above the active region, the loop structure is revealed to be extended, forming an arcade (Liu et al., 2013).

is demonstrated in Figure 1.8 which shows a coronal ‘loop’ observed from the Earth’s point of view, with its 3D nature as an arcade revealed by the Extreme Ultraviolet Imager (EUVI; Wuelser et al., 2004) onboard the Solar–TERrestrial RElations Observatory (STEREO), which was located above the active region at the time (Liu et al., 2013). Arcade formation is strongly linked with the production of flux ropes, which are thought to be the originating structures of CMEs (Fan & Gibson, 2007; Roussev et al., 2003; Webb, 1992).

A related visible and ultra–violet (UV) phenomenon which is strongly associated with loops and arcades is that of flare ribbons. Ribbons are commonly identifiable as brightenings in broadband white light and the 6563 Å line emis-
sion produced by the 3 to 2 Balmer transition in neutral hydrogen, known as Hα emission. They generally appear as roughly parallel brightenings, usually separated by the polarity inversion line of the magnetic field (e.g., Asai et al., 2003; Zirin & Tanaka, 1973). Importantly, ribbons are consistently associated with HXR footpoints (see below for details), which indicates that the emission is produced by chromospheric plasma which has been nonthermally heated by the accelerated electrons producing the HXRs (Canfield & Gayley, 1987; Qiu et al., 2001). Thus, ribbon formation and evolution played an important role in the development of the standard model of solar flares.

1.1.1.1 X–Ray Structure

X–ray images provide detailed information on both the thermal and nonthermal processes which occur in a solar flare. As originally shown in Figure 1.2, X–ray images often exhibit a structure well–described by a pair of high–energy footpoints, and a larger extended loop structure in the softer X–rays. Notably, SXRs often originate exclusively from higher up in the loop, as this is where heated plasma gathers in the gradual phase of the flare (see Section 1.1.2.3). Examples of this overall structure remain common in the RHESSI era. However, the improved capabilities of RHESSI have lead to further discovery, for example that of the presence of a ‘super–hot’ 44 MK distribution of plasma located slightly above the standard flare looptop in one observed event (Caspi & Lin, 2010).

Separate from thermal loop emission, nonthermal looptop and ‘above–the–looptop’ sources have been of great interest since their discovery by Masuda et al. (1994), as they can potentially reveal details on the location and nature of the acceleration region itself. Indeed, RHESSI and EUV studies have shown that at
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**Figure 1.9:** RHESSI nonthermal loop HXR images for ten solar flare events. Lower energy emission is shown at the top, and higher at the bottom. For each energy, the top row is produced by CLEAN, and the bottom by VIS FWDFIT (see Section 3.1.1 for details on these imaging methods). As shown, loop length appears to increase with increasing photon energy (Xu et al. 2008).

these Masuda sources, all electrons are energised, suggesting a direct observation of the acceleration region (Krucker & Battaglia 2013; Krucker et al. 2010). In some circumstances, loop emission can also be produced nonthermally, and it has been shown by numerous studies that more extended loop-like sources appear for higher photon energies (see Figure 1.9), consistent with a beam of particles penetrating downwards from above, with higher energy particles reaching further along the loop before braking (Guo et al. 2012; Jeffrey & Kontar 2013; Xu et al. 2008). Motion of these sources during a flare have also been put forward as evidence for reconnection between the looptop and a previously disconnected plasmoid (Bárta et al. 2008; Milligan et al. 2010).

A great focus of RHESSI studies has been on the spatial location, structure and size of nonthermal footpoints. A number of studies have been performed to identify source centroid locations, with the goal of gaining an understanding on
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how accelerated electrons deposit their energy in the corona and chromosphere (Aschwanden *et al.* 2002; Battaglia *et al.* 2012). In terms of source size, advanced image reconstruction techniques were used to reveal that HXR footpoints are 3–6 times larger than predicted, based on the collisional thick target model, leading to the suggestion of a threaded loop structure rather than a single monolithic loop (Dennis & Pernak, 2009; Kontar *et al.* 2010). This issue is a focus of the analysis outlined in Chapter 5.

1.1.1.2 Structure in Other Wavelengths

HXR footpoints are frequently associated with the H$\alpha$ and UV ribbons mentioned above (e.g., Tanaka & Zirin 1985; Temmer *et al.* 2007). This supports the interpretation that ribbons are produced at the footpoints of flare arcades, where the accelerated electrons causing the nonthermal heating are emitting bremsstrahlung in HXRs. However, a consistent distinction between HXR footpoints and visible/UV ribbons is that HXR sources are often much more localised, usually confined to points along ribbons. It has been suggested by Asai *et al.* (2002) that the HXRs occur at regions of the flare arcade which exhibit locally increased magnetic field strength. However, a small number of events have occurred in which HXRs exhibit a ribbon–like structure (Masuda *et al.* 2001). An example of such an event is shown in Figure 1.10. In this observation by Liu *et al.* (2007), it is hypothesized that these counterexamples occur when a uniformity of energy deposition occurs along a ribbon, perhaps due to a uniquely clear transition from sheared loops to the standard post–flare arcade.

Microwave emission is also known to be associated with impulsive HXR emission as far back as the early observations of Peterson & Winckler (1959). These
microwaves are interpreted as gyrosynchrotron and synchrotron emission produced by nonthermally accelerated electrons spiraling along the strong magnetic fields in solar flare loops (Kundu, 1961). Indeed, microwave imaging has revealed a strong association in spatial structure with SXR emission, as demonstrated for example in 1.11 (see also e.g., Kundu et al., 2004, 2009). However, a large number of studies have shown that the spectral index of electrons producing the microwave emission does not match that producing the HXRs, suggesting a
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Figure 1.11: Comparison of RHESSI 10–20 keV images (black contours) and 34 GHz images produced by the Nobayama Radioheliograph (NoRH; Nakajima et al., 1994), shown as grey contours. As shown, the NORH microwave emission originates from a pair of footpoints and occasionally a more extended loop structure, while the RHESSI SXRs trace out the overlying loop of heated plasma (Sui et al., 2005b).

break in the power law between the HXR energies and \(\sim 300\) keV, above which microwaves are produced (Kundu et al., 1994; Silva et al., 2000).

In extremely energetic events, some flares have exhibited the impulsive emission of gamma radiation. Gamma rays originating from the Sun were first observed as an enhancement over background flux at all energies below 7 MeV, with clear lines at 0.5 MeV, 2.2 MeV, 4.4 MeV, and 6.1 MeV (Chupp et al., 1973).
1.1 Solar Flares

Figure 1.12: Three gamma–ray flares observed by RHESSI. Trace images are shown for (from left to right) flares occurring on October 28, October 29, and November 2 of 2002. Overlaid in red and blue are the contours of the 200–300 keV (HXR) and 2218–2228 keV (gamma ray) RHESSI images. In the middle image, the blue circle is centred on the centroid location with a radius of $1\sigma$ of this position. The purple lines trace the motion of the HXR footpoints (Hurford et al., 2006).

Suri et al. (1975). Line emission at these energies is interpreted as emission from chromospheric atoms which have been excited by accelerated ions and protons during a solar flare (Kuzhevskii, 1969; Lingenfelter & Ramaty, 1967). A series of gamma–ray emitting events were imaged by RHESSI and TRACE, as shown in Figure 1.12. The 2.2 MeV emission used to produce the images is dominated by the neutron capture line, produced when neutrons created in flare ion collisions are captured on $^3$He nuclei (see e.g., Biermann et al., 1951; Hua & Lingenfelter, 1987). These rare observations of impulsive gamma–ray emission normally reveal single unresolved sources, such as in the latter two events, which have indicated that these bursts are produced in flares. However in the first event of October 28 2002, two ‘footpoint–like’ sources appeared about 15 arcsec offset from the HXR footpoints, which may be explained partly by gradient and curvature drifts, which will force the oppositely–charged electrons and ions in opposing directions.
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1.1.2 Temporal Evolution

As we have now established the standard spatial structure of solar flare emissions, we will now outline a detailed description of solar flare emission as the event evolves over time. Below we detail the behaviour of a standard solar flare in terms of morphological and spectral evolution by dividing the flare progression into three phases; the preflare phase, the impulsive phase, and the gradual phase.

1.1.2.1 The Preflare Phase

In the minutes to hours leading up to impulsive flare events, it has been noted by many authors that a gradual brightening appears in SXR, EUV and UV wavelengths (Acton et al. 1992; Bumba & Křivský 1959; Fárnák et al. 1996). This brightening is commonly attributed to some sort of pre–heating of the chromospheric plasma up to temperatures as high as 10 MK, despite the lack of particle acceleration at the time (e.g., Fletcher et al. 2013). This requires that an additional model for energy transfer from corona to chromosphere in the pre–flaring phase be developed. For example, one study has suggested that this heating is caused by a conduction front, originating in the corona, which travels to the chromosphere, heating and expanding plasma there (Battaglia et al. 2009).

1.1.2.2 The Impulsive Phase

The impulsive phase is characterised by a rapid increase in the nonthermal emissions outlined above, namely gamma–rays, HXRs, white light continuum, and nonthermal radio and microwave emissions. During this time, most of the stored magnetic energy is converted to power the solar flare and eventual eruption, de-
1.1 Solar Flares

Figure 1.13: Flare HXR footpoint evolution during the impulsive phase. *Left:* 20–50 keV RHESSI HXR centroids produced by CLEAN (crosses) and PIXON (circles; Metcalf et al. 1996) overlaid on a EIT 195 Å image of a flare ribbon. The time each source is measured at is indicated by the symbols colour, and the polarity inversion line is overplotted in gray. HXR sources appear at the northern edge of the flare ribbons and evolve southward between 13:12 and 13:24 UT. *Right:* Evolution of photon spectral index and flux at 35 keV over the impulsive phase (Grigis & Benz 2005).

spite lasting on the order of minutes. Studies which make use of magnetic field extrapolations from the photosphere have supported this idea, showing that active regions with a large amount of measured free energy tend to produce more flares (Jing et al. 2010). Such studies have also shown a drop in magnetic energy and magnetic helicity during a solar flare (Bleybel et al. 2002, Murray et al. 2013), again indicating that this energy is released to power the flare.

It is during the impulsive phase that the HXR and Hα footpoints discussed above appear and evolve. From the standard model of solar flares it is expected that new magnetic field lines are brought into the reconnection region as a flare progresses, resulting in motion of these footpoints (Forbes & Lin 2000).
2.1.5 for details). In fact, footpoints are regularly observed to separate over time, interpreted as the growth of the flaring loop system above \cite{Fletcher_Hudson_2002, Qiu_etal_2002}. It has also been shown that these HXR sources show a consistent motion along flare ribbons, as shown in Figure 1.13. Here, HXR sources appear to move clearly along the footpoints of the flare arcade, and so was interpreted as a disturbance which propagated smoothly along the arcade, triggering acceleration as it progressed \cite{Grigis_Benz_2005}.

Conversely, footpoint convergence is also a reasonably common observation, for example from 13:12 to 13:14 UT in Figure 1.13. Indeed there are several observations which show convergence followed by separation \cite{Ji_etal_2008, Zhou_etal_2008}. Such observations can also be accompanied by a HXR looptop source which exhibits a decrease in altitude during the footpoint convergence \cite{Liu_etal_2009}. These observations are partially explained by the concept of a coronal implosion, or inflow of plasma towards the region of energy release as magnetic pressure rapidly drops \cite{Hudson_2000, Simoes_etal_2013}.

Accelerated particles which produce these nonthermal signatures deposit their energy in the chromosphere, causing a process of heating and expansion of plasma known as chromospheric evaporation \cite{Antiochos_Sturrock_1978}. This process has been categorised as either gentle or explosive chromospheric evaporation. In the former case, the chromospheric plasma is heated by either the accelerated particles or a conduction front and so expands, while in the latter, the plasma is heated so rapidly that radiative cooling is insufficient, and plasma expands both upwards and downwards \cite{Fisher_etal_1984, Mariska_etal_1989}. With modern UV and HXR imaging spectrometers, both types of evaporation have now been directly observed, as shown for the case of explosive evaporation in Figure 1.14.
Figure 1.14: Evidence for explosive chromospheric evaporation. Shown are velocity maps produced by the Coronal Diagnostics Imager (CDS; Harrison et al., 1995) for He I and Fe XIX line emission in the top and bottom panels, respectively. Red and blue pixels denote down- and up-flowing material, respectively, with the velocity scale shown on the right. Overplotted are 10 % and 40 % contours of RHESSI 25–60 keV emission (Milligan et al., 2006a).
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[Milligan et al. 2006a,b]. In this work, upward and downward velocities on the order of $\sim 230$ and $\sim 30$ km/s were detected from blue- and red-shifted thermal plasma emission, respectively. More recently, evaporation has been detected indirectly through RHESSI HXR nonthermal emissions [Liu et al. 2006] and by loop-filling microwave emission [Gan & Li 2012].

Finally it is important to consider the spectral evolution during a solar flare. The RHESSI HXR spectrum is generally comprised of an exponential thermal component and a power-law nonthermal component (see Section 3.1.2 for further details). During the impulsive phase, the nonthermal component dominates the HXRs. Often it first appears relatively steep before flattening to reach the higher energies, then softens again as the flare transitions into the gradual phase (e.g., Benz 1977; Fletcher & Hudson 2002; Parks & Winckler 1969). This behaviour is commonly referred to as ‘soft–hard–soft’ evolution, and is apparent for example in Figure 1.13, where during the course of the event, a large number of impulsive HXR bursts occur, along with multiple instances of fall and rise of spectral index. This evolution is generally attributed to an increase and decrease in acceleration efficiency, which should also have an effect on HXR spatial structure. This concept is one of the science goals of this thesis, and is explained fully in Chapter 4.

1.1.2.3 The Gradual Phase

Following the rapid energy conversion and release associated with the impulsive phase, a great deal of emission still continues to take place for hours afterward, in a period of time known as the gradual phase of the flare. During this time, nonthermal HXR bursts and radio emission are generally no longer observed, and SXRs exhibit a gradual decay.
1.1 Solar Flares

Figure 1.15: Cooling in a canonical solar flare. *Top:* RHESSI 12–25 keV and 6–12 keV looptop images shortly after the peak in HXRs. *Middle:* Sequences of images produced using TRACE and SOHO/CDS, with responses to coronal temperatures decreasing from 8 MK (top) to 0.25 MK (bottom). Time during the flare progresses from left to right. *Bottom* RHESSI (left axis) and GOES (solid line, right axis) lightcurves, with noted times corresponding to the images above (Raftery *et al.*, 2009).
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The primary physical process taking place in the post-flare loop or arcade system is that of plasma cooling. A well-observed example of flare cooling is shown in Figure 1.15. As shown in the middle grid of images, hot plasma originally appearing in the hottest channels at $\sim 8$ MK disappears and simultaneously appears in the cooler channels, progressing down to $\sim 0.25$ MK, indicating the plasma in the post-flare loop is cooling (Raftery et al. 2009). During flare evolution, the dominant cooling mechanism can transition between thermal conduction (e.g., Culhane et al. 1994) and radiation (Cargill et al. 1995; Fisher & Hawley 1990). Conduction has been observed to originally dominate, followed by radiation as the flare cools (e.g., Aschwanden & Alexander 2001). Recent modelling work has however indicated a third phase of ‘catastrophic cooling’, characterised by a rapid drop in temperature and a roughly constant density (Cargill & Bradshaw, 2013; Reale & Landi 2012). Conversely, observational results have indicated a delay in cooling, interpreted as an additional post-flare heating mechanism (e.g., Jiang et al. 2006; Ryan et al. 2012).

During the gradual phase, the plasma at the top of the post-flare loops might exhibit a ‘cusp’-shaped structure (as previously seen at the limb in Figures 1.7 and 1.8). These cusps tend to be hotter at higher altitudes, and are commonly interpreted as pinched magnetic fields underneath a current sheet (Liu et al. 2009; Tsuneta et al. 1992). These loops can be seen to rise and expand outwards during the gradual phase (Gallagher et al. 2002; Svestka et al. 1987). This is commonly interpreted as motion and evolution of the acceleration region itself (Sui & Holman 2003; Veronig et al. 2006). This interpretation is supported by the presence of supra-arcade-downflows, interpreted as low-density contracting flux tubes (McKenzie & Hudson 1999; Sheeley et al. 2004) and, more recently
as wakes behind smaller retracting loops (Savage et al. 2012b).

1.2 Radio Noise Storms

Solar radio emissions have been enthusiastically studied since their first observation by Reber (1944). Various physical mechanisms result in radio emission in the solar corona (see Nindos et al. 2008 for review). Incoherent emission mechanisms include free–free thermal emission produced by heated flare plasma (e.g., Dulk & Gary, 1983; Kundu et al., 1982), and gyrosynchrotron and synchrotron emission produced by accelerated electrons bound to magnetic fields in active regions (e.g., Benz et al., 2005; Ramaty, 1969). Coherent radio emission can be produced by electron cyclotron masers (Ergun et al., 2000), and plasma emission (see Melrose 1987 for review). Plasma emission is an important tool in understanding the active Sun, as it is produced (indirectly) by nonthermal processes (see Section 2.3.2 for details), and so will be our focus here.

Plasma emission is generally categorised into a number of different ‘types’. As illustrated in Figure 1.16, these types are generally distinguishable based on their appearance in a dynamic spectrum (Dulk 1985). Type II storms are commonly associated with shock acceleration in the solar corona, and appear as relatively slowly–drifting features, often with a harmonic counterpart at twice the frequency (e.g., Carley et al., 2013; Mann et al., 1996). Type III bursts drift much more rapidly, as they are produced by accelerated particles escaping the solar surface along open magnetic field lines (e.g., Krupar et al., 2014; Morosan et al., 2014; Wild, 1950), even being used alongside HXR observations to narrow down the height of the flare–related acceleration region (Reid et al., 2011). Our focus here
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Figure 1.16: A schematic solar dynamic spectrum. Plasma emission produced by different processes appear as varying spectral structures. Type III bursts, believed to be caused by accelerated electrons escaping along open magnetic fields appear as very steep linear or curved features, while Type II bursts, more associated with shock acceleration sites, appear as segmented slowly drifting features, often exhibiting harmonic emission (Dulk 1985).

will be on Type I radio noise storms, which we will outline in detail below.

1.2.1 Spectral Observations

As with other types of plasma emission, radio noise storms can be identified based on their appearance in a dynamic spectrum, which is an image showing how the radio emission spectrum evolves, with frequency on the vertical axis, and time on the horizontal axis. Noise storms are generally observed to have two components; long-duration, wide-band emission, overlaid by short-lived (<1 s), narrow-band (<10 MHz) spikes in emission, called Type I bursts. An example high-resolution spectrum is shown in Figure 1.17 from Iwai et al. (2013). Type I burst peak flux distributions have previously been shown to exhibit a power-law structure, with
Figure 1.17: A type I radio noise storm dynamic spectrum. *Top:* Left- and right-circularly polarised spectra for a four hour period, exhibiting a storm particularly in the latter two hours of the observation. *Bottom:* Zoomed-in subsets of the above spectra, denoted there by the red box in the RCP image. Here the fine structure of the storm is revealed. Overplotted diamonds denote where the automated detection procedure has marked the presence of bursts (Iwai *et al.* 2013).
1. INTRODUCTION

an index of 2–3 (Mercier & Trottet, 1997; Ramesh et al., 2013). Iwai et al. (2013) however made use of the high spectral resolution of the Assembly of Metric-band Aperture TElescope and Real-time Analysis System (AMATERAS; Iwai et al., 2012b) and an automated burst–finding procedure to show this spectral index may actually be between 4 and 5. In either case, such a power–law form suggests, similar to flare distributions, the presence of an avalanche process producing the accelerated particles responsible for noise storms.

Spectral observations of radio noise storms often reveal highly polarised emission with extremely high brightness temperatures, leading to their common interpretation as plasma emission (Kerdraon & Mercier, 1983; Sundaram & Subramanian, 2004). As storms are often long–lived and comprised of numerous small bursts, they have been presented as evidence for ambient or quiet–sun acceleration outside of solar flares (Raulin & Klein, 1994). Indeed, the possibility of storms being produced in nano– and pico–flares have had important implications for coronal heating (Mercier & Trottet, 1997; Ramesh et al., 2013).

1.2.2 Spatial Structure

Type I storms have been observed to originate from, or above, active regions since their very early observations (Payne-Scott & Little, 1951; Wild & Zirin, 1956). As shown in Figure 1.18, these emissions are interpreted as a column of material, emitting at the plasma frequency, and so at higher frequencies closer to the solar surface (McLean, 1981). The sources themselves have sizes of several arcminutes for the continuum storm, and arcminutes for individual bursts, with source size generally decreasing with emission frequency (Malik & Mercier, 1996).
1.2 Radio Noise Storms

Interestingly, noise storms are commonly seen in the high corona, and associate with areas where active regions interact, rather than above them individually (Brueckner 1983; Lang & Willson 1989; Willson et al. 1990).

Noise storms, like flare emission, are commonly used as tools to gain information on the nature of reconnection and particle acceleration in the corona. For example, joint observations of storms and consistent coronal upflows have been put forward as evidence for gradual interchange reconnection in the corona (Del Zanna et al. 2011; see Figure 1.19). Observations of storms have also been associated with transequatorial reconnection between active regions (Willson 2005) and ‘slipping’ reconnection during a solar flare (Dudík et al. 2014). However, an

Figure 1.18: Interpreted spatial structure of Type I noise storms. Shown are the solar limb as a black curve, with marked sunspots. Above this active region is the volume producing the type I storm. Higher frequency plasma emission originates from lower heights in the corona, where density is higher (McLean 1981).
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Figure 1.19: Noise storm association with SXR active regions. Shown are three full-disk X-ray images produced by the Hinode X-Ray Telescope, for three different times during the passage of an active region across the disk. Overplotted contours show 80% of peak flux levels of 408 (blue), 327 (green), 237 (orange), and 151 MHz (red), observed by the Nançay Radioheliograph (Del Zanna et al., 2011).

alternative acceleration mechanism has been put forward, which notes emerging flux could be super-Alfvénic, and so produce shocks causing acceleration and eventually plasma emission making up a noise storm (Spicer et al., 1982).

While the long duration of storms indicates a lack of clear correlation with solar flares or CMEs, there still exists a relationship with solar activity. A study has shown that noise storms tend to precede CMEs (Ramesh & Sundaram, 2001), and more recently, a decline in noise storm emission has been noted during the release of a CME (Iwai et al., 2013). A number of studies have also shown that storm brightness temperature decreases in association with solar flares (Aurass et al., 1990, 1993; Boehme & Krueger, 1982). This behaviour is explored and interpreted in a study of a recent coronal event in Chapter 6.
1.3 Thesis Outline

The observations described in this chapter are built upon in three work chapters (Chapters 4 to 6), each related to the acceleration and propagation of particles in the active solar atmosphere. Firstly, we advance the understanding of nonthermal processes in solar flares in their earliest phases. During these early phases, it is expected based on the collisional thick target model (CTTM; see Section 2.3.1), that the commonly-observed ‘soft–hard–soft’ evolution of the electron spectral index should have a measurable effect on nonthermal HXR footpoints – producing a variation in height as the beam hardens and softens. This effect has not been observed to date, likely due to the fact that in most flares, emissions at lower energies – which are most sensitive to this variation – and earlier in the flare are dominated by thermal SXRs from heated chromospheric and coronal plasma. This issue is addressed by performing a detailed study of an ‘early impulsive flare’, when thermal SXRs are absent. This study reveals the expected evolution of the HXR footpoints, and has been published in O’Flannagain et al. (2013).

Secondly, we address the issue of HXR source sizes outlined above. As mentioned, it has been revealed by detailed RHESSI studies of HXR footpoints that source vertical extents appear 3–6 times larger in observations than those predicted by theory for standard chromospheric targets. Modelling work has been performed to address this issue by taking into account various coronal density structures and other physical processes than collisions, such as pitch–angle scattering and magnetic mirroring, with no major predicted change in source size as a result. This problem is addressed with a model of locally ionised chromospheric plasma produced by the electron beam, which should serve to vertically
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extend HXR sources. This successfully produces an increase in vertical extent to levels matching observations within commonly observed energy ranges of $\sim 30$ to 70 keV. A paper outlining this work has been published in O’Flannagain et al. (2015).

Finally, a potential observation of the source of accelerated particles in active regions is outlined. To date, a handful of observations of magnetic reconnection exist, usually of apparently 2D reconnecting EUV structures at the solar limb. However, as reconnection in the corona is a 3D process, it is important to also look for examples of null–point reconnection in solar observations (see Section 2.2.1.1 for details). To assist in this, we outline an observation of a collapsing transequatorial EUV loop structure detected by AIA, with an accompanying dynamic Type I noise storm imaged by the Nançay Radioheliograph. The behaviour of the storm emission during the passage of the active region across the disk, and particularly during its collapse, is interpreted as a process of gradual and then enhanced reconnection along a separator produced by the quadrupolar magnetic field. This work is currently in preparation as a paper for publication (O’Flannagain et al. 2015).
In this chapter we outline some of the fundamental physics required to understand the work described in Chapters 4–6. As solar activity is driven by magnetic fields in coronal plasma, the chapter begins by introducing plasma physics and magnetohydrodynamics. These topics are then built upon to describe the various theories of magnetic reconnection, which is believed to be the primary energy release mechanism in solar flares and CMEs. We then introduce the standard solar flare model as it is currently understood. Particular attention is given to the role of reconnection in the nonthermal acceleration of particles and their propagation through the corona and chromosphere. Finally, we outline the emission mechanisms for the two primary types of emission analysed in this thesis: collisional thick target X–rays, and plasma emission.
2. THEORY

2.1 Plasma Physics

A plasma is often defined as an ionised gas, but in fact need not be fully ionised, and can be more accurately described by the three following criteria. First, the plasma oscillation period, $\tau_p$ must be much smaller than the collision timescale, $\tau_c$:

$$\frac{\tau_c}{\tau_p} >> 1, \text{ where } \tau_p = \frac{2\pi}{\omega_p}, \omega_p = \frac{ne^2}{m_e\epsilon_0}$$  \hspace{1cm} (2.1)

A plasma oscillation is caused when a perturbed charged particle experiences a restoring force due to charge separation. Second, for the length scale being considered, $L$, the Debye length, $\lambda_D$ must have a value such that:

$$L >> \lambda_D.$$ \hspace{1cm} (2.2)

The Debye length is the distance at which an electron, given an amount of kinetic energy equal to the average thermal energy of the plasma, will travel before returning due to the electrostatic force. This length therefore defines a scale above which collective effects such as shielding of magnetic charge take place, which are a crucial component of plasma behaviour. Finally, a Debye sphere, or a sphere with a radius equal to the Debye length, must be sufficiently populated:

$$\Lambda >> 1.$$ \hspace{1cm} (2.3)

Here, $\Lambda$ is the plasma parameter, equal to three times the number of particles contained in a Debye sphere, or $\Lambda = 4\pi\lambda_D^3 n$, where $n$ is the plasma electron number density. These criteria indicate that the defining quality of a plasma is
that, on the scale being considered, it is sensitive to the effects of electric and magnetic fields. As such, it is useful now to outline the fundamental equations governing the evolution of these fields.

### 2.1.1 Maxwell’s Equations

Maxwell’s equations form the foundation of classical electrodynamics by describing the interaction between magnetic and electric fields, $\mathbf{B}$ and $\mathbf{E}$, and how they evolve in time and space. In a vacuum, they are given as:

\begin{align*}
\nabla \cdot \mathbf{E} &= \frac{\rho_c}{\epsilon_0} \quad (2.4) \\
\nabla \cdot \mathbf{B} &= 0 \quad (2.5) \\
\nabla \times \mathbf{E} &= -\frac{\partial \mathbf{B}}{\partial t} \quad (2.6) \\
\nabla \times \mathbf{B} &= \mu_0 \mathbf{J} + \mu_0 \epsilon_0 \frac{\partial \mathbf{E}}{\partial t} = \mu_0 \mathbf{J} + \frac{1}{c^2} \frac{\partial \mathbf{E}}{\partial t} \quad (2.7)
\end{align*}

where $\rho_c$ is the charge density and $\mathbf{J}$ is the current density. It is important to note that, as $\mu_0 \epsilon_0$ is by definition equal to $1/c^2$, then the second term in Equation 2.7, which includes the displacement current, can be neglected when typical plasma velocities are significantly below the speed of light, which we hereafter assume to be valid.
2. THEORY

2.1.2 Plasma Kinetic Theory

Maxwell’s equations can be used to calculate the dynamics of single particles when combined with the equation for the Lorentz force:

\[
F_L = q(E + v \times B)
\]  (2.8)

where \( q \) and \( v \) are the charge and velocity of the particle upon which the force acts. However, in order to gain an understanding of how a plasma acts on larger scales, one must investigate the behaviour of distributions \( f(v, r, t) \) with respect to velocity, \( v \), position \( r \) and time \( t \). This is called plasma kinetic theory.

We can develop an expression for the evolution of a particle distribution by starting with a continuity equation, which relates the rate of change of particle number within a phase–space volume \( d^3r \, d^3v \) to the integrated flows of particles into and out of that volume. These flows pass through surface \( dS \) in space and \( dS_v \) in velocity space.

\[
\frac{\partial}{\partial t} \left[ \int f \, d^3r \, d^3v \right] = - \left[ \int f \mathbf{v} \cdot dS \, d^3v + \int f \mathbf{a} \cdot dS_v \, d^3r \right]
\]  (2.9)

Here, \( \mathbf{a} \) is the acceleration where \( \mathbf{a} = d\mathbf{v}/dt \), and so the first and second terms on the right account for flows through \( dS \) and \( dS_v \), respectively. By invoking Gauss’ divergence theorem, we can selectively exchange our surface integrals with volume integrals on the right–hand side of the equation:

\[
\frac{\partial}{\partial t} \left[ \int f \, d^3r \, d^3v \right] = - \left[ \int \frac{\partial}{\partial \mathbf{r}} \cdot (f\mathbf{v}) \, d^3r \, d^3v + \int \frac{\partial}{\partial \mathbf{v}} \cdot (f\mathbf{a}) \, d^3r \, d^3v \right]
\]  (2.10)
Now, by assuming the volume element to be arbitrarily small, and thus removing constant terms from the integrands, and finally substituting $F/m$ for $a$, where $F$ is any external force, we arrive at the collisionless Boltzmann equation:

$$\frac{\partial f}{\partial t} + \mathbf{v} \cdot \frac{\partial f}{\partial \mathbf{r}} + \frac{F}{m} \cdot \frac{\partial f}{\partial \mathbf{v}} = 0 \tag{2.11}$$

which, if the external force $F$ is the Lorentz force, becomes the Vlasov equation:

$$\frac{\partial f}{\partial t} + \mathbf{v} \cdot \frac{\partial f}{\partial \mathbf{r}} + \frac{q}{m}[\mathbf{E} + (\mathbf{v} \times \mathbf{B})] \cdot \frac{\partial f}{\partial \mathbf{v}} = 0 \tag{2.12}$$

Now, in order to evaluate macroscopic terms of plasmas, which are of greater interest observationally, we can take the moments of the Vlasov equation, where the $n^{th}$ moment of a distribution $f(x)$ is $\int f(x)x^n dx$. It can be demonstrated that the zeroth, first, and second moments of the Vlasov equation will produce the equations of conservation of mass, conservation of momentum, and conservation of energy, respectively (for a detailed derivation see for example Goossens, 2003).

### 2.1.3 Magnetohydrodynamics

By deriving these conservation equations we are establishing laws for the dynamics of plasma in electric and magnetic fields, thereby laying the foundations of magnetohydrodynamics (MHD). These conservation equations are given here. The zeroth moment gives the conservation of mass:

$$\frac{Dn}{Dt} = -n \nabla \cdot \mathbf{v} \tag{2.13}$$
where \( n \) is the particle number density. Here, the derivative on the left–hand side is a convective derivative, where \( D/Dt = \partial/\partial t + \mathbf{v} \cdot \nabla \). This essentially accounts for both evolution of the density itself (due, for example, to changes in pressure or temperature), as well as bulk flows.

The second moment gives the conservation of momentum:

\[
mn \left[ \frac{\partial \mathbf{v}}{\partial t} + (\mathbf{v} \cdot \nabla)\mathbf{v} \right] = qn(\mathbf{E} + \mathbf{v} \times \mathbf{B}) - \nabla \cdot \mathbf{P} + \mathbf{P}_{ij} + \sum F \quad (2.14)
\]

where \( \mathbf{P} \) is the pressure tensor, \( \mathbf{P}_{ij} \) is the sum of forces due to collisions, and the \( \sum F \) term accounts for any other forces, such as gravity. In the hot tenuous corona, \( \mathbf{P}_{ij} \) can often be neglected due to the low collision frequency. Additionally, we can take only the diagonal (pressure) terms in \( \mathbf{P} \), neglecting the off–diagonal (shear) terms, giving \( \nabla \cdot \mathbf{P} = \nabla p \), where \( p \) is the (scalar) pressure of the plasma. This leaves us with the simplified expression for momentum conservation:

\[
mn \left[ \frac{\partial \mathbf{v}}{\partial t} + (\mathbf{v} \cdot \nabla)\mathbf{v} \right] = qn(\mathbf{E} + \mathbf{v} \times \mathbf{B}) - \nabla p + \sum F \quad (2.15)
\]

Notably, upon removal of the Lorentz and other forces, this gives the equation of motion for a neutral gas in hydrostatic equilibrium.

Based on this equation of motion, we can demonstrate an important property of plasma in the solar atmosphere. From Equation (2.15) we can derive a general form of Ohm’s law, which relates the time–variation of the electric current density with the Lorentz force applied to the plasma:

\[
\mathbf{J} = \sigma(\mathbf{E} + \mathbf{v} \times \mathbf{B}) \quad (2.16)
\]

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By combining this with Ampère’s law (Equation 2.7), the plasma induction equation is produced:

$$\frac{\partial \mathbf{B}}{\partial t} = \nabla \times (\mathbf{v} \times \mathbf{B}) + \eta \nabla^2 \mathbf{B}$$

(2.17)

where $\eta$ is the resistivity and $\eta = 1/\sigma \mu$. The first and second terms on the right-hand side are called the magnetic advection and magnetic diffusion terms, respectively, and their relative importance is quantified by the magnetic Reynolds number $R_m = \nabla \times (v \times B)/\eta \nabla^2 B \approx vL/\eta$, where $L$ is the length scale over which $B$ varies. Advection is the component of magnetic field line evolution that results from motion of the plasma, while diffusion explains magnetic fields passing through plasma. In the Sun, the conductivity varies at different depths, but above the photosphere it becomes so high that the advective term strongly dominates magnetic field evolution ($R_m \propto 1/\eta = \sigma \mu \gg 1$). The plasma and magnetic field move together – this is known as the ‘frozen-in condition’, and is key to the explanation of solar activity, as it means that plasma motion can distort and stress magnetic fields, which results in the buildup of energy believed to be released in solar eruptive events.

In a simple plasma model, we can identify three types of pressure in a plasma. The forces acting on a portion of plasma under the solar atmosphere are the sum of gravity ($\rho g$), the Lorentz force ($\mathbf{J} \times \mathbf{B}$) and the fluid pressure gradient ($-\nabla p$):

$$\rho_c \frac{dv}{dt} = -\nabla p + \mathbf{J} \times \mathbf{B} + \rho g$$

(2.18)

In equilibrium there is no acceleration so these terms sum to zero, and the gravitational force is weak relative to pressure and electromagnetic effects, so using
2. THEORY

Ampère’s Law, the force balance becomes:

\[
\nabla p = -\frac{1}{\mu_0} \mathbf{B} \times (\nabla \times \mathbf{B})
\]

(2.19)

or, using the vector identity \( \nabla (\mathbf{A} \cdot \mathbf{A})/2 = \mathbf{A} \times (\nabla \times \mathbf{A}) + (\mathbf{A} \cdot \nabla)\mathbf{A} \):

\[
\nabla \left( p + \frac{B^2}{2\mu_0} \right) = \frac{1}{\mu_0} (\mathbf{B} \cdot \nabla)\mathbf{B}
\]

(2.20)

The term on the right hand side is the magnetic tension term, which acts to straighten curved magnetic field lines, but can be neglected if the field does not vary considerably along its axis. Thus the equilibrium state of the plasma and field are determined only by the pressure balance of the portion of plasma with its surroundings, thereby depending on the balance between the gas pressure \( p \) and the magnetic pressure \( B^2/2\mu_0 \). The ratio of these two terms is called the plasma beta parameter, or:

\[
\beta = \frac{p}{B^2/2\mu_0}
\]

(2.21)

The variation of \( \beta \) through the layers of the solar atmosphere is shown in Figure 2.1. This variation is of great importance for various processes related to solar activity.

For example, the high value of \( \beta \) in the photosphere allows hydrodynamic plasma flows to carry magnetic fields, which are connected to chromospheric and coronal plasma in a low–\( \beta \) environment. This means that fields in the chromosphere and corona can become twisted and stressed due to photospheric flows, potentially resulting in energy release and other forms of solar activity.
2.1 Plasma Physics

Figure 2.1: The plasma beta parameter in the solar atmosphere. The plasma beta parameter, or the ratio of gas to magnetic pressure of a plasma, is shown for the photosphere, chromosphere and corona (Gary, 2001).

2.1.4 The Force–Free Approximation

A number of approximations can be made in Equation 2.18 for conditions encountered in the solar atmosphere. Firstly we can again assume the plasma is in magnetohydrostatic equilibrium by replacing the left-hand side of the equation with zero. Then, if we assume that the length scale of interest is much smaller than the density scale–height of the plasma, $L_0 << H$, we can say that $\rho c g << \nabla p$, and remove it from the equation. Finally, if the plasma beta parameter is low such that $L_0 << 2H/\beta$, then we can also neglect the pressure gradient
\( \nabla p \). This leaves us with, simply:

\[
J \times B = 0
\]  
(2.22)

which is known as the force–free approximation, and is widely used to model solar coronal fields.

Equation 2.22 tells us that, if the current density \( J \) is non–zero, then it must be parallel to the magnetic field \( B \), or:

\[
J = \alpha(r) B
\]  
(2.23)

where \( \alpha(r) \) is a parameter which may vary over some space \( r \). Three common forms of \( \alpha \) are generally used in coronal field modelling. First, one can simply assume that \( \alpha = 0 \), meaning there is no current, which is called the potential case. Second, \( \alpha \) can be given a value which varies from one line of magnetic field to another, but is constant along each, called the linear force–free approximation (LFF). Finally, and most generally, \( \alpha \) can be allowed any value at any point, called the non–linear force–free approximation (NLFF).

### 2.1.4.1 Potential Field Extrapolations

It is useful here to describe a common data analysis technique used to estimate the strength and topology of coronal magnetic fields, based on the force–free approximation of coronal plasma. Given a solar photospheric magnetogram such as those introduced in Section 3.2, magnetic fields can be extrapolated upwards into the corona, using either the potential, LFF or NLFF approximations. Here
we outline the theory behind potential field extrapolations, and in particular the ‘mpole’ method \cite{Longcope1996}, which is used in the work described in Chapter 7.

The original method of performing potential extrapolations was to use Green’s functions as first introduced by \cite{Schmidt1964} and expanded upon by \cite{Sakurai1982}. An important property of potential magnetic fields is that, since the current density is zero, we have from Ampère’s Law:

\[ \nabla \times \mathbf{B} = 0 \]  

which has the general solution

\[ \mathbf{B} = \nabla \phi \]  

where \( \psi \) is a scalar potential. Now, from Gauss’ law for magnetism, we know that \( \nabla \cdot \mathbf{B} = 0 \), and therefore \( \phi \) satisfies the Laplace equation:

\[ \nabla^2 \phi = 0. \]  

With measured photospheric magnetic fields \( B_{ph} \) from solar magnetograms, we also have a boundary condition:

\[ -\hat{n} \cdot \nabla \phi(r) = B_{ph}(r) \quad (z = 0) \]  

where \( z \) and \( \hat{n} \) are the height above and unit vector normal to the photosphere, and \( r \) is the three–dimensional spatial coordinate of the volume above and including the photosphere, defined as where \( z = 0 \). Finally, we apply the condition
2. THEORY

\[ \phi(r) \to 0 \text{ as } r \to \infty. \]  

(2.28)

If we define \( r' \) as the location on the \( z = 0 \) surface such that \( r' = (x', y', 0) \), then given these conditions, the Green’s function \( G(r, r') \) for this problem must satisfy:

\[ \nabla^2 G(r, r') = 0 \quad (z > 0) \]  

(2.29)

\[ G(r, r') \to 0 \text{ as } |r - r'| \to 0 \]  

(2.30)

\[ -\hat{n} \cdot \nabla G(r, r') = 0(r) \quad (z = 0, r \neq r') \]  

(2.31)

Finally, we require that as we move towards the \( z = 0 \) surface, unit flux is maintained:

\[ \lim_{z \to 0} \int -\hat{n} \cdot \nabla G(r, r')dS = 1 \]  

(2.32)

where \( dS \) is the surface element. These conditions require that \(-\hat{n} \cdot \nabla G(r, r')\) is a delta function centred on \( r' \):

\[ -\hat{n} \cdot \nabla G(r, r') = \delta(r - r') \]  

(2.33)

thereby establishing our \( G(r, r') \) as the Green’s function for our linear differential operator \(-\hat{n} \cdot \nabla\). The explicit form of our Green’s function is then found to be

\[ G(r, r') = \frac{1}{2\pi |r - r'|} \]  

(2.34)

which is mathematically equivalent to a representation of a magnetic monopole at \( r' \).
This solution provides a basis for treating the photospheric magnetic field as a surface occupied by discrete magnetic monopoles of various polarities and strengths, which is the basis for the ‘mpole’ procedure introduced by Longcope (1996). Using this approximation, the magnetic field in the volume above the $z = 0$ layer can be calculated based on a contribution by each monopole, or

$$B(r') = \sum_i \frac{\phi_i}{2\pi} \frac{r - r_i}{|r - r_i|^3}$$

(2.35)

where $\phi$ is the magnetic flux, and the subscript $i$ corresponds to the $i^{th}$ measured magnetic charge. These charges can be purely simulated, or can be generated from a magnetogram by isolating regions of positive or negative flux based on a chosen threshold parameter. Each region is then simplified to a point with a location $r_i'$ and flux $\phi_i$.

In order to determine topology based on these extrapolated field strengths, linkage between charges is found by tracing field lines $r_B$ using the equation

$$\frac{d r_B(s)}{ds} = \frac{B[r_B(s)]}{|B[r_B(s)]|}.$$  

(2.36)

An example of the resulting extrapolated field is shown in 2D in Figure 2.2 and in 3D in Figure 2.3. The former image represents a 2D slice of an mpole extrapolation of three colinear magnetic charges, while the latter shows a full 3D extrapolation in the case of a quadrupolar active region, produced by two parallel bipoles. Importantly, the topology produced by this code can be used to define separatrix surfaces, which are three-dimensional surfaces bordering regions of different magnetic connectivity. Separators occur at the intersection of these sur-
2. THEORY

Figure 2.2: Potential magnetic field extrapolation from three colinear photospheric magnetic charges using the mpole algorithm. The solid horizontal line represents the photosphere at \( z = 0 \), while curved solid lines represent extrapolated potential fields. Dotted lines represent the ‘mirror’ extrapolated fields underneath the photosphere, which can also be produced by the same method. However, the sub-surface field is generally treated as a cylindrical flux tube, as shown. Triangles represent the location of magnetic null points [Longcope, 1996].

faces, and so are lines in 3D space. Separators link magnetic null points, and play an important role in three-dimensional reconnection theory, which we will cover in the following section.

2.1.5 Magnetic Reconnection

With a foundation in plasma physics and MHD established, we can now investigate what is believed to be the triggering mechanism of impulsive solar activity—magnetic reconnection. Figure 2.4 (a) shows a magnetic geometry where closely-
Figure 2.3: 3D Potential magnetic field extrapolation from a quadrupolar active region using the mpole algorithm. The locations of the magnetic charges are marked by plus signs on top of the synthetic magnetogram. The location of the separatrix surfaces at $z = 0.15$ are shown as solid and dashed lines, demonstrating that their point of intersection denotes a separator, shown as the solid line that passes upwards into the corona. This separator connects the two null points A and B, denoted by triangles (Longcope, 1996).

Separated, oppositely-directed magnetic field lines promote reconnection. This geometry is expected in coronal loops that develop a ‘waist’, where the two legs of the loop approach one another. This leads to a unique scenario where there is a strong magnetic field gradient – from positive at one leg, to equal but negative at the other (Figure 2.4 b). It is normally assumed that the resistivity in the corona is low, such that diffusion (the latter term of Equation 2.17) is negligible. However, in this scenario of strong field gradient, the length scale $L$ over which
2. THEORY

Figure 2.4: The Sweet-Parker reconnection model. a) A schematic of the magnetic field line topology in the diffusion region. b) The magnetic field strength as it varies in the horizontal direction, demonstrating the large magnetic field gradient around the neutral line (Sweet, 1958).

Thus, the field varies is small. Thus, $R_m \approx vL/\eta \ll 1$, and now diffusion dominates:

$$\frac{\partial B}{\partial t} \approx \eta \nabla^2 B \quad (2.37)$$

Thus, the field lines are no longer anchored or frozen in to the plasma, and are able to pass through it into orientations of lower energy state – magnetic reconnection can occur.
2.1.5.1 Sweet-Parker Reconnection

The process explained above is modelled in detail by the Sweet-Parker model of magnetic reconnection [Parker 1957; Sweet 1958]. An important conclusion to be drawn from this is the timescale of energy release, as this is what determines if reconnection can account for the rapid conversion of energy expected to occur in flares. This can be calculated by deriving the electric field, \( E \), inside and outside the diffusion region, and using the boundary layer condition to extract an inflow velocity \( v_i \), which is key to determining how rapidly reconnection will occur.

Outside of the diffusion region, magnetic field lines are nearly straight, so \( \nabla \times \mathbf{B} = 0 \) and so there is no induced current. Therefore, by Ohm’s law the electric field is:

\[
E = -v \times B
\]  

outside the diffusion region. Inside the diffusion region there is a strong current (as the field lines are now sharply curved), and the velocity has decreased significantly, so again using Ohm’s law the electric field inside is:

\[
E = \sigma J
\]  

In the steady state at the boundary between these two regimes, the electric field inside and outside is equal, so that, in the plasma inflow direction, \( v_{in} B_x = \sigma J_z \).

Now, approximating \( B_x \) by Ampère’s law, we have \( B_x = \mu_0 J_z \delta \), where \( \delta \) is the diffusion region width. Thus \( v_{in} = \eta/\delta \). The square of the magnetoacoustic Mach number can now be written, assuming the outflow velocity is the Alfvén
2. THEORY

velocity or magnetoacoustic sound speed $v_A$:

\[
\frac{\nu_m^2}{v_A^2} = \frac{\eta}{\delta} \frac{(v_A \delta/L)}{v_A^2} = \frac{\eta}{v_A L} = \frac{1}{R_A}
\]  

(2.40)

where $L$ is the (longer) length of the diffusion region, and $R_A$ is the magnetic Reynolds’s number using the Alfvén speed, also called the Lundquist number, usually about $10^{14}$ in the solar corona. This indicates a Sweet-Parker Mach number of $M_A \equiv v_m/v_A = 1/\sqrt{R_A}$.

The timescale for reconnection is $t_{rec} = L/v_m = t_A/M_A$, where $t_A$ is the Alfvén transit time, $L/v_A$. Typically, $t_A$ is on the order of $10^1$–$10^2$ s. Using the Sweet-Parker Mach number of about $10^{-7}$ in the corona, the timescale is $t_{rec} \approx 10^8$–$10^9$ s. This indicates a major limitation in the Sweet-Parker model – the reconnection timescale is far too long compared to flare activity which lasts minutes – a problem that was addressed in part by a new model introduced by Petschek in 1964 (Shibata & Magara, 2011).

2.1.5.2 Petschek Reconnection

Petschek’s model is an expansion on the Sweet-Parker model in that it contains a Sweet-Parker diffusion region at the origin (Figure 2.5; Petschek, 1964). However, Petschek also took into account the magnetic field far from this region, where we still have magnetic field annihilation but no diffusion, as it is again a region of advection-dominated magnetic field evolution. As diffusion is no longer an option, standing shock fronts develop where the inflowing plasma is forced to rapidly change its velocity to the Alfvén speed in the horizontal.

In this model, a uniform strong magnetic field $B_e$ external of the reconnection
region decreases to $B_i$ at the upper border to the diffusion region. As the field curves, it is altered by the addition of a component normal to the shock front, $2B_n$. This distortion is accounted for by first assuming that it is produced by a continuous array of magnetic dipoles of moment $m$ (Priest & Forbes 2000). The magnetic flux at a distance $r$ from one of these dipoles is $r\pi m/r = \pi m$. However, if the dipoles are separated by $dx$, the flux contributed by one dipole is also $2B_n dx$. Equating these we have $m = 2B_n/\pi dx$. To determine the field at the boundary, we integrate the moment over $x$, using $L$ and $L_e$ as the distances.
at the edge of the diffusion region and at large distances from it, respectively:

\[
\frac{1}{\pi} \int_{-L}^{L} - \frac{2B_n}{x} \, dx - \frac{1}{\pi} \int_{L_e}^{L} - \frac{2B_n}{x} \, dx = \frac{4B_n}{\pi} \ln \left( \frac{L}{L_e} \right) = - \frac{4B_n}{\pi} \ln \left( \frac{L_e}{L} \right)
\]  

(2.41)

and so the field at the boundary becomes:

\[
B_i = B_e - \frac{4B_n}{\pi} \ln \left( \frac{L_e}{L} \right).
\]

(2.42)

In the switch-off shock regime, the shock travels at \(v_s = B_n/\sqrt{\mu \rho} \), or in the case of a standing shock, the inflowing external plasma has this velocity and the shock is stationary, \(v_e = v_s = B_n/\sqrt{\mu \rho c} \). This means that \(M_A = v_e/v_A = B_n/B_e \), so that Equation 2.42 can be rewritten:

\[
B_i = B_e (1 - \frac{4M_A}{\pi} \ln \left( \frac{L_e}{L} \right))
\]

(2.43)

Petschek used this derived relation to determine a maximum reconnection rate based on \(M_A \), where this field distortion component was enough to equal the original external field, so that the factor in parentheses in Equation 2.43 is equal to \(1/2 \). From this, and again assuming an outflow velocity of \(v_A \) the maximum Petschek Mach number is \(M_A \approx \pi/8ln(R_A) \). Given the values used in the above calculations for the Sweet-Parker reconnection timescale: \(t_{rec} = t_A \pi/8ln(R_A) \), which, using the same values, is on the order of seconds, much more in line with the observed duration of flaring emission.

It should also be noted that modern simulations have shown that uniform resistivity in the Petschek model leads once more to elongated Sweet–Parker layers \cite{UzdenskyKulsrud2000}. These elongated layers have variously been
2.2 Solar Flare Models

Figure 2.6: Geometry of a Sweet–Parker layer that is unstable to plasmoids (Daughton & Roytershteyn, 2012).

shown to be unstable to breakup into smaller magnetic islands, or plasmoids as shown in Figure 2.6 (Matthaeus & Lamkin, 1985). Recently, plasmoids have been proposed to play a key role in mediating fast reconnection (Kliem et al., 2000). The formation of current sheets between islands, which are also unstable to plasmoids, leads to rapid reconnection with average rates of $R \approx 0.01$ (Huang & Bhattacharjee, 2010).

2.2 Solar Flare Models

These emissions can briefly be described by the interpretation presented in Figure 2.7, which shows schematically the summary of the CSHKP model (Carmichael, 1964; Hirayama, 1974; Kopp & Pneuman, 1976; Sturrock, 1966). Magnetic field dissipation occurs in the corona, causing particle acceleration along new field lines towards the chromosphere (see Section 2.2.1). As these nonthermally-accelerated electrons spiral downwards, they emit synchrotron radiation as microwaves. Since
2. THEORY

Figure 2.7: A schematic of a flaring loop. Top: Accelerated particles at the flare looptop produce nonthermal radio, microwave and X-ray emission. Energy is transported to the footpoints by thermal conduction and the accelerated particles, which produce nonthermal HXRs, extreme UV and Hα emission. Heated plasma here then expands to fill the loop, emitting thermal radiation in a wide range of wavelengths (Dennis & Schwartz, 1989).
the electron energies are of the order of 1keV and above, they traverse the ‘leg’ of the flare loop in a matter of seconds, when they lose their energy in free-free processes at the footpoints, emitting HXRs and γ-rays. Depending on their energy, they may penetrate further into the atmosphere, emitting X-ray bremsstrahlung in the chromosphere, and for large flares, white light in the photosphere (Hudson et al. 1992).

2.2.1 Particle Acceleration

Magnetic reconnection, as has been stated, serves as a mechanism to convert stored magnetic potential energy into kinetic energy and heat by the acceleration of coronal electrons and ions. The simplest cause of acceleration is of course that of the application of a strong electric field.

The reconnecting current sheet (RCS) describes the plane located between two magnetic fields of opposite direction during magnetic reconnection (the central horizontal plane in Figure 2.5). As plasma and magnetic field lines flow in (vertically in Figure 2.5) at inflow velocity \( v_i \), they are accelerated outwards (horizontally) by an induced electric current (Zharkova & Gordovskyy, 2004) derived above:

\[
E = \frac{1}{\sigma}J - v \times B = \eta \nabla \times B - v \times B
\]

(2.44)

where the first term dominates in the diffusion region, while the second dominates in the frozen-in scenario far from the origin. Given an inflow velocity of 1% the Alfvén speed (Priest et al., 1997), and solar coronal plasma parameters, efficient particle acceleration is produced simply by the Lorentz force.

However, there are further mechanisms that may contribute to particle ac-
celeration at reconnection sites. The standing shocks required in the Petschek model are an attractive mechanism as they would produce the power-law distribution in X-ray spectra that are almost uniformly observed. However, shocks are unable to accelerate low-energy particles effectively, requiring a ‘seed’ injection of accelerated electrons, and the topology of the shocks required for efficient acceleration does not seem to be present in flares (although shocks may be the source of other astrophysical emissions). A more likely contender would be stochastic acceleration in slow magneto-acoustic waves, where electrons are scattered by other randomly-oriented electrons. Head-on collisions become more likely with increasing energy, resulting in bulk acceleration of electrons tied to newly-reconnected magnetic field lines (Zharkova et al. 2011).

2.2.1.1 Acceleration in 3D Reconnection at Null Points

The fundamental theories of reconnection outlined previously rely on 2D models of reconnection, which are an idealised version of the magnetic field in the solar corona. 3D reconnecting fields are more representative of the mechanisms taking place, and so are a field of increasing interest, with clear distinctions being made between 2D and 3D reconnection theory (e.g., Birn & Priest 2007). As such, we will now outline the background behind models of acceleration during reconnection at a 3D null point as described by Dalla & Browning (2005) and Browning et al. (2010), which is of particular relevance to the work presented in Chapter 7.

In these models, a single–particle approach is taken. In order to determine the primary forces acting on a charged particle, we begin again with our equation of motion (Equation 2.15), excluding all forces but the Lorentz force, and assuming a negligible electric field such that \( m(dv/dt) = q(E + v \times B) \). By taking the dot
2.2 Solar Flare Models

product with \( \mathbf{v} \), and setting the right–hand side to zero, as \( \mathbf{v} \perp (\mathbf{v} \times \mathbf{B}) \), we have

\[
\frac{d}{dt} \left( \frac{m|\mathbf{v}|^2}{2} \right) = 0
\]  

(2.45)

which demonstrates that a static magnetic field cannot change the kinetic energy of a particle. Then, if we split \( \mathbf{v} \) into its perpendicular and parallel components, \( \mathbf{v} = \mathbf{v}_\perp + \mathbf{v}_\parallel \), and noting that \( \mathbf{v}_\parallel \times \mathbf{B} = 0 \), our equation of motion can be split into these two components also:

\[
\frac{d\mathbf{v}_\parallel}{dt} = 0
\]

(2.46)

\[
\frac{d\mathbf{v}_\perp}{dt} = \frac{q}{m} (\mathbf{v}_\perp \times \mathbf{B})
\]

(2.47)

The former equation demonstrates that the velocity along the magnetic field is unaffected by this force, while the latter equation describes an acceleration that forces particles to orbit a guiding centre. As a result, charged particles can be highly magnetised, and therefore are treated with the guiding–centre approximation, which approximates particle motion as directly along these centres, neglecting gyration.

A force which works against the magnetisation of particles is the electric drift force, which imparts a velocity of the particle perpendicular to the guiding centre in an electric field \( \mathbf{E} \) and magnetic field \( \mathbf{B} \) field of

\[
\mathbf{v}_E = c \frac{\mathbf{E} \times \mathbf{B}}{|\mathbf{B}|^2}.
\]

(2.48)
2. THEORY

This results in a radius of gyration, or Larmor radius of

\[ r_L = \frac{mc^2E}{qB^2}. \]  

(2.49)

The degree of magnetisation, which determines whether the drift force is enough to discount the guiding centre approximation is therefore

\[ \epsilon = \frac{r_L}{L} = \frac{mc^2E}{qB^2L}. \]  

(2.50)

In the solar corona, on global length scales, it is often the case that \( \epsilon \ll 1 \), and therefore particles can be safely treated to move along their guiding centres. However, this approximation is shown not to hold close to magnetic null points.

A simple 3D magnetic null point can be defined by the magnetic field \( \mathbf{B} = (B_x, B_y, B_z) \), where

\[ B_x = B_0 \frac{x}{L} \]  

(2.51)

\[ B_y = B_0 \frac{y}{L} \]  

(2.52)

\[ B_z = -2B_0 \frac{z}{L}. \]  

(2.53)

where \( B_0 \) is the magnetic field strength near the null, and \( L \) is the characteristic length scale. A visual example of this geometry is given in Figure 2.8 Here, the null point is located at \((0,0,0)\), where each component of the magnetic field linearly approaches zero. Here, the degree of magnetisation, \( \epsilon \) asymptotically approaches infinity, and so the guiding centre approximation rapidly becomes insufficient, and the full equation of motion must be used to determine particle
Figure 2.8: Geometry of a magnetic null point. 

a) Sampled magnetic field lines based on Equations 2.51 - 2.53. b) During reconnection, inflow and outflow respective to the null point are expected, as highlighted in the 3D shaded regions here. (Dalla & Browning 2005).

Models using this setup have shown, for a random distribution of test particles, that electrons are efficiently accelerated to energies as high as 100 keV. Furthermore, the steady–state distribution of particles was found to be that of a power–law (Browning et al. 2010), as shown in Figure 2.9. For a magnetic field strength of both 100 G and 20 G, the distribution rapidly evolves from a thermal distribution to a nonthermal power–law, as indicated by flare HXR inversions.

2.2.2 Density Models and Nonuniform Ionisation

Regardless of the initial acceleration mechanism or location primary energy release, the density, temperature and ionisation fraction of the corona and chromo-
2. THEORY

Figure 2.9: Model evolution of a proton spectrum accelerated at a null point. On the left, evolution of the spectrum is shown (from black to blue to green to orange to red) for a magnetic field of 100 G, with the initial (black) spectrum representing the chosen initial distribution. On the right, the evolution is shown with the same colour evolution in the case of a magnetic field of 20 G \cite{Browning2010}.

sphere are still crucial in understanding their role as a target in the HXR and radio emission mechanisms outlined in the following section. As such, we outline the current solar atmospheric models here, for quiet–sun conditions as well as in flaring active regions.

The most commonly used models of solar density and temperature are the VAL \cite{Vernazza1981} and FAL \cite{Fontenla1990} models. The VAL model is produced using the full radiative transfer and hydrodynamic modelling of Fraunhofer lines detected in the Sun by the Skylab instrument. As a result of fitting emission from H, H\textsuperscript{−}, C, Si, Fe, Mg, Al, He, He II, Ca II, Mg II among others, a model distribution over height of plasma temperature and density were produced, as shown in Figure 2.10. This model provides a basis for the target of a nonthermal electron beam impacting on a quiet–Sun plasma. However, to account for the enhanced and redistributed density in active and flaring regions, more specific models have been developed.
2.2 Solar Flare Models

Figure 2.10: The VAL model of the solar atmosphere. Shown are the temperature (solid curve) and mass density (dashed curve) variation with height above the solar surface, which is on the right of the plot. Notable are the sharp drop in temperature and rise in density at the transition region, followed by the roughly exponential rise in density with depth into the chromosphere, where the temperature reaches its minimum (Vernazza et al., 1981).

As shown in Section 2.3.1 and expanded further in Chapter 4, the collisional thick target model can be used to predict HXR source heights from HXR spectra when using an input density model such as those outlined here. Conversely, HXR spectra can be informed by images, which provide source locations, to produce empirical density models (e.g., Liu et al., 2006). Saint-Hilaire et al.
2. THEORY

Figure 2.11: Statistically derived density structure of the flaring solar atmosphere. The black solid line represents the density structure based on a fit to results of a statistical RHESSI study based on HXR spectra and locations (Saint-Hilaire et al., 2010).

(2010) outline a survey, performing this technique on 838 flares in order to produce a model density structure of an ‘average flaring chromosphere’, shown in Figure 2.11. In this study it was found that the atmosphere is well described by a double–exponential density structure, with a scale–height of $131 \pm 16$ km in the chromosphere, and $5.4 \pm 0.6$ Mm in the corona.

These target models have been used as starting points in modelling the solar atmospheric response to an injection of nonthermal electrons. Allred et al. (2005) use the RADYN (Carlsson & Stein, 1992, 1997) suite of radiative and hydrodynamic models to determine how the temperature, density, and ionisation fraction
of the corona and chromosphere will respond to impulsively injected electron beams over short timescales. This results in a detailed form of non-uniform ionisation (NUI), which is the concept that the atmosphere encountered by the beam has a varying ionisation fraction. The NUI established at the transition region between the completely ionised corona and neutral chromosphere has previously been presented as a potential cause of the power-law break seen in RHESSI spectra (Su et al., 2009). The effect of the local NUI presented in the models of Allred et al. (2005) on the size and structure of RHESSI sources is explored in detail in Chapter 5.

2.3 Nonthermal Emission Mechanisms

In this final section we move away from the fundamental physics required to understand flares, and towards the physics of the generation of electromagnetic radiation during solar activity. In particular, we describe the primary models for interpreting the nonthermal HXR and radio emission analysed in the work presented in this thesis, namely collisional thick target model (CTTM) emission and plasma radiation.

2.3.1 The Collisional Thick Target Model

The collisional thick target model (CTTM Brown, 1971; Hudson, 1972) describes the relationship between the HXR emission rate spectrum, $I(\epsilon_x)$ (photons cm$^{-2}$ s$^{-1}$ keV$^{-1}$), and the distribution of injected electrons $f_0(\epsilon_0)$. A bremsstrahlung cross-section, $\sigma(\epsilon, \epsilon_x)$, determines the probability that an electron of energy $\epsilon$ will undergo a free-free interaction with a positively charged nucleus and emit a photon of energy $\epsilon_x$. 

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\( \epsilon_x \). The total X-ray spectrum at distance \( r \) to the observer is

\[
I(\epsilon_x) = \frac{1}{4\pi r^2} \int_{\epsilon_x}^{\infty} f_0(\epsilon_0) \left( \int_{t(\epsilon=\epsilon_x)}^{t(\epsilon=\epsilon_0)} \sigma(\epsilon, \epsilon_x)v(\epsilon)n(\epsilon)dt \right) d\epsilon_0
\]  

(2.54)

where \( n(\epsilon) \) and \( v(\epsilon) \) are the electron density and velocity, and \( \epsilon_0 \) is the electron energy at injection. In this way, the integral inside the brackets describes the total number of electrons that emit a photon of energy \( \epsilon_x \) while braking.

To solve this integral we first need to know how the energy of an injected electron varies as it propagates through the target plasma. Here we make use of the important assumption of the CTTM; that energy losses experienced by the injected electrons are purely due to collisions; there are no losses due to the magnetic field. The rate of change of an electron’s energy is therefore

\[
\frac{d\epsilon}{dt} = -Kn(\epsilon)v(\epsilon) \frac{\epsilon}{\epsilon}
\]

(2.55)

where \( K = 2\pi e^4 A \), and \( A \) is the Coulomb logarithm, which accounts for the range of distances of nearest approach between an ion and electron which would result in significant energy loss.

With this relation we can simplify Equation 2.54:

\[
I(\epsilon_x) = \frac{1}{4\pi r^2} \int_{\epsilon_x}^{\infty} f_0(\epsilon_0) \left( \int_{\epsilon_0}^{\epsilon_x} \sigma(\epsilon, \epsilon_x) \frac{\epsilon}{K} \right) d\epsilon_0
\]

(2.56)

Given an observed HXR spectrum \( I(\epsilon_x) \), determining the injected electron spectrum \( f_0(\epsilon_0) \) represents an inverse problem, which was solved in Brown (1971). In this work, it was assumed that the photon spectrum had a power–law form, given
2.3 Nonthermal Emission Mechanisms

by

\[ I(\epsilon_x) = I_1 \frac{\gamma - 1}{\epsilon_1} \left( \frac{\epsilon_x}{\epsilon_1} \right)^{-\gamma} \]  

(2.57)

where \( \gamma \) is the power–law index or slope of the spectrum, and \( I_1 \) is the reference photon flux at energy \( \epsilon_1 \).

On inversion of Equation 2.56, the solution for the injection spectrum was found to be:

\[ f_0(\epsilon_0) = 2.56 \times 10^{33} \gamma^2(\gamma - 1)^3 B \left( \gamma - \frac{1}{2}, \frac{3}{2} \right) \frac{I_1}{\epsilon_1^2} \left( \frac{\epsilon}{\epsilon_1} \right)^{-(\gamma+1)} \]  

(2.58)

where \( B(a, b) \) is the Beta function \( B(a, b) = \int_0^1 x^{a-1} (1 - x)^{b-1} dx \).

With this model, observations made of HXR spectra can be used to determine the injected electron distribution, potentially revealing the nature of the acceleration process itself. The CTTM has also been expanded upon to produce useful predictions on HXR source height distribution (Brown & McClymont, 1975) based on the concept that higher-energy (or faster) nonthermal electrons penetrate deeper before losing their energy. This method is described in further detail in Section 4.3.

2.3.2 Plasma Emission

The generation of plasma emission is a non–linear process, generally summarised into three steps: (1) A nonthermal process, here taken to be a nonthermal distribution of charged particles, is generated in a thermal plasma, and either initially exhibits a positive slope \((df(v)/dv > 0)\), or this property is reached by the process of fast particles outpacing slower particles (shown in Figure 2.12); (2)
2. THEORY

Figure 2.12: The ‘bump–on–tail’ instability. The solid line represents a particle distribution which will lose energy to Langmuir waves with phase velocity between \( V_1 \) and \( V_2 \), which is where the slope is positive. This mechanism will cause a flattening of the bump. If the distribution was produced by faster particles outpacing slower ones, the bump will evolve toward lower energies, as shown by the dashed line [Lin et al. 1981].

This distribution becomes unstable to the generation of Langmuir waves; and (3) The generated Langmuir waves interact, resulting in emission of radiation at the plasma frequency and its harmonics.

The generation of Langmuir waves can be found by introducing a perturbation \( f_1(\mathbf{r}, v, t) \) to a stable distribution \( f_0 \), where \( f_1 << f_0 \), in what is called ‘quasi–linear theory’ [Vedenov 1963]. This perturbation is periodic in space and time,
and so is given a functional form \( f_1 \propto e^{i(\omega t + kr)} \). To model this perturbation we substitute it into the Vlasov equation and rearrange to produce:

\[
f_1 = \frac{q}{m} \left( \frac{j}{\omega - \mathbf{k} \cdot \mathbf{v}} \right) \mathbf{E} \cdot \nabla f_0.
\]  

(2.59)

Importantly, the denominator of the function describing the perturbed quantity includes the term \( \omega - \mathbf{k} \cdot \mathbf{v} \), which approaches zero as the speed of the wave, \( \omega/\mathbf{k} \) approaches the speed of the particles, \( \mathbf{v} \). This produces a singularity which requires a special treatment around this condition, known as contour integration (Melrose, 1989). This method reveals a complex form of the dispersion relation, such that the solution of our perturbation in the region \( \mathbf{v} \approx \omega/\mathbf{k} \) goes as

\[
e^{i[(\omega + i\gamma) t]} = e^{i\omega t} e^{-\gamma t}.
\]  

(2.60)

This is a damped wave with damping factor \( \gamma \), which is the essence of Landau damping, wherein a Langmuir wave loses energy to a particle distribution. However, as the damping factor can be either positive or negative, this can also explain wave growth, which occurs as a result of the bump–on–tail instability.

The Langmuir wave produced in this way will have the plasma frequency given by

\[
\omega_p = \sqrt{\frac{n_e e^2}{m \epsilon_0}}
\]  

(2.61)

and can produce fundamental plasma emission by scattering off an ion-acoustic wave. For these interactions, the momentum equations must be obeyed, and so
2. THEORY

for two waves, 1 and 2, producing a wave, 3, the following must hold:

\[ \omega_1 + \omega_2 = \omega_3 \]  \hspace{1cm} (2.62)
\[ k_1 + k_2 = k_3 \]  \hspace{1cm} (2.63)

However, as the ion–acoustic wave frequency in this interaction is much smaller than the plasma frequency, fundamental plasma emission is expected to oscillate at approximately \( \omega_P \). The second harmonic emission occurs when two Langmuir waves interact, propagating in opposite directions, resulting in observable emission at \( \omega = 2\omega_P \). These emissions, usually observed as radio bursts in the solar case, can then be observed and used to gain information on accelerated particles in the active Sun.
In this Chapter, the instruments used to make observations of solar flares and related coronal phenomena are described. A major focus of this thesis is the observation of coronal and chromospheric sources of nonthermal hard X-ray emission, and so the RHESSI instrument, which makes high–resolution images and spectra of solar flare X–rays, is described in detail. Following this, two of the Solar Dynamics Observatory (SDO) instruments, the Atmospheric Imaging Assembly (AIA) and the Helioseismic and Magnetic Imager (HMI) are introduced, as they are both useful in developing an understanding of the magnetic backdrop for coronal reconnection. Finally, the Nançay Radioheliograph (NRH) is described, including an introduction to interferometric imaging. The NRH was used in identifying ambient and rapid reconnection in the solar corona.
3. INSTRUMENTATION

3.1 The Ramaty High Energy Solar Spectroscopic Imager (RHESSI)

The Ramaty High Energy Solar Spectroscopic Imager (RHESSI; Lin et al., 2002) is a NASA Small Explorer mission designed to produce images and spectra of solar X-rays and γ-rays. The objective of the mission is primarily to gain insight into the nonthermal processes responsible for the production of X-rays in solar flares, namely reconnection, particle acceleration and energy transport. RHESSI was launched on 5 February 2002, and has therefore provided over thirteen years of X–ray observations of the Sun, covering the decay of solar cycle 23, and the rise of solar cycle 24.

Due to RHESSI’s ability to image HXRs at a maximum spatial resolution of 2.25 arcsec while simultaneously providing spatially-integrated spectra, it is ideal for studying the positions and geometries of HXR sources of emission, as is the focus of the science goals detailed in Chapters 4 and 5.

3.1.1 RHESSI Imaging

RHESSI records counts of solar X-rays using nine cooled hyperpure germanium detectors, each with no inherent imaging capability but a spectral resolution of \(~1\) keV below 100 keV and \(~3\) keV up to 1 MeV (Smith et al., 2002). Each of these detectors is located behind a pair of aligned grids which are made of X-ray opaque materials (tungsten for all grids except grid pair 1, which uses molybdenum). See Figure 3.1 for an outline of the grid and detector geometries. Grid pairs are numbered 1–9, with logarithmically increasing slit spacings, resulting in nominal
3.1 The Ramaty High Energy Solar Spectroscopic Imager (RHESSI)

Figure 3.1: A schematic of the RHESSI imaging apparatus. Incoming solar X-rays enter at the top, encountering the first and rear set of grids on their path to the germanium detector at the rear of the instrument. The fraction of X-ray flux which reaches the detector is sensitive to their angle of incidence. As the entire spacecraft rotates once every four seconds, this results in a modulated lightcurve which is encoded with spatial information on the incident X-rays (Hurford et al., 2002).

full–width at half–max (FWHM) resolutions of 2.26, 3.92, 6.78, 11.8, 20.4, 35.5, 61.1, 105.8, and 183.2 arcseconds. Rotation the spacecraft once every four seconds causes the incident X-rays to be periodically blocked, resulting in the recording of a modulated time profile at the detectors. As the orientation of the spacecraft as well as the geometric properties of the grids are known, this time profile can be used to reconstruct information on the origin of incident X-rays.

The encoding of the HXR spatial parameters into the modulated time profiles can be understood upon inspection of Figure 3.2. As shown, the high–resolution time profile of recorded HXRs can vary in amplitude, phase, and frequency de-
3. INSTRUMENTATION

Figure 3.2: Demonstration the relationship between ideal RMC modulated profiles and HXR source brightness, extent, and location on the solar disk. On the left are idealised HXR fluxes plotted against time, with the corresponding model source on the right inset (Lin et al., 2002).

pending on the geometry of the source. The phase and frequency of the profile depend on the angular and radial coordinates of the model Gaussian HXR source, respectively. The latter is due to the fact that a source further from the centre of the solar disk will be passed over by RHESSI’s slats a higher number of times within a single spacecraft rotation than a source near disk centre. Further, RHESSI is sensitive to source size based on the fact that a source larger than the
grid spacing can be visible through more than one slit at a time, and so is never fully attenuated, resulting in a lightcurve of higher troughs and lower peaks, as shown in panel 6 of Figure 3.2.

To reconstruct a RHESSI image, an algorithm is first selected. The list of available algorithms increases with time, as new techniques are constantly being developed. In the case of this thesis, the Back Projection (Mertz et al., 1986), CLEAN (Högboom, 1974), and Visibility Forward Fit (VIS FWDFIT; e.g., Kontar et al., 2008b, 2010) techniques were used. These techniques are detailed below.

3.1.1.1 Back Projection

Back projection is the most basic and fastest method available for RHESSI image reconstruction, and often serves as a starting point for many of the more involved algorithms. If over a given time interval, $\Delta t_i$, with a HXR count flux distribution $F_m$ across the sky with some spatial coordinate $m$, we expect RHESSI to register a total number of counts $C_i$ given by:

$$C_i = \sum_m P_{im} F_m \Delta t_i.$$  \hspace{1cm} (3.1)

Here, $P_{im}$ is the modulation pattern, which describes the 2-D probability distribution of a given photon originating from a position $m$ in the sky (Hurford et al., 2002). This pattern is dependent on the known geometry and orientation of RHESSI’s grid pairs, and neglecting slat transparency at high energies and photon scattering, can be approximated by a triangular wave with minima of 0% and maxima of 50%. These two extremes correspond to scenarios when all light from a given direction passes through the front grid also passes through the rear
Figure 3.3: Demonstration of the RHESSI Back Projection algorithm. To the far top–left, a map \((F_m)\) containing a simple Gaussian source at an arbitrary position is input. From left to right, (and in terms of Equation 3.1), the modulation pattern \((P_m)\), measured count flux \((C_i)\), and their product, the weighted probability map \((P_m \cdot C_i)\) are then shown. To the right of these is a running sum of the weighted map. Here, the left column covers the first half–rotation of the instrument, and the right covers the second. As shown, this process gradually builds up a Back Projection image, known as a ‘dirty map’, exhibiting heavy sidelobes. The use of multiple detectors, and post–processing algorithms such as CLEAN, can be used to reduce these sidelobes. Figure credit: NASA.
grid to be detected, and when all light passing through the front grid hits a slat on the rear grid resulting in no detection, respectively. The period of the triangle wave is then given by the spacing of the grids as projected onto the Sun.

As the recorded count flux, $C_i$, is summed over our spatial coordinate $m$, Back Projection must use the known modulation profile, $P_{im}$, to restore information about the spatial distribution of the HXR source. It does so by producing a weighted probability map, $P_{im} \cdot C_i$, for each time interval within a full rotation. These probability maps essentially tell us how likely it is that emission is coming from the areas of the solar disc which are unobscured by RHESSI’s grids at a given phase of the spacecraft rotation. By summing the maps over a full rotation (or a number of full rotations), a ‘dirty’ map of HXR intensity can be produced:

$$I_m = \frac{1}{A} \sum_{i=1}^{N} \frac{C_i}{\Delta t_i} \cdot P_{im}$$  \hspace{1cm} (3.2)

Here, $A$ is the effective detector area ($cm^{-2}$). The resulting map, $I_m$ has units of $counts\ cm^{-2}\ s^{-1}$, and is our first estimate of the distribution of HXR flux across the Sun. Figure 3.3 serves as a demonstration of how summing these weighted maps over a full cycle reproduces a map of the source. As RHESSI’s nine grids can only sample a finite number of spatial frequencies, maps produced by Back Projection tend to exhibit strong sidelobes, or rings of intensity around the primary source. These sidelobes can be removed afterward by more involved algorithms, such as CLEAN.
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Figure 3.4: The CLEAN algorithm as performed by the RHESSI software suite. The peaks are partially removed from the ‘dirty’ map and convolved with a Gaussian to produce the COMPONENT map (top-right), which is scaled to fill the colour table (bottom right), and is finally summed with residuals left over from the dirty map to produce the CLEAN image (bottom-left).

3.1.1.2 CLEAN

 CLEAN (see Figure 3.4) is the most commonly used method of removing sidelobes from a ‘dirty’ map, and is inherited from the field of radio astronomy (Högboom, 1974). The algorithm works by locating the pixel of peak flux, taking some
3.1 The Ramaty High Energy Solar Spectroscopic Imager (RHESSI)

fraction of the flux value, known as the gain (normally 10 %) and producing a point source with a flux of that value. This point source is then convolved with a Gaussian, and placing it on a new ‘CLEANed’ map. The ‘beam-width’ of this Gaussian is chosen by the user, but is generally intended to reflect the maximum resolving power of the grids being used. This process is iterated a predefined number of times, or until the model time profile based on the CLEANed map matches the recorded profile within acceptable limits.

3.1.1.3 Visibility Forward Fit

Visibility Forward fit (VIS_FWDFIT) makes use of a RHESSI observable known as its visibility. A RHESSI visibility is a measure of the contrast (or amplitude) of a periodically modulated HXR lightcurve. RHESSI visibilities are produced by subdividing a modulated lightcurve, such as those shown in Figure 3.2, into a number of bins, and stacking these bins over subsequent rotations in order to increase signal–to–noise. The oscillating count flux in these stacked bins are then fit with sinusoids, from which a visibility phase and amplitude can be calculated. Visibilities are useful because they contain information on the spatial distribution of HXRs, but are not dependent on instrumental effects. Additionally, because of the wide use of visibilities in radio astronomy, a large number of fast imaging algorithms have already been developed, and can now be used with RHESSI data.

VIS_FWDFIT works by comparing these visibilities with those produced by a synthetic source. The user defines a source structure (usually some combination of 2D Gaussian components, which can be symmetric, elliptical, or curved), and starting parameters (e.g., central coordinates, FWHM, aspect ratio, radius of curvature). These parameters are then varied by VIS_FWDFIT until an ac-
ceptable $\chi$–squared, or measure of goodness–of–fit is reached. Thus the output of this algorithm is not strictly a map of HXR intensity, but parameters defining the Gaussian source structure. These parameters can be used directly if source size or location is of interest to the user, or can be used to reconstruct maps.

This method has some advantages and disadvantages over CLEAN. In direct comparison with a CLEAN image, a VIS_FWDFIT image will have no remaining sidelobes or spurious sources (because these are not included in the model fit). This however is an indicator of the primary disadvantage of VIS_FWDFIT; it is heavily dependent on the user–defined model source structure. For this reason, a Back Projection or CLEAN image is often inspected first for the time and energy interval of interest, so a decision on the accuracy of some combinations of Gaussian sources can be made. If a structure described by a combination of Gaussian components is accurate, VIS_FWDFIT can theoretically provide source positions and FWHMs with an accuracy surpassing the spacing of the finest grid ([Kontar et al., 2010]). Such observations are built upon in Chapter 5.

3.1.2 RHESSI Spectroscopy

RHESSI produces spectra based on the process of photoelectric absorption, Compton scattering and electron/hole production within the germanium detectors. An X-ray or $\gamma$-ray photon creates an electron/hole pair, which are attracted to the cathode/anode, respectively, registering as a time-tagged pulse. These detections are binned over a certain amount of time, producing a count spectrum ($CS$).
3.1 The Ramaty High Energy Solar Spectroscopic Imager (RHESSI)

This can be inverted to a photon spectrum (PS) via

\[ CS = BG + DRM \ast PS, \]

where BG is background count rate and DRM is the detector response matrix.

The X-ray background count rate is made up of a combination of instrumental, solar, and non-solar emissions. As RHESSI orbits the Earth, it enters a 30-minute eclipse roughly every 90-minutes. During eclipse, data is not recorded by RHESSI, except for a five-minute buffer on either end. These intervals can be used for background subtraction, as they should represent all non-solar and instrumental flux. If there are other active regions or events occurring at the same time as a given observation, these can often be sufficiently well subtracted by sampling time intervals shortly before and after the flare of interest.

The RHESSI DRM is a matrix relating the recorded count spectrum to the incident photon spectrum. The DRM accounts for a wide range of physical processes which cause incident photons to either be registered at their true energy with less than 100% efficiency (diagonal elements), or to be shifted to another incorrect (usually lower) energy (off-diagonal elements). These processes include absorption or Compton scattering by various components of the spacecraft, Compton scattering of photons off the Earth’s atmosphere, noise in the electronics, and resolution degradation due to radiation damage accumulated in the detector over RHESSI’s life cycle. The DRM matrices for each detector were produced before launch by measuring the response of the detectors to 3.7 and 6.1 keV emission from radioisotopes at various angles, and fitting the recorded countrate to the known spectrum (Smith et al., 2002).
Figure 3.5: A sample RHESSI spectrum including a thermal (red) and nonthermal (green) fit.

Once background subtraction has been done, and a photon count spectrum recovered, model-based fits can be applied in an attempt to determine the components of the emission (Figure 3.5). $v_{th}$, the variable thermal fit component, produces the expected photon distribution emitted via thermal bremsstrahlung, requiring abundance information, temperature, and emission measure ($EM = \int_V n_e^2 dV$ where $n_e$ is electron density and $V$ is volume of emitting material). The model is produced using the following photon spectrum ([Aschwanden 2005]):

$$I_{vth}(\epsilon) = 2.6 \times 10^7 \left( \frac{EM_{49}}{T_7^{1/2}} \right) \frac{1}{\epsilon_{keV}} \exp\left(-\frac{\epsilon_{keV}}{0.86T_7}\right) \left( photons \ cm^{-2} s^{-1} keV^{-1} \right)$$

(3.4)

where $EM_{49}$ and $T_7$ are the emission measure and temperature in units of $10^{49} cm^{-3}$.
and $10^7 K$ respectively. The thermal model can also include Gaussian fits to the Fe/Ni line complexes at 6.7 and 8 keV. There is also a common emission line at 10 keV, however this is dominated by instrumental emission in the germanium detectors, and so is usually removed in background subtraction. Another fit component, thick2, produces a model spectrum based on the collisional thick target model (see Section 2.3.1). Given a bremsstrahlung cross-section, the input parameters to this model are the parameters describing the nonthermal electron spectrum which goes on to cause the HXR emission.

The electron spectrum is usually observed to be well-described by a double power-law with a break in the slope at an energy $E_b$. The parameters required to describe this distribution are the total electron flux, $f_1$, the low- and high-energy cutoffs $E_{low}$ and $E_{high}$, the break energy $E_b$, and the spectral indices below and above the break energy, $\delta_{low}$ and $\delta_{high}$. The spectral index in a power-law model defines the relative number of electrons which are accelerated to high energies, and since the CTTM predicts that high-energy electrons penetrate deeper into chromospheric and coronal plasma, this would suggest that a hardening spectrum should drive emission to lower radial altitudes. This expected relationship is explored fully as the main focus of Chapter 4.

# 3.2 The Solar Dynamics Observatory (SDO)

The Solar Dynamics Observatory is a NASA-designed satellite, launched on 11 February 2010, with the mission of studying the solar atmosphere on small spatial scales and high temporal cadence in order to better understand the physics of the Sun–Earth connection \cite{Pesnell2012}. At launch, three instrument suites
were included onboard in pursuit of this goal, namely the Atmospheric Imaging Assembly (AIA; Lemen et al., 2012), the Helioseismic and Magnetic Imager (HMI; Scherrer et al., 2012), and the Extreme Ultraviolet Variability Experiment (EVE; Woods et al., 2012). The former two were used in this thesis as part of a rare observation of magnetic reconnection (see Chapter 6), and so are described in detail in this Section.

3.2.1 The Atmospheric Imaging Assembly (AIA)

One method of investigating energy release and transport in the solar corona and chromosphere is to analyse extreme ultraviolet (EUV) images of plasma at the temperatures characteristic of those regions of the atmosphere. As this plasma is ‘frozen in’ to the magnetic field in the corona, coronal loop structures are often used as a proxy of concentrations of magnetic field. As the theoretical driver of solar active energy release, magnetic reconnection, is believed to occur on short timescales and over small areas, it is ideal to make these EUV images with as high a temporal cadence and spatial resolution as possible. For this reason, the Atmospheric Imaging Assembly (AIA) was used to assist in the tracking of field line motion in the work outlined in Chapter 6.

The AIA expands upon the observations of previous instruments such as the Transition Region And Coronal Explorer (TRACE; Handy et al., 1999) and the Extreme Ultraviolet Imaging Telescope (EIT; Delaboudinière et al., 1995) onboard the Solar and Heliospheric Observatory (SOHO), which both continue to observe the Sun in the extreme ultraviolet (EUV) wavelength range. The AIA builds upon the capabilities of these instruments by producing full–disk images
3.2 The Solar Dynamics Observatory (SDO)

Figure 3.6: Schematic of AIA telescope design. Visible or UV light enters through the aperture on the left, and passes to the primary mirror located around the right baffle. Solar light is reflected into the left baffle, where it is reflected again off the active secondary mirror. The secondary mirror takes information recorded by the guide telescope and stabilises the image for the CCD. The filter wheel and shutter are finally located between the right baffle and the CCD. These two components determine the wavelength of light to be recorded and the exposure time, respectively [Lemen et al. 2012].

through eight EUV and visible narrowband and two continuum filters with a resolution of $\sim 1.5$ arcsec every 12 seconds.

The basic imaging apparatus of one of AIA’s Cassegrain telescopes is shown in Figure 3.6. The AIA is comprised of four of these telescopes, each differing in the available filters in the filter wheel. Each telescope has two sets of multilayer filters comprised of Zirconium, to center transmitted light on 131 Å and 93 Å, and Aluminium to focus on 335 Å, 193 Å, 211 Å, 171 Å, and 304 Å. An exception is telescope 3, which instead of one filter, has three Al/MgF$_2$ filters to focus on lower–energy ultraviolet light; 1600 Å, 1700 Å, and 4500 Å.

These wavelengths are chosen with the specific purpose of detecting line emission from C IV, He II, various ionization levels of Iron, and in the case of telescope 3, continuum emission (See Table 3.1). The lines chosen are known to be pro-
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<table>
<thead>
<tr>
<th>Channel</th>
<th>Primary ion(s)</th>
<th>Region of atmosphere</th>
<th>Char. log(T)</th>
</tr>
</thead>
<tbody>
<tr>
<td>4500 Å</td>
<td>continuum</td>
<td>photosphere</td>
<td>3.7</td>
</tr>
<tr>
<td>1700 Å</td>
<td>continuum</td>
<td>temperature minimum, photosphere</td>
<td>3.7</td>
</tr>
<tr>
<td>304 Å</td>
<td>He II</td>
<td>chromosphere, transition region</td>
<td>4.7</td>
</tr>
<tr>
<td>1600 Å</td>
<td>C IV + cont.</td>
<td>transition region, upper photosphere</td>
<td>5.0</td>
</tr>
<tr>
<td>171 Å</td>
<td>Fe IX</td>
<td>quiet corona, upper transition region</td>
<td>5.8</td>
</tr>
<tr>
<td>193 Å</td>
<td>Fe XII, XXIV</td>
<td>corona and hot flare plasma</td>
<td>6.2, 7.3</td>
</tr>
<tr>
<td>211 Å</td>
<td>Fe XIV</td>
<td>active–region corona</td>
<td>6.3</td>
</tr>
<tr>
<td>335 Å</td>
<td>Fe XVI</td>
<td>active–region corona</td>
<td>6.4</td>
</tr>
<tr>
<td>94 Å</td>
<td>Fe XVIII</td>
<td>flaring corona</td>
<td>6.8</td>
</tr>
<tr>
<td>131 Å</td>
<td>Fe VIII, XXI</td>
<td>transition region, flaring corona</td>
<td>5.6, 7.0</td>
</tr>
</tbody>
</table>

Table 3.1: Summary of AIA channel details, including wavelength, primary ion or source of emission, targeted region of the chromosphere, and characteristic temperature.

duced at specific temperatures, thus allowing us to probe different regions of the solar atmosphere, depending on their temperature. As shown in Table 3.1, AIA images cover the temperature range of log(T)=3.7–7.0, from the photosphere to the hot flaring corona.

The response of each of the optically thin coronal channels can be seen in Figure 3.7. These temperature response functions were produced using the atomic database, CHIANTI (Dere et al., 1997, 2009), to produce a contribution function $G(T, \lambda)$ of an ion at temperature $T$ to emission at wavelength $\lambda$, assuming the coronal abundances given by Feldman & Widing (1993). This contribution function is then convolved with the wavelength response of the instrument to produce the temperature responses shown in Figure 3.7. These functions can be used to associate observed coronal features with a most–probable plasma temperature, keeping in mind the width and the dual-peaked nature of the curves.

An example set of images produced using AIA is shown in Figure 3.8. The hot-
3.2 The Solar Dynamics Observatory (SDO)

Figure 3.7: AIA temperature response functions for each of the six ‘coronal’ channels. Each curve corresponds to the digital number (DN) recorded at the AIA CCD as a result of incident emission from plasma of a given temperature.

...ter coronal channels (top: 131 Å, 94 Å, 335 Å, middle: 171 Å, 193 Å and 211 Å) exhibit a complex twisted coronal loop. The bottom row, which contains images using the cooler channels (1600 Å, 335 Å, and a line–of–sight magnetogram) demonstrates that the coronal loops are connecting sunspots and other concentrations of opposite polarity located underneath in the photosphere. These images demonstrate the capability of AIA to resolve fine structure in coronal loops. This capacity is explored further by imaging apparent reconnection inflow and outflow in Chapter 6.
3. INSTRUMENTATION

**Figure 3.8:** Sample early images produced using AIA shortly after launch. *Top row, left to right:* 131 Å, 94 Å, 335 Å. *Middle row:* 171 Å, 193 Å and 211 Å. *Bottom row:* 1600 Å, 304 Å, and a line-of-sight magnetogram produced with SDO’s Helioseismic and Magnetic Imager (HMI) (Lemen et al., 2012).

### 3.2.2 The Helioseismic and Magnetic Imager (HMI)

The HMI was designed to measure Doppler shift, line-of-sight magnetic field, continuum intensity, and vector magnetic field on the solar surface. A successor to the Michaelson Doppler Imager (MDI; Scherrer et al., 1995), the HMI includes a number of significant improvements. A notable difference is the use of the Fe I 6173 Å absorption line which is believed to have greater sensitivity to magnetic fields than MDI’s Ni I 6768 Å line (Norton et al., 2006). As the work outlined in this thesis only makes use of line-of-sight magnetograms from HMI, the process
by which these are produced is outlined here.

HMI produces magnetograms by first measuring the Zeeman effect. The Zeeman effect takes place when emitting or absorbing atoms are exposed to a magnetic field \((\text{Zeeman} 1897)\). This field will exert a torque on the magnetic dipole moment associated with an electron’s orbital angular momentum. This removes the degeneracy of different electrons with identical principal and orbital quantum numbers \((n \text{ and } l)\), but different magnetic quantum numbers, \(m_l\). For an orbital of angular momentum \(l\), magnetic quantum numbers range from \(-l\) to \(l\), such that in an atom exposed to a magnetic field, there are \(2l + 1\) energy states within each orbital. A key element is that the difference in energy between two adjacent split levels \((\Delta m_l = 1)\) is directly related to the applied magnetic field \(B\), given by:

\[
\Delta E = \frac{e\hbar}{2m}B
\]

(3.5)

where \(e\), \(\hbar\) and \(m\) are the electron charge, the reduced Planck constant, and the electron mass, respectively. As a result of this energy splitting, transitions to or from split energy levels produce multiple emission or absorption lines. In any given transition, selection rules dictate that the change in magnetic quantum number must be \(\Delta m_l = -1, 0, \text{ or } +1\). Converting Equation 3.5 from energy to wavelength of emitted or absorbed light, and substituting known values, we have:

\[
\Delta \lambda = \pm 46.67g_j\lambda_0^2B
\]

(3.6)

where \(g_j\) is the Landé g–factor, and \(\lambda_0\) is the central wavelength of the line from material at rest without the presence of a magnetic field.

It is shown in Equation 3.6 that an absorption or emission line will exhibit
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**Figure 3.9:** Schematic diagram of the HMI instrument onboard SDO. Sunlight enters from the top–right into the front window filter, importantly then passing through the Lyot filter and wide– and narrow–band Michaelson filters. The light then terminates at the front and side CCDs (Couvidat et al. 2012).

A splitting, or in the case of wide, closely–spaced lines, an apparent broadening when exposed to a magnetic field, and this broadening will be proportional to the strength of the field, $B$. The HMI is designed to accurately measure this broadening in the Fe I absorption line produced in the solar photosphere. To scan across the Fe I absorption feature, the HMI includes a tunable Lyot filter and two tunable Michaelson filters, as indicated in Figure 3.9. The Lyot filter has five elements, with passband FWHMs ranging from 0.69 Å to 344 mÅ, while the two Michaelson filters have FWHM of 172 mÅ (wide band) and 86 mÅ (narrow band). The resulting range of all included filters is 76 mÅ at the center of the tuning position (Schou et al. 2012).
3.2 The Solar Dynamics Observatory (SDO)

Figure 3.10: HMI tuning positions. Six positions are shown in different colours which sample the Fe I solar line which is shown in black. Each position is sampled once every 45 seconds (Schou et al., 2012).

This combination of filters results in narrow, tunable transmission profiles presented in Figure 3.10. As shown, six filtergrams are produced for each pixel every 45 seconds, each centred at a different wavelength along the Fe I absorption feature, shown in black. Line-of-sight magnetic field strengths are produced as a by-product of the observed Doppler velocities. In order to do so, theoretical Fe I line profiles were convolved with each of the detector responses shown in Figure 3.10 for a range of velocities, producing a lookup table of the Fourier components of the series describing the line profile. Once velocities are produced for left-circularly polarised (LCP) and right-circularly polarised (RCP) light, the
Figure 3.11: An example full–disk magnetogram produced by the HMI instrument. These data were taken from the date of an active region collapse observed on 6 July 2013 at the active region just east of disk–center. This event is explored fully in Chapter 6.

The line–of–sight magnetic field is produced by:

\[ B = (V_{LCP} - V_{RCP})K_m \]  

(3.7)
where $K_m$ is a calibration constant produced by Equation 3.6 and has the value:

$$K_m = 9.34 \times 10^{-5} \lambda_0 g_j c = 0.231 \ G \ s \ m^{-1}. \quad (3.8)$$

The result of this data processing is a full-disk magnetogram as shown in Figure 3.11. In these magnetograms, black and white areas represent regions where the magnetic field is determined to point away from and towards the observer, respectively. The active region to the left of disk-center in this image is explored in much greater detail in Chapter 6.

### 3.3 The Nançay Radioheliograph (NRH)

Like hard X-ray observations, measurements of solar radio emission can provide insight into the nonthermal processes that occur during solar activity, and for this reason, the Nançay Radioheliograph (NRH: Kerdraon & Delouis, 1997) was of great use in making the observation of magnetic reconnection outlined in Chapter 6. NRH can produce interferometric images of solar radio emission in up to ten frequency bands between 150 and 450 MHz every 125 ms. The resolution of the images is approximately 3.2–5.5 arcmin at 164 MHz and 1.25–2.2 arcmin at 432 MHz.

The current iteration of NRH consists of an array of antennas oriented in a ‘T’ shape (see Figure 3.12). The east–west subset is made up of 19 antennas, four of which have parabolic collectors consisting of four thick dipoles, providing orthogonal linear polarisation between 150 and 400 MHz (Radioheliograph Group, 1989). The remaining 15 are variations of thick dipoles, with only one linear polarisation. This set of antennas provide a mixture of baselines ranging from
Figure 3.12: Schematic of the current NRH setup, highlighting the East–West and North–South baselines of 3200 m and 1295 m, respectively (Kerdraon & Delouis, 1997).

50 m to 3200 m. The north–south portion of the array is made up of 24 five meter dishes providing baselines from 54.3 m to 1248 m.

The signal recorded at each detector is first passed through filters which remove the 200–210 MHz and 88–108 MHz (FM) frequency bands, in order to negate terrestrial interference. It is then passed into a frequency mixer, which changes the observing frequency to 113 MHz before transmission to the NRH lab through buried coaxial cables. The receiver for the north–south detectors is located on pedestals, and so corrections must be made for changes in gain due to temperature variation throughout the day. The east–west cables are buried and
so do not require corrections as frequently.

An additional step in the preparation of NRH images is calibration, which is performed in real time on the receiver outputs by a Concurrent Computer Maxion biprocessor. Calibration data are taken from observations of a known intense radio source, usually Cygnus A. From these observations gain and phase offsets are determined, and these are then applied in real time to recorded solar data.

One of the major upgrades from the previous version of NRH is the replacement of the analog correlator with a digital one. Previously, NRH was limited to producing fast 1D images only, as the analog correlator produced just 55 correlations. With the upgraded digital correlator, all potential correlations between receivers are available, numbering 861, with 576 being produced after neglecting redundant correlations. This new setup allows for the production of full-disk 2D images produced at a cadence of 5 ms over up to 10 observing frequencies between 150 and 450 MHz. The spatial resolution ranges from 0.3 to 6.0 arcmin, depending on frequency and direction. Example images produced during a solar radio burst shortly after the NRH upgrade are shown in Figure 3.13. The method by which these radio images are produced is described here.

### 3.3.1 Interferometric Imaging

In order to produce 2D images, Fourier imagers such as RHESSI and radio interferometers such as the NRH make use of an observable known as the visibility. The meaning of visibility can be understood by examining the output of a simple two-element interferometer. As shown in Figure 3.14, such a system consists of
two receivers separated by a baseline vector $\mathbf{b}$. Radio emission from an angle $\theta$ in the sky will produce voltages at each detector with a phase offset $\tau_g$. As the delay is the time taken for the radio wave (travelling at the speed of light, $c$) to
3.3 The Nançay Radioheliograph (NRH)

Figure 3.14: A schematic of a simple two-element interferometer. Two detectors are separated by baseline vector $\mathbf{b}$. The detected signals $V_1$ and $V_2$ at each detector are sent to a correlator where the multiplied and time-averaged response $R$ is produced. *Figure Credit: National Radio Astronomy Observatory.*
3. INSTRUMENTATION

traverse the extra distance \(|b|\cos(\theta)\) to receiver 1, it is given by:

\[
\tau_g = \frac{|b|\cos(\theta)}{c}. \tag{3.9}
\]

The two detectors thus record voltages at time \(t\) of

\[
V_1 = V\cos[\omega(t - \tau_g)] \text{ and } V_2 = V\cos(\omega t). \tag{3.10}
\]

The purpose of the correlator is to multiply the signals, then average them over time \(\delta t\). This time interval is chosen such that \(\delta t >> \omega^{-1}\), thus producing a correlated voltage, or interferometer response, of:

\[
R = \langle V_1 V_2 \rangle = \frac{V^2}{2}\cos(\omega \tau_g). \tag{3.11}
\]

Responses to different values of \(\tau_g\) can then be recorded by either manually moving detectors, or allowing the Earth’s rotation to do so, effectively sampling different projected baselines. This response represents the real part of the visibility of the brightness distribution across the sky, and is the direct observable of an interferometer.

The relationship between the observable visibility and the sky brightness distribution is given by the Van Cittert–Zernike Theorem (van Cittert, 1934; Zernike, 1938). From this theorem we can state that:

\[
V(u, v) = \int I(l, m)e^{i2\pi[ul+vm]}dldm \tag{3.12}
\]

where \(V(u, v)\) is the visibility, sampling spatial frequencies \(u\) and \(v\). \(I(l, m)\) is
the brightness distribution across the sky, with angular coordinates $l$ and $m$. This gives a mathematical relationship between interferometric visibilities and the spatial distribution of flux across the sky. Thus, to determine the desired
3. INSTRUMENTATION

\( I(l, m) \), we simply take the inverse Fourier transform:

\[
I(l, m) = \int V(u, v)e^{-i2\pi[u(l+m)]}dudv,
\]  

(3.13)

which produces an image. It is interesting to note here that the back projection process is equivalent to a discrete inverse fourier transform, highlighting the analogy to RHESSI imaging (see Section 3.1.1).

However, this is an ideal case, notably with an infinite sampling of the \((u, v)\) plane. Each point on the \((u, v)\) plane is sampled by using interferometer pairs of different projected baselines, so since an interferometer with \(N\) detectors can only produce \(N(N - 1)/2\) pairs, only a finite sampling can be achieved. The result of various levels of sampling is shown in Figure 3.15. As shown, for fewer antenna pairs, strong sidelobes are present. Even with a large number of detector pairs, sidelobes are diminished but still present. Remaining sidelobes can be further removed by the use of image deconvolution algorithms such as CLEAN or the more complex algorithms outlined in Section 3.1.1.
Coronal Hard X-ray Source Response to Electron Spectral Hardening

The collisional thick target model (CTTM) predicts that more energetic electrons penetrate to greater column depths along the flare loop. This requires that sources produced by harder power-law injection spectra should appear further down the legs or in the footpoints of a flare loop. Therefore, the frequently observed hardening of the injected power-law electron spectrum during flare onset should be concurrent with a descending hard X-ray source. In this Chapter we test this implication by making a detailed observation of an early impulsive event. We analyse RHESSI spectra to determine that observed emission is nonthermal above $\sim 7$ keV, and investigate images to demonstrate for the first time that the predicted behaviour exists, and should be visible in all early impulsive events.

This work has been published in Astronomy and Astrophysics (O’Flannagain et al., 2013). All of the analysis outlined in this Chapter was performed by the first author, with regular guidance provided by the co–authors.
4. CORONAL HARD X-RAY SOURCE RESPONSE TO 
ELECTRON SPECTRAL HARDENING

4.1 Introduction

The analysis of nonthermal X-ray emission is extremely important in explaining the process which causes impulsive energy release. Observable properties such as nonthermal X-ray source position are expected to depend on the nature and evolution of the accelerated electron spectrum. However, nonthermal emission is frequently masked by thermal emission in the early phase of the flare, making it difficult to investigate nonthermal processes before the peak in hard X-rays (HXRs). There exist a small number of recorded events in the database of Ramaty High Energy Spectroscopic Imager (RHESSI; Lin et al., 2002) called early impulsive flares, which can be identified by a delay of $\sim 30$ s or less between the initial rise in soft X-ray flux and the impulsive rise in HXR flux. Sui et al. (2007) outline analysis of 33 such events, in which plasma preheating is minimal, and so nonthermal emission may be the primary contributor to the RHESSI spectrum even before the peak in HXRs. Early impulsive flares are essential to gaining an understanding of the behaviour of nonthermally accelerated electrons at the earliest phases of an event. In addition, due to their predominantly nonthermal spectra, even down to low energies, these events reveal the behaviour of low-energy electrons which are stopped in the tenuous corona.

Nonthermal coronal X-ray sources have previously been suggested as evidence for coronal magnetic reconnection (Frost & Dennis, 1971; Masuda et al., 1994) and plasmoid-looptop reconnection (Milligan et al., 2010). In the RHESSI era, numerous studies have been carried out on occulted flares, where the bright nonthermal footpoint emission is masked by the solar limb, allowing observations of possibly nonthermal looptop emissions which are normally outside of the dynamic
4.1 Introduction

range of the instrument (e.g. Balciunaite et al. 2002, Krucker et al. 2007). Coro-
nal nonthermal emission has been shown to be temporally correlated with Type
III radio bursts (Krucker et al. 2008), further supporting the argument for the
existence of a nonthermally accelerated electron population in the corona. Loop-
top source motion has previously been interpreted as a signature of transition
from X-type to Y-type reconnection during a flare (Sui & Holman 2003).

Early impulsive flares provide an opportunity to observe faint looptop non-
thermal emission without sacrificing information on the behaviour at the foot-
points during the HXR peak. This therefore allows for the detection of any source
motion between the coronal looptop and chromospheric footpoints. During the
rise phase of a typical flare, the flux of HXRs reaches a peak and the spectral
Based on the theoretical derivations of nonthermal X-ray intensity with height in
the coronal acceleration scenario (Brown & McClymont 1975), this is expected
to result in a descent of the location of peak nonthermal emission in the time
coming up to the HXR peak. This downward motion of HXR sources during
the hardening of the electron spectrum is visualised in Figure 4.1. Previously, it
was suggested that such downward motion was observed in the C1.1 class early
impulsive flare that occurred on 28 November 2002 (SOL2002-11-28T04:37, Sui
et al. 2006). In this event, a faint looptop source appeared, split into two, de-
cended down both loop legs, and reached the footpoints at the time of the peak
in HXRs. An in-depth analysis of this behaviour will help to test the thick target
model during the earliest phase of nonthermal emission, which is rarely observed.

In this Chapter, we model descending X-ray sources by taking into account
the time variation in the spectral index of the electron injection spectrum. We
4. CORONAL HARD X-RAY SOURCE RESPONSE TO ELECTRON SPECTRAL HARDENING

Figure 4.1: *Top:* Model HXR photon spectra for injected power-law electron beams with spectral index of 5 (left), 4.5 (middle), and 4 (right). Overplotted are the photon spectrum from each figure to the left, in order to clarify the difference between subsequent spectra. Thus, the rightmost plot includes spectra produced by nonthermal electrons of index 4 (solid line), 4.5 (dashed line) and 5 (dotted line). *Bottom:* Synthetic images based on the injected beam parameters used to produce the spectra above. The 1D profile of HXR flux along the path of the electron beam were produced using the model outlined in Section 4.3, then multiplied across by a Gaussian and curved to demonstrate what a flare loop observation might look like for a hardening spectrum.

suggest that a descent of HXR sources in the rise phase of a flare can be explained by hardening of the electron injection spectrum. In Section 4.2 the 28 November 2002 flare observations and analysis are presented. Section 4.3 the model used to determine theoretical source positions is described, predicting the dependence of source height on spectral index and observed photon energy. In Section 4.4 we present the results of this analysis, and in Section 4.5 interpretations are drawn based on the comparison of our theoretical models and these observations.
4.2 RHESSI Observations

A C1.1-class solar flare was observed by RHESSI on 28 November 2002, beginning at 04:35:30 UT, with HXR emission observed for roughly 50 s (Figure 4.2a). The flare was located near the Sun’s western limb, with unocculted footpoints. RHESSI was in attenuator state A0, meaning there were no aluminium attenuators in front of the detectors during the event. As a result, RHESSI was able to detect X-rays with energies as low as 3 keV. Throughout the event, flare emission was observed up to energies of \( \sim 50 \) keV.

Time intervals were selected to produce as many independent images as possible without creating noise-dominated X-ray source maps of this low-count flare detection. One 16 s interval was used from the start of the flare at 04:35:24 UT until 04:35:40 UT. From that point on, images were made by integrating flux over 8 s, until the end of the final interval at 04:38:00 UT. In order to aid in the automated tracking of source peaks, we laid overlapping time intervals in between each of these intervals, resulting in a total of 36 images per chosen energy band. Energy bands were selected to focus on the low-energy part of the spectrum and set at 3–6 keV, 6–8 keV, and 8–10 keV, producing reliable imaging of source motion in all energy ranges. Images produced using higher energy bands were noise dominated for all time intervals except during the peak in HXRs. Consequently they were excluded from this analysis, with the exception of 25–50 keV emission at the HXR peak, which was used to estimate the location of the flare loop footpoints.

Figure 4.2 gives a summary of the RHESSI observations. The descent of X-ray sources down two legs of an apparent loop documented by Sui et al. (2006) is
**Figure 4.2:** (a) X-ray Lightcurve of the flare of 28 November 2002. The 12–25 keV curve is scaled by a factor of five for clarity. Four times are marked, corresponding to the start times of the four images and spectra shown below. Overplotted is the electron power-law index derived from the spectral fits (dotted line), demonstrating the concurrence of maximum spectral hardness (minimum spectral index) with peak in HXR emission. (b) Spatially integrated spectra for the times of their corresponding images using pre-flare background subtraction. Overlaid are thermal and nonthermal fits constructed using the OSPEX spectral analysis suite. Residuals, or the difference between observed and model-based X-ray flux, normalized to the one-sigma uncertainty in the photon flux, are plotted below each spectrum. (c) RHESSI image contours corresponding to energy bands of 3–6, 6–8, and 8–10 keV, with a contour showing 20–50 keV at $t_3$, the HXR peak. These images are centred on the west limb, denoted by the solid line. Contours represent 75% of the peak emission of the image, with a second 50% contour included for the 8–10 keV image at interval $t_2$, in order to show the location of the southern source. Images are generated using the CLEAN algorithm available in the RHESSI image analysis software. Each of the intervals used for these images are eight seconds in duration, beginning at the time shown in (b).
4.2 RHESSI Observations

immediately evident upon study of RHESSI images (Figure 4.2c). A crucial step in modelling this behaviour was determining at what times and energies emission appeared to be nonthermal, especially within the energy range of 3–10 keV, which is well below the more common estimates of upper limits to the nonthermal low-energy cutoff of \(~20–40\) keV (e.g. [Holman et al., 2003]). However, more recent work which corrects for albedo effects suggests cutoffs of less than 12keV ([Holman, 2012] Kontar et al., 2008a, for recent discussion). Thus, before images could be interpreted based on thick-target emission of X-rays, we analysed high-resolution RHESSI spectra in order to separate nonthermal from thermal emission.

4.2.1 Spectroscopy

Full-Sun spatially integrated spectra were produced over the duration of the flare, using the same time intervals as those chosen for the imaging. We used detector 4 on RHESSI due to its high spectral resolution of \(~1\) keV at energies below \(~100\) keV ([Smith et al., 2002]). Because of the low average count flux of the flare, significant noise was present, especially in the time before the HXR peak. This meant that, for many of the time intervals selected, different combinations of thermal and nonthermal fit components could be used with equally good comparisons with data. These components included the thermal, thick target, and Gaussian line options provided in the OSPEX suite of algorithms ([Kaastra et al., 1996]). We found that the thermal component could be fitted by a continuum variable thermal (\(v_{\text{th}}\)) model, with a Gaussian line to account for the iron line complex emission at 6.7 keV. In some fit attempts, a thermal continuum component was not even necessary prior to the HXR peak. However, a full (line plus continuum)
model could also be used to achieve equally good fits to the observed spectrum, based on the $\chi^2$ test provided in OSPEX. In order to remain consistent with the thermal interpretation of the production of the iron line complex, the full thermal model was selected for this work.

During the fitting process, we investigated the sensitivity of the fit to the variation of the low-energy cutoff. This cutoff is a notoriously difficult parameter to derive from RHESSI spectra (Sui et al., 2005a) as usually only an upper limit to its value can be established. For all time intervals prior to the HXR peak, it was found that the $\chi^2$ value of the fit was almost constant with different initial values of low-energy cutoff, ranging from 1 to 15 keV. This further indicates that the highest values that still produced good fits can only be seen as upper limits to this parameter. As such, the cutoff was assumed to be at an energy less than 5 keV for this analysis, which allowed the use of a smooth injection spectrum without a cutoff for the modelling. Further justification for this fully nonthermal interpretation is given by analysis of the images (see Sect. 2).

As shown in Figure 4.2b, the resulting spectral fits demonstrate that emission is predominantly nonthermal for the phase of the flare prior to the HXR peak, except for the iron line feature at 6.7 keV. Given RHESSI’s dynamic range of $\sim$ 1:10, it is likely that both types of emission are observable simultaneously below $\sim$7 keV (Hurford et al., 2002). It could be argued that this significant thermal emission is accounted for by the apparently thermal looptop source present during the early phase of the flare, even after the initial sources have descended down the loop. However, comparison of the total counts associated with this source and with the footpoint sources indicate that the looptop emission cannot alone produce all of the thermal emission indicated by the spectra. If the descending
sources are produced by an injection of nonthermal electrons, localised heating and thus thermal emission are to be expected at the site where energy deposition is at its peak. Therefore low-energy footpoint emission may be a combination of thermal and nonthermal emission.

An estimate of the displacement between thermal and nonthermal footpoint emission can be made by approximating the distance covered by evaporating plasma over the time since the initial beam penetration. If ablation of chromospheric material begins at 04:35:48UT, and given standard evaporation velocities of $\sim 100-200$ km s\(^{-1}\) (e.g., Milligan et al. [2006b]), this would result in a displacement of thermal emission by $\sim 2.4-4.8$ Mm at 04:36:12UT, the latest interval used in this study. In reality, the thermal source would be continually replenished by the ablation at the footpoint, so these displacement values are an upper limit only. Indeed, by modelling radiative and convective energy release following beam heating, Allred et al. (2005) determine that, for an impulsive flare such as this, the displacement between the location of peak energy deposition and peak footpoint temperature can be as little as 0.3 Mm after 6 s for an impulsive event. Therefore, while the emission may have a thermal contribution, this displacement error is small enough so that the X-ray sources of all energies will be hereafter used as a proxy for location of peak nonthermal energy deposition. As such, it is appropriate to model their motion based purely on the location of the peak of the simulated nonthermal photon distribution with height, which is derived in Sect. 4.3.
Figure 4.3: RHESSI images produced by scanning through finely-spaced energy bins for two time intervals. Top: Images are shown for increasing energy from left to right, continuing in a second row, for energies from 4 to 51.5 keV, with finer spacing at the lower part of the spectrum, where counts are higher. The time interval for this set of images was 68 seconds long, starting at 04:35:00 UT, thus covering the full rise phase of the flare. A long interval was chosen in order to achieve the required counts for imaging. Bottom: Images for identical energy bands, produced over a time interval of 24 seconds, starting at 04:36:00. This time interval covers the peak and decay phases of the flare.

4.2.2 Imaging

Images were reconstructed using the CLEAN algorithm, with detectors 3, 4, 5, 6, 8, and 9 \cite{Hogbom1974, Hurford2002}. Detector 1 was excluded because the fine spatial resolution (~2.3 arcsec) tended to add small peak emission near the larger sources, making automated source-tracking unreliable. Detectors 2 and 7 were excluded as their imposed lower threshold energy is at least ~9 keV.
4.2 RHESSI Observations

Imaging revealed the motion of X-ray sources down and up the legs of the flare loop previously noted by Sui et al. (2006) (Figure 4.2c). A 3–10 keV source appears just west of the limb at 04:35:40 UT and descends ~12 Mm down the apparent flare loop to reach the footpoints at 04:36:08 UT, which coincides with the peak in hard X-rays. Following this, the source rises ~11 Mm to return to a looptop position, where it remains until soft X-ray emission returns to pre-flare level. This motion is seen in all three energy bands used for imaging, although the sources exhibit different qualitative behaviours before and after the HXR peak. Before the peak, the higher energy sources are located lower in the loop, descending slower and at different rates, covering ~13 Mm, ~9 Mm, and ~3 Mm in the 3–6, 6–8, and 8–10 keV bands respectively. After the peak, the distance travelled by each source is roughly constant with emitted photon energy, and higher energy emission originates higher in the loop, contrary to the ordering observed during the descent.

It is also informative to investigate the HXR images at different finely spaced energies, as shown in Figure 4.3. If the source is nonthermal, one would expect that for increasing photon energy the source will appear lower in the corona or chromosphere. For the early time interval covering the rise phase of the flare, shown at the top of Figure 4.3, this appears to be the case, excluding energies around the iron and iron–nickel lines at 6.7 and 8 keV, where emission relocates to the looptop. This shows that emission at higher, and perhaps even lower, energies originates from the footpoints and is potentially nonthermal, in agreement with the spectra. In the images produced over the peak and decay phase of the flare, shown in the lower panel of Figure 4.3, we see that the bright footpoints dominate the looptop source even around 6–8 keV, but below 6 keV, emission is primarily
located at the looptop. This suggests that at this stage of the flare, enough chromospheric evaporation has taken place to allow thermal emission to begin to dominate the lower part of the spectrum.

In order to compare with predictions of the thick target model, source position with time and energy was quantified (Figure 4.4). We chose the southern leg of the loop for analysis because the sources travelled further along this leg, resulting in better defined height values. The position of the source was represented by the peak of a two-dimensional Gaussian fit to the southern CLEAN source, which was isolated by removing all flux lower than 30% of the brightest pixel. Height was then defined as the distance from the southern footpoint to the position of the source along a curve that passed through these two points as well as the northern footpoint (Figure 4.4 inset).

The footpoints were defined as the peaks of the X-ray sources in the 20–50 keV range at the HXR peak of the flare (Figure 4.2 c, orange contour). The location of these footpoints was used as a reference point for the heights of the low-energy sources. These relative heights were then converted to absolute ones by adding the predicted height of peak 25–50 keV emission, based on the CTTM (see Sect. 3). This analysis was repeated for all three energy bands used to create images.

With the evolution of the source height for each energy band quantified as a function of time and values of the nonthermal power-law index derived from spectra, we then compared the RHESSI observations directly to predicted height-time evolution based on the thick target model.
4.2 RHESSI Observations

Figure 4.4: Height of X-ray source peak with time for the 3–6, 6–8, and 8–10 keV energy bands. Vertical lines represent the one-sigma width of the two-dimensional Gaussian which was fit to the RHESSI source to determine peak location. They illustrate the size of the source, which is sensitive to the point spread function (PSF) of the instrument. Height is defined as distance in megametres from the source peak to the southern footpoint along the circle defined by the source peak position itself and both footpoints (see inset). Footpoints are defined as the peak position of 25-50 keV emission at 04:36:08–04:36:12 UT, the peak in HXRs. The temporal spacing of the data points here does not represent the integration time of the associated RHESSI images. For all images but the first, the integration time is 8 s, while the spacing between them is 4 s, resulting in an overlap of 4 s. Inset: An example image of 3–6 keV emission at 04:35:40–04:35:48 UT. The source, just prior to splitting into two, can be seen to the right of the image, at the assumed looptop. Overlaid on the image are locations of the peaks of Gaussian fits to the current descending source (open square) and the 20–50 keV footpoints seen at the HXR peak (filled diamonds). The definition of height is visualised as the distance along the circle between the southern footpoint and the southern source.
4.3 Thick Target Modelling

In this Section we outline the method by which a model nonthermal X-ray source height is calculated for a given injected spectral index, $\delta$, and photon energy, $\epsilon$. This builds upon the CTTM which was introduced in Section 2.3.1. A power-law electron injection spectrum describes the distribution of electrons with their kinetic energy, $E_0$, before any interaction with coronal or chromospheric plasma. It has the form $f_0(E_0) = (\delta - 1) (f_1/E_1) (E_0/E_1)^{-\delta}$, where $E_1$ and $f_1$ constitute reference points in electron flux and energy. Following acceleration, electrons travel down the flare loop and undergo Coulomb collisions with the ambient plasma, reducing their energy from $E_0$ to $E$. Thus, at a given distance, $z$, measured from the footpoint along the loop, the spectrum becomes $f(E, N(z))$, where $N(z) = -\int n(z) dz$ is the column depth and $n(z)$ is the number density of the ambient plasma \cite{Brown1972}. The energy lost to collisions is given by $E^2 = E_0^2 - 2KN$, \cite{Brown1972}, where $K = 2\pi e^4 \Lambda$ and $\Lambda$ is the Coulomb logarithm for an ionised plasma, which is used here as observed emission originates from heights at which the solar atmosphere is well ionised \cite{Brown1975, Emslie1978}.

In this work the goal is to determine the peak location of nonthermal X-ray emission by exploring different density models and injection spectral indices. \cite{Brown2002} derived this distribution of nonthermal photon flux with height as

$$\frac{dI}{dz} = \frac{A f_1 \sigma_0}{8\pi r^2 E_1} (\delta - 1) \frac{1}{\epsilon} n(z) \left( \frac{E_1^2}{2KN(z)} \right)^{\delta/2} B \left( \frac{1}{1 + u(z)}, \frac{\delta}{2}, \frac{1}{2} \right), \quad (4.1)$$

where $r$ is the distance from source to observer, $A$ is the cross-sectional area of
4.3 Thick Target Modelling

the loop, \( u(z) = \epsilon^2 / 2KN(z) \) and \( B(...) \) is the Incomplete Beta Function,

\[
B \left( \frac{1}{1 + u}, \frac{\delta}{2}, 1 \right) = \int_0^{1/(1+u)} x^{\delta/2-1} (1 - x)^{-1/2} dx. \tag{4.2}
\]

As we are only interested in the height of peak HXR flux, for neatness we hereafter remove the constant factor \( \alpha = Af_1\sigma_0/(8\pi r^2E_r) \) and express the distribution as \( (dI/dz)^* = (dI/dz)/\alpha \). From Equation 4.2 we therefore obtain

\[
\left( \frac{dI}{dz} \right)^* = (\delta - 1) \frac{1}{\epsilon} n(z) \left( \frac{E_1^2}{2KN(z)} \right)^{\delta/2} B \left( \frac{1}{1 + u(z)}, \frac{\delta}{2}, \frac{1}{2} \right). \tag{4.3}
\]

To evaluate Equation 4.3, a model providing density \( n(z) \) and column depth \( N(z) \), which are related for all \( z \) by \( n(z) = -dN/dz \), is required. Using this relationship one can say \( n(z) = -Nd(logN)/dz \) and define \( H(N) \equiv -1/d(logN)/dz \) the local scale height, such that \( n(z) = N(z)/H(N(z)) \). Thus a depth-varying scale height is implemented through the choice of \( H(N) \). In this work the chosen model for scale height is \( H(N) = H_r(N_r/N)^a \), where \( H_r \) and \( N_r \) are reference scale heights and column depths, which along with \( a \) can be varied freely, where \( a > 0 \). While this model is described by three variable parameters as presented, we note that one of these parameters can be set constant. As \( H_r \) always appears in the factor \( H_rN_r^a \), it will be left fixed at the constant value of \( 10^9 \) cm, while \( N_r \) and \( a \) are allowed to vary. To constrain the model, limits can be set on \( n(z) \) and \( H(N) \) based on previously measured and physically expected values for the low solar atmosphere.
Following this choice of $H(N)$, Equation (4.3) becomes

$$
\left( \frac{dI}{dz} \right)^* = \frac{(\delta - 1)}{H_r N_r^a} \frac{\epsilon}{\epsilon} \left( \frac{E_i^2}{2K} \right)^{\delta/2} N^{1+a-\delta/2} B \left( \frac{1}{1+\frac{a}{2}}, \frac{\delta}{2} \right),
$$

which we can now use to produce a plot of $dI/dz$ versus $z$ (see Figure 4.5), from which the height $z_{\text{max}}$ at peak $dI/dz$ can be calculated. In order to convert from a column depth to a height in the solar atmosphere, the relation $n(z) = -dN/dz = N/H(N) = N^{1+a}/(H_r N_r^a)$ was used to form a differential equation, integration of which between the limits $N = N(z = z_{\text{max}}) = N_{\text{max}}$ and $N = 0$ then gives

$$
z_{\text{max}} = \frac{H_r}{a \left( N_r/N_{\text{max}} \right)^a},
$$

which gives an absolute height of the model nonthermal source. The observed source heights, however, are measured as distance above the 25–50 keV footpoint. Therefore, the model height of the 25–50 keV footpoint is calculated and added to the observed values before comparison is made. We note that this reference height is relatively small, roughly 0.25 Mm.

The model $(dI/dz)^*$ distribution for nonthermal emission of 7 keV photons is shown in Figure 4.5 at four different electron spectral indices. The position of peak emission is highlighted, illustrating the result that harder injection spectra (lower index values) result in a lower location of peak emission. This reflects the fact that if electrons are accelerated to higher kinetic energies, they will propagate further into the coronal and chromospheric plasma, radiating bremsstrahlung by short range electron-ion interactions, before losing their energy to long-range Coulomb interactions.

This model will be used to produce expected nonthermal source heights based
4.3 Thick Target Modelling

Figure 4.5: Modelled HXR flux distribution with distance from the footpoint along the flare loop, produced using Equations (4.4) and (4.5) and assuming example density parameters $H_r = 10^9\text{cm}$, $N_r = 2.43 \times 10^{19}\text{ cm}^{-2}$, and $a = 0.9$. Four sample spectral index values are input at a photon energy of 7 keV. Since z- and $\delta$- independent factors are neglected, the distributions are normalised such that they peak at 1; however, the location of the peaks and the relative scaling between plots of different index are accurate. The height at which $dI/dz$ distributions are at their maximum ($z_{\text{max}}$) is noted, illustrating the HXR source height’s dependence on spectral index. For lower spectral indices, the height at which $dI/dz$ is at its maximum value is lower in the model flare loop.
on the density model, spectral index, and photon energy. These CTTM-based heights of peak photon flux can be fitted to those recorded by RHESSI, using the electron spectral index evolution, $\delta(t)$, provided by the fits to RHESSI spectra. A close match between the observed and modelled source heights will indicate whether or not the CTTM prediction of source descent is a possible interpretation of the observed motions in this flare. Additionally, the density model required by the fit can be compared with previous observations given in other work to determine if the densities required to produce this result are commonly encountered in flaring plasma.

### 4.4 Results

The comparison between model and observed height-time evolution is shown in Figure 4.6. While we determined heights earlier and later in the flare, only the portion of the height-time evolution that was fitted with our model is shown. We did not use the first time interval because images and spectra were noise-dominated, and so both the measured height value and spectral index were inaccurate. Data after 04:36:00UT have also been neglected from the fitting algorithm, as it is believed here that the emission becomes predominantly thermal at 3–10 keV and is thus not expected to be predictable based on a nonthermal electron flux model. Vertical lines at each data point represent one-sigma widths of the Gaussians that were fitted to the RHESSI sources, which remained roughly 2–3 Mm, corresponding to the RHESSI PSF’s half width at half maximum.

An initial observation of importance is the distribution in height for the three energies before the HXR peak at 04:36:12 UT. The 8–10 keV source is located
4.4 Results

Figure 4.6: (a) Electron spectral index, based on fits to RHESSI spectra. (b–d) Model and observed nonthermal source height evolutions for photon energies of 3–6 keV, 6–8 keV and 8–10 keV respectively. Source heights derived from RHESSI observations are denoted by diamonds, with vertical solid lines indicating the one-sigma width of the Gaussian which was fitted to the X-ray source. The heights corresponding to the peak in model intensity are shown as a solid line. The shaded gray area extending above and below the solid line represents a one-sigma width of the model intensity distribution, calculated using its full width at half maximum, where FWHM = 2.35σ. Two alternate model height evolutions are shown as dashed and dotted lines, which use different fit parameters $N_e$, $H_e$, and $a$. Along with the model given by the solid line, these all produce minimal $\chi^2$ values when fit to the data. The alternative results are presented to show the range of possible fits to the data, with corresponding density models shown in Figure 4.7. The dashed horizontal line represents the absolute height of the 25–50 keV footpoint, approximately 0.24 Mm. Fit parameters of the solid line are shown in (d).
lower in the loop than the 6–8 keV source, which likewise is lower than the 3–6 keV source. This distribution does not hold for the full duration of the flare; there is a reversal at the HXR peak of the flare (Figure 4.4). In the nonthermal scenario, a flux of nonthermal electrons travels through an increasingly dense chromospheric plasma. Higher energy electrons are stopped by higher densities, and so the fastest electrons will penetrate deeper before their bremsstrahlung emission peaks. Thus, in this regime, high-energy emission is expected to be located lower in the loop than low-energy emission. However, for thermal emission, the reverse is true if magnetic reconnection is progressing above the loop. In this scenario, the upper loops are newly reconnected and hotter, while the plasma underneath has had time to cool, leading to the highest energy thermal emission being located nearer the looptop (Tsuneta et al., 1992). With this in mind, the imaging analysis performed here suggests that the tracked X-ray emission in the 3 - 10 keV energy band is nonthermal until the HXR peak, at which point thermal emission becomes dominant as the sources appear to rise. This is consistent with the spectroscopic results and furthermore suggests that the tracked emission before the HXR peak can be treated as nonthermal. As a result, it is during this phase that the CTTM can be fitted to the data.

As shown in Figure 4.6 b–c, for suitably chosen scale height parameters, a model source descent can be simulated which shows strong agreement with observation. Three source descents are shown, one for each of the three energy bands used for imaging. The 3–6 keV source first appears ∼15 Mm above the footpoint, while the 6–8 and 8–10 keV emission appears at ∼12 and ∼5 Mm respectively. All three sources then descend to reach approximately the same height, ∼5 Mm, above the footpoint, which we previously defined by the location
of the southern 20–50 keV source at the HXR peak. The early 6–8 keV emission appears to originate higher up, likely due to the contribution by the thermally–
generated iron line complex to this energy band. This difference in apparent rate of descent of the nonthermal source was predicted in the CTTM through the use of a depth-varying hydrostatic scale height, an essential part of the density model used in the fitting process.

This density model is summarised in Figure 4.7. To help constrain the fit parameters, we ensured that the resulting density and scale height models agreed reasonably with previous observations (e.g. Aschwanden et al. 2002; Liu et al. 2006).

### 4.5 Discussion and Conclusion

In order to treat the observed source motion with the CTTM, we first needed to establish energies at which emission could be considered nonthermal. Spectroscopic analysis suggests that, prior to the HXR peak, emission is predominantly nonthermal above 7 keV and contributed to by both thermal and nonthermal components below that energy. As the flare progresses, the lowest energy bands become dominated by thermal emission. This can be explained by heating of the plasma in the flare loop from 8 MK to 11 MK within 15 s, as derived from Geostationary Operational Environmental Satellite (GOES) observations, using the background subtraction method outlined in Ryan et al. (2012). As the plasma reaches greater temperatures, it emits thermal radiation at higher energies. This heating period is expected to take place in all flares; however, in this case it was gradual enough so that we could make a significant number of lower–energy
4. CORONAL HARD X-RAY SOURCE RESPONSE TO ELECTRON SPECTRAL HARDENING

Figure 4.7: Density profile required to obtain the fits shown in Figure 4.6. Density \(n(z)\) and local scale height \(H(z)\) versus height, \(z\), above the footpoint are shown. The input parameters \(H_r\), \(N_r\), and \(a\) are the best-fit parameters resulting from the fitting process outlined in Figure 4.6. Two alternate density models, which are derived from the alternate fits shown in Figure 4.6, are given here as dashed and dotted lines. The shaded region represents the range of heights within which observations of HXR sources were made, and so densities and scale heights outside of this range are not expected to be accurate. It was assumed that the density structure of the flare plasma was approximately constant over the \(\sim 20\) s rise phase of the HXRs. Vertical solid lines indicate the location of peak emission for the denoted energies, which represent the three energy bands used in this study. The values shown correspond to the first observation, where the injection spectral index is \(\delta = 5\).
4.5 Discussion and Conclusion

HXR detections. Therefore, we deemed it appropriate to analyse observed source motions based on the CTTM.

We obtained a close match between model and observed X-ray source heights in this work (Figure 4.6); however, many important assumptions were made in order to do so, including that of a model density structure. Densities ranging from $10^{11}$ to $10^{13}$ cm$^{-3}$ over 20 Mm within a flare loop have been observed in previous RHESSI studies (e.g. Liu et al., 2006). The required density distributions in this work show similar structure and are also in line with derived densities of Aschwanden et al. (2002). An interesting requirement for this analysis was the introduction of a depth-varying scale height, which is responsible for the difference in apparent descent rate of the nonthermal emission between different photon energies. Close to $z = 0$, the required scale height is on the order of $10^7$ cm, or hundreds of kilometers, in agreement with previous RHESSI-based calculations of $\sim 130$–140 km (Kontar et al., 2008b; Saint-Hilaire et al., 2010), as well as with scale heights derived from temperatures put forward by modelling of visible and UV emission (Vernazza et al., 1981). The latter work, as well as that laid out by Allred et al. (2005), also suggest coronal temperatures consistent with a density scale height on the order of a number of megametres, as was also required by this fit. Without a variation in scale height (and thus temperature), the distance between sources of different energy would be constant, contrary to observations of this event.

In the process of modelling the distribution of nonthermal emission with height, we have shown theoretically that a strong asymmetry should be present in observed sources. As CLEAN was used to reproduce the images, this asymmetry may have been diminished and the peak of the model distribution therefore could
4. CORONAL HARD X-RAY SOURCE RESPONSE TO ELECTRON SPECTRAL HARDENING

have been shifted by the reconstruction process. To test this, we ran the model intensity distributions through a one-dimensional version of CLEAN. Some loss of the asymmetry of the source was seen for a small number of iterations (100), with the resulting distribution approaching a Gaussian shape. This may explain the near-Gaussian shape of the sources in RHESSI images. The resulting peak position was seen to shift by at most \( \sim 1 \) Mm in the case of the model used. Finally, the presence of a low energy cutoff substantially higher than the photon energy would have the effect of removing the low-energy electrons that contribute most to the “tail” of the asymmetric model source. In this way, a cutoff could also explain the observation of symmetric sources.

The model used in this work relied on the assumption that the thick target model is accurate and that the density structure of the target is the dominating factor on an X-ray source position. We note that other relevant mechanisms have been discussed but were not taken into account here. One could consider pitch-angle diffusion where, immediately following energy release, electron flux exhibits a large pitch angle to the magnetic field and is thus contained to the higher parts of the loop (Fletcher, 1997). Over time, diffusion causes a lowering in pitch angle and most of the accelerated electrons move gradually further down the loop, which may contribute to a source descent. Another important consideration concerns the evolution of \( n(z) \) with time. As electrons are accelerated in the flare loop, they cause heating and expansion. This results in a redistribution of local plasma density, which should lead to a predicted change in height of nonthermal HXR sources. This would work against mechanisms which cause a descent in HXR emission.

Battaglia et al. (2012) make use of Fokker-Planck modelling to determine the
4.5 Discussion and Conclusion

degree by which various mechanisms displace peak heights from their location as determined by collisional effects alone. They find that overall displacements of \( \sim 10\% \) in source position can be caused by magnetic mirroring and the implementation of a non-uniformly ionised flare loop, while pitch-angle scattering can cause a more stark displacement of up to 20\%. It would therefore be important to allow for these effects in a complete model.

Keeping these remarks in mind, we have shown here that the hardening of the electron injection spectrum is, with suitably chosen model densities and injection spectrum, sufficient to drive downward motion of nonthermal X-ray sources during the hardening of the injection spectrum. This model requires a flux of electrons from the looptop, or at least from \( \sim 20 \text{ Mm} \) above the footpoint towards the footpoints of the flare loop. Models invoking torsional Alfvén waves as the mechanism of primary energy transfer from the corona (Fletcher & Hudson 2008) or cascading reconnection in the chromosphere with re-acceleration there (Brown et al. 2009) still offer no explanation of such a relationship between spectral index and HXR source heights. However, following further development of these relatively new models, early impulsive events such as this will be useful in testing these predictions of nonthermal source behaviour before the peak in HXRs.
4. CORONAL HARD X-RAY SOURCE RESPONSE TO ELECTRON SPECTRAL HARDENING
Hard X-ray Source Sizes in a Beam-Heated and Ionised Chromosphere

Solar flare hard X-rays (HXRs) are produced as bremsstrahlung when an accelerated population of electrons interacts with the dense chromospheric plasma. HXR observations presented by Kontar et al. (2010) using the Ramaty High-Energy Solar Spectroscopic Imager (RHESSI) have shown that HXR source sizes are 3–6 times more extended in height than those predicted by the standard collisional thick target model (CTTM). Several possible explanations have been put forward including the multi-threaded nature of flare loops, pitch–angle scattering, and magnetic mirroring. However, the nonuniform ionisation (NUI) structure along the path of the electron beam has not been fully explored as a solution to this problem. Ionised plasma is known to be less effective at producing nonthermal bremsstrahlung HXRs when compared to less ionised plasma. If the peak HXR emission were produced in a locally ionised region within the chromosphere, the intensity of emission will be preferentially reduced around this peak, resulting
in a more extended source. Due to this effect, along with the associated density enhancement in the upper chromosphere, injection of a beam of electrons into a partially ionised plasma should result in a HXR source which is substantially more vertically extended relative to that for a neutral target. Here we present the results of a modification to the CTTM which takes into account both a localised form of chromospheric NUI and an increased target density. We find 50 keV HXR source widths, with and without the inclusion of a locally ionised region, of \( \sim 3 \) Mm and \( \sim 0.7 \) Mm, respectively. This helps to provide a theoretical solution to the currently open question of overly-extended HXR sources.

This work has been published in the Astrophysical Journal (O’Flannagain et al. 2015). The modelling and original concept for the work outlined in this Chapter was provided by the first author, while the co-authors provided guidance and context.
5.1 Introduction

While RHESSI observations have resulted in well-established spatial HXR properties in terms of centroid locations, the vertical extent of a HXR footpoint source is still found to be substantially underestimated by the Collisional Thick Target Model (CTTM; Brown 1971, Hudson 1972). Indeed, a detailed RHESSI study performed by Kontar et al. (2010) aimed to measure the sub-arcsecond spatial properties of one particular solar flare that occurred on 6 January 2004. In that work, the visibility forward fit (VIS FWDFIT) method was used, which by fitting model visibilities of pre-defined Gaussian sources to observed RHESSI visibilities, can produce moments of the HXR distribution with a greater accuracy than RHESSI’s finest grid pitch (Schwartz et al., 2002). With this method, it was shown that observed vertical source extents were in the range of 3–6 arcsec (∼2–4 Mm), depending on energy. This was a consistent factor of 3–6 times more extended than their simulated counterparts (see also Dennis & Pernak, 2009). This presents a clear disconnect between observation and the most commonly-used model for solar flares.

The standard model of solar flares describes an energy release in the corona, believed to be via magnetic reconnection (Petschek, 1964, Sweet, 1958). Particle acceleration results from either the strong electric fields associated with the changing magnetic field resulting from reconnection, or a secondary mechanism such as shocks or turbulence formed around the diffusion region (see Zharkova et al. (2011) for a recent review). These accelerated particles travel down the reconnected magnetic field lines into the dense chromosphere, where they lose their energy to Coulomb collisions and emit bremsstrahlung (CSHKP model;
5. HARD X-RAY SOURCE SIZES IN A BEAM-HEATED AND IONISED CHROMOSPHERE

Carmichael, 1964; Hirayama, 1974; Kopp & Pneuman, 1976; Sturrock, 1966). The CTTM describes the relationship between the HXR spectrum produced and the injected spectrum of electrons. It has also been used to make predictions on the nonthermal HXR footpoint vertical intensity distribution (Brown & McClymont, 1975; Brown et al., 2002), and so can produce estimates of the vertical extent of chromospheric HXR sources for a given density structure along the flare loop. In order to reproduce sources which are 3–6 times larger than the standard CTTM predicts, as are observed by RHESSI, one could alter the density structure until the observed extent is reached. However, Battaglia et al. (2012) have shown that this would require a strong enhancement of coronal density. This was shown to result in electrons of energy as high as $\sim 20$ keV being stopped in the corona, failing to produce chromospheric footpoints. Additionally, the coronal densities required for a substantial increase in source size are larger than $\sim 10^{11}$ cm$^{-3}$, which were noted to be atypically high.

A further category of density models, which are based on simulating the response of a quiet-Sun plasma to an injection of nonthermal electrons, may provide a solution to this problem. Theoretical work invoking full radiative and hydrodynamic models shows that upon heating by an electron beam, a portion of the chromosphere rapidly becomes ionised (Abbett & Hawley, 1999; Allred et al., 2005; Fisher et al., 1985). It has been previously shown that variation in the ionisation fraction along the path of the beam, or non-uniform ionisation (NUI) should produce a break in the HXR spectrum, as ionised plasma is a less efficient bremsstrahlung target (Brown, 1972; Su et al., 2009). It has been demonstrated that a rise or fall in ion fraction across the transition region will cause a variation in the vertical structure of HXR emission (Emslie, 1978). However, the effect
of a local variation of ion fraction within the chromospheric target has not been explored.

In this work, we demonstrate how this reduction of HXR production efficiency in a locally ionised chromosphere, together with the increased density in the upper chromosphere can result in a substantial increase in the vertical extent of the emitted source size. We discuss this increase in the context of RHESSI’s finest spatial observations. We also briefly demonstrate the effect this local ionisation has on the RHESSI spectrum, as this may contribute to the break normally attributed to the variation in ionisation profile encountered in the transition region.

5.2 Method

In this Section, we outline the details of expressions for HXR distributions in height (profiles; Section 2.1) and energy (spectra; Section 2.2) produced by a beam of nonthermal electrons injected into a plasma of freely-varying ionisation fraction and temperature. For the injected distribution $f_0$ of electron energies $E_0$, we will take the standard power-law form without a low-energy cutoff:

$$f_0(E_0) = (\delta - 1) \frac{f_1}{E_1} \left( \frac{E_0}{E_1} \right)^{-\delta}$$

(5.1)

where $f_1$ and $E_1$ are a reference flux and energy, and $\delta$ is the spectral index. For the remainder of the Chapter, we will only consider HXR photons with energies greater than 20 keV, and will assume a low–energy cutoff below this value. As a low–energy cutoff will have no effect on emission with a greater photon energy, it is satisfactory to exclude one from our injected beam model.
5. HARD X-RAY SOURCE SIZES IN A BEAM-HEATED AND IONISED CHROMOSPHERE

5.2.1 Target Density Model

In this work, a modified form of the standard hydrostatic equilibrium (HSE) density structure is used. The standard form describes an exponential drop in density with height, $z$, above the photosphere: $n(z) = n_{ph} \exp(-z/H)$, where $n_{ph}$ is the density at the top of the photosphere, and $H$ is the scale height. The scale height in this case assumes a constant temperature $T$ and ionisation fraction $X$ and has the form $H = k_B T / [(X + 1)m_p g]$, where $g$ is the acceleration due to gravity at the solar surface. However, to take into account an arbitrary variation in temperature and ionisation fraction, we adopt a heuristic form for the exponent change such that

$$n(z) = n_{ph} \exp \left[ -\frac{m_p g}{k_B} \int_0^z \frac{dz}{(X(z) + 1)T(z)} \right] \quad (5.2)$$

where $k_B$ and $m_p$ are Boltzmann’s constant and the proton mass, respectively. In this case we have made the assumption that the dominant term in the expansion of $d(n(X+1)T)/dn$ is $(X+1)T$. This assumption may not hold in all cases for the solar atmosphere, but for the purposes of producing a density enhancement, this assumption is found to be satisfactory at least as an empirical means of exploring the NUI effect. This expression for density takes into account the rapid variation in scale height both at the transition region and as a result of localised beam heating and ionisation.
5.2 Method

5.2.2 HXR Height Profile

The distribution of HXR emission along a flare loop leg in the chromosphere can be determined by taking into account the collisional energy loss of an electron of initial energy $E_0$ to an energy $E$. The electron loses energy as it travels to a column depth $N(z)$, which is the integral of target density, $n(z)$ along the path of the electron: $N(z) = - \int_z^\infty n(z)dz$. This energy loss is given by $E_0^2 - E^2 = 2KN$, where $K = 2\pi e^4 \Lambda$, and $\Lambda$ is the Coulomb logarithm, which accounts for the range of collisional impact factors that result in significant energy loss, and so depends on the ionisation state of the plasma (Brown, 1972).

To arrive at a HXR spectrum of photons of energy $\epsilon$, this collisionally modified electron distribution, $f(E)$, is multiplied by Kramers bremsstrahlung cross-section $\sigma(\epsilon, E) = \sigma_0/(\epsilon E)$ (Kramers, 1923). Brown et al. (2002) derived the distribution of nonthermal photon flux with height to be

$$\frac{dI}{dz} = \frac{Af_1\sigma_0}{8\pi r^2E_1} \left(\delta - 1\right) \frac{1}{\epsilon} n(z) \left(\frac{E_1^2}{2KN(z)}\right)^{\delta/2} \times B\left(1 + u(z); \frac{\delta}{2}, \frac{1}{2}\right),$$

where $A$ is the cross-sectional area of the loop, $r$ is the distance to the observer, $u(z) = \epsilon^2/2KN(z)$ and $B(\ldots)$ is the Incomplete Beta Function

$$B\left(1 + u(z); \frac{\delta}{2}, \frac{1}{2}\right) = \int_0^{1/(1+u)} x^{\delta/2-1} (1 - x)^{-1/2} dx.$$  

In order to account for variations in ionisation fraction $X$, we follow Brown et al. (1998) in defining an effective column depth $M$ such that our energy loss
5. HARD X-RAY SOURCE SIZES IN A BEAM-HEATED AND IONISED CHROMOSPHERE

is

\[ E_0^2 - E^2 = 2K'M \]  \hspace{1cm} (5.5)

\[ M = \int_0^N (\lambda + X(N'))dN' \]  \hspace{1cm} (5.6)

where \( K' = 2\pi e^4 \Lambda, \Lambda = \Lambda_{ee} - \Lambda_{eH}, \) and \( \lambda = \Lambda_{eH}/\Lambda \approx 0.55 \) (Brown, 1973). By replacing the \( 2KN \) factors in Equation 5.3 with \( 2K'M \), we can produce a model HXR height profile which takes into account the variation in ionisation fraction both at the transition region and within the target.

5.2.3 HXR Spectrum

The predicted HXR spectrum resulting from an electron beam passing through plasma which contains a step-function in ionisation fraction at the transition region has been given in a number of works (Brown, 1973; Brown et al., 1998; Su et al., 2011). This work was expanded upon to account for a region of linear variation from a fully ionised to a fully neutral target by Su et al. (2009). Here we present a further expansion which allows for any possible variation in ionisation fraction.

5.3 Generalised HXR NUI Spectrum

Following the work listed above we have the general form for distribution of HXR flux with energy, assuming again that the Kramers bremsstrahlung cross-section is adequate:

\[ J(\epsilon) = \frac{1}{4\pi r^2} \frac{\sigma_0}{\epsilon K'} \int_\epsilon^\infty \int_\epsilon^\infty \frac{f_0(E_0)}{\lambda + X(M)} dE_0 dE \]  \hspace{1cm} (5.7)
where all symbols have their previous meaning. We introduce an ionisation fraction $X(M)$ which will vary arbitrarily over $L$ linear steps:

$$X(M) = \begin{cases} 
X_0 : 0 \leq M < M_1 \\
X_1 : M_1 \leq M < M_2 \\
\vdots \\
X_L : M_L \leq M < M_{L+1}
\end{cases} \quad (5.8)$$

As the limit of the inner integral of Equation 5.7 is a function of $X(M)$, we take into account variation of $X(M)$ over $L$ linear steps by breaking up the integral into $L$ terms.

$$J(\epsilon) = \frac{1}{4\pi r^2} \sigma_0 \int_{\epsilon}^{\infty} \left[ \int_{E}^{\sqrt{E_0^2 + 2K'M_1}} \frac{f_0'(E_0')}{\lambda + X_0} dE_0' \\
+ \int_{\sqrt{E_0^2 + 2K'M_1}}^{\sqrt{E_0^2 + 2K'M_2}} \frac{f_0'(E_0')}{\lambda + X_1} dE_0' \\
\cdots + \int_{\sqrt{E_0^2 + 2K'M_L}}^{\infty} \frac{f_0'(E_0')}{\lambda + X_L} dE_0' \right] dE \quad (5.9)$$

We have essentially arrived at Equation 4 of Su et al. (2009), however in this case we have accounted for an arbitrary number of linear steps in $X(M)$ rather than one. This leads us to the following general expression of the HXR spectrum, following the substitution of our power-law injection spectrum, given at the beginning of this section.
5. HARD X-RAY SOURCE SIZES IN A BEAM-HEATED AND IONISED CHROMOSPHERE

\[ J(\epsilon) = \frac{f_I}{E_1^{1-\delta}} \frac{1}{4\pi r^2} \frac{\sigma_0}{\epsilon K'} \left( \frac{E^2-\delta}{(2-\delta)(\lambda + X_0)} \right) + \sum_{l=1}^{L} \left( \frac{1}{2(\lambda + X_{l-1})} (2K'M_l)^{-\delta+1/2} B \left( \frac{1}{1 + \frac{e^2}{2K'M_l}}, \frac{\delta}{2}, \frac{1}{2} \right) \right) \]

This expression can now be used to produce a HXR spectrum for any given variation in ionisation fraction along the path of the nonthermal beam, assuming a finite number of linear steps is an acceptable approximation of this variation. We are now able to produce both model HXR height profiles and spectra for any input ionisation fraction \( X(M) \).

5.4 Results

5.4.1 HXR Height Profile

The parameters of our target plasma, as well as a sample HXR height profile for photons of 50 keV are shown in Figure 5.1. Black lines in all cases describe a fully ionised corona and fully neutral chromosphere. Red lines show the atmospheric parameters and HXR fluxes when local heating and ionisation are taken into account using the model presented here. As shown in Figure 5.1 (a) and (b), we have approximated a localised beam energy deposition with a rise in both temperature and ionisation fraction, which produce a density which is enhanced in the chromosphere, exhibiting a more gradual decrease with height than the
neutral case. As a result of the combination of this density enhancement and the reduction of HXR emission around the chromospheric ionised region, we see a strong increase in vertical extent of the 50 keV emission as shown in Figure 5.1 (c).

Vertical sizes of the model sources are determined by taking the FWHM of emission. This size ignores the various imaging limitations imposed by RHESSI’s response, such as its dynamic range (Hurford et al. 2002, see Section 3.2 for a complete treatment of these effects). As shown in Figure 5.1 (c), the inclusion of NUI and heating effects increases the extent of our source from 0.63 Mm to 2.0 Mm, an increase by a factor of 3.2 for a 50 keV source. This effect can be seen for energies between 20 keV and 200 keV upon examination of Figure 5.2.

A complete set of resulting HXR height profiles are given in Figure 5.2 (a), in the form of an image of HXR flux against both height and energy. That is, a vertical slice of Figure 5.2 (a) is a height profile as presented in Figure 5.1 (c). In Figure 5.2 (b) and (c), the vertical size and centroid height of model HXR source are shown, respectively. Again, the red lines correspond to the NUI case and solid to the neutral chromosphere case. Also shown in blue are the observed values for size and height as measured using RHESSI, presented in Kontar et al. (2010), the work which initially highlighted the disparity between observed and predicted source vertical extents.

It is immediately apparent that, above \( \sim 30 \) keV, a substantial increase in vertical extent results from the inclusion of NUI in the target plasma. The ratio of source FWHM in a NUI target to that of a neutral one reaches a peak of a factor of 3.3 at 40 keV, and slowly falls with increasing energy to a value of \( \sim 2.1 \) at 100 keV, and \( \sim 1.4 \) at 200 keV. For the energy range of \( \sim 30–70 \) keV,
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Figure 5.1: (a) Target ambient plasma temperature variation with height. The quiet-Sun case is shown in black, and the flaring target case is shown in red. In the latter case, a localised increase in temperature, peaking around 2 Mm, is present due to beam heating. The sudden rise in temperature around 2.5 Mm constitutes the transition region. (b) Target plasma density (solid lines) and ionisation fraction (dashed line) distributions with height. In the flaring scenario, the chromosphere has been ionised around the region of peak beam energy deposition which, along with the rise in temperature, causes an increase in scale height. This results in an enhancement of density in the upper chromosphere. (c) HXR flux distribution versus height for photons of 50 keV. The black and red flux profiles correspond to the neutral and NUI target cases, respectively, and both profiles are divided by the value of peak emission in the neutral case. Also shown are the full-width-half-max (FWHM) of both profiles, and their ratio.
5.4 Results

Figure 5.2: (a) Distribution of HXRs in height and energy, in the case where nonuniform ionisation is taken into account. Emission is normalised to the maximum flux in its energy band. (b) Vertical size and (c) centroid height of the model sources versus energy, for the nonuniform (red) and fully neutral (black) case. Overplotted in blue are the observational values recorded in Kontar et al. (2010), as shown in Figure 6 of that paper.
the NUI-adjusted vertical extent matches those measured by RHESSI. Outside of this range, source sizes gradually return to that of the neutral case. This may be due to the fact that electrons emitting photons outside this energy range are penetrating to depths above and below the location of the peak in ionisation, and therefore result in standard neutral target emission. The source height on the other hand is much less strongly affected within the same energy range. As shown in Figure 5.2 (c), the height of the NUI source centroid is a factor of 1.6 and 1.5 higher than its neutral counterpart at 20 keV and 40 keV, and this disparity continues to diminish with increasing energy.

### 5.4.2 Instrumental Effects

A further important contribution to HXR source size which requires discussion is the effect of RHESSI’s imaging response. Due to the finite spacing of RHESSI’s finest grids, it has a lower resolution limit, which may result in an increase in the apparent size of small sources. This point-spread effect has previously been shown to be insufficient to completely account for the difference between predicted and observed source sizes \(^{(\text{Battaglia et al. 2012})}\). However, it is useful to determine if RHESSI would be able to detect the difference between the neutral and NUI HXR sources presented here.

Model two-dimensional HXR sources were produced by taking the matrix product of the model flare emission profiles and a normalised Gaussian with a FWHM of 3 arcsec. The resulting maps were then used as input to RHESSI simulation software developed by R. Schwartz \((\text{e.g., Sui et al. 2002})\). This software produces a RHESSI calibrated eventlist from an input simulated map \((\text{Schwartz})\).
et al. [2002]). From the eventlist, any of the available standard image reconstruction methods can then be used to produce an image. For our purposes, we use the VIS FWDFIT method [Schmahl et al. 2007]. VIS FWDFIT uses a characteristic shape for an X-ray source such as a circular or elliptical Gaussian, and fits the resulting simulated visibilities to those observed by RHESSI. The resulting Gaussian parameters, along with their standard deviations are returned as an output, and can be conveniently used to directly measure properties such as position and extent of the HXR source.

In order to accurately compare with the observations made by Kontar et al. [2010], we used the same set of detectors in the reproduction of RHESSI images from the synthetic input maps. These were detectors 2–7, as in that observation, detector 1 showed no significant signal, and detectors 8 and 9 were deemed too coarse for the event, exhibiting no modulation in their lightcurves. In addition, the input total counts parameter was chosen in order to match that of the flare which occurred on 6 January 2004, in the timeframe used to produce images in that study. This was estimated to be $7 \times 10^3$ counts detector$^{-1}$, as that was roughly the number of counts collected in each of the energy bands used in Kontar et al. [2010]. This selection was made in an effort to demonstrate that any distinction between sources emitted by a neutral and NUI target would be measurable by RHESSI in a real flare.

The result of running VIS FWDFIT on model 50 keV HXR sources emitted by a neutral and NUI target are shown in Figure 5.3. Input maps are shown in panels (a) and (c), and the maps resulting from running VIS FWDFIT on the calibrated eventlist produced by these input maps are shown in (b) and (d). As we are assuming an ideal case where the source is viewed side–on, the vertical
5. HARD X-RAY SOURCE SIZES IN A BEAM-HEATED AND IONISED CHROMOSPHERE

Figure 5.3: Demonstration of the effect of RHESSI’s imaging response to HXR sources produced by neutral and NUI target chromospheres. (a/c): A model HXR input map produced by taking the matrix product of a 3 arcsec FWHM Gaussian with the neutral/NUI HXR profile of 50 keV emission as shown in Figure 5.1 (c). The vertical line represents the western solar limb. The sharp dropoff in flux at the right of the source in (c) is due to the sharply–defined transition region, shown at ~2.4 Mm in Figure 5.1 (c). (b/d): The map produced by running the VIS FWDFIT routine on the calibrated eventlist resulting from the map shown in (a). This serves as an approximation of a RHESSI image given a HXR source produced in a neutral/NUI chromosphere.
line in these images denotes the point where the density function switches from photospheric to chromospheric, and so can be seen as the western limb. As shown, both sources are radially broadened to a degree by the simulated RHESSI imaging process. There still remains a clear distinction in source size between the neutral sources and those produced in a density–enhanced and NUI region for the realistic detector choice and countrate used. The FWHM of the minor axes of the neutral and NUI sources after instrumental effects are 1.6 and 2.3 arcsec, with standard deviations of 0.2 arcsec in both cases. It could be inferred that sources of FWHM <1.6 arcsec are extended up to this limit, but sources that are already larger are unaffected. This highlights the requirement for a physical mechanism to increase the vertical HXR extent in order to match observations. It is noted here that these measurements should be treated as lower limits only, due to the fact that the calibrated eventlists are produced using the same grid parameters that are then used to reproduce images (see for example, Battaglia et al., 2011).

5.4.3 HXR Spectrum

The HXR spectrum resulting from interaction between our power-law injected spectrum of electrons and the NUI plasma shown in Figure 5.1 is shown in Figure 5.4(a). As no low-energy cutoff has been included in this model, the resulting HXR photon spectrum appears to be a straightforward power-law spectrum with a spectral index that is constant with energy. However, the disparity is made clear in the rescaled spectra in 5.4(b). As shown upon examination of Figure 5.4(c), there is some variation in photon spectral index. As energy rises to 200 keV, the index of the spectrum rises from ~2.20 to ~2.27, a rise of ~0.07. This varia-
Figure 5.4: (a) Model HXR spectra resulting from the interaction of a power-law beam of electrons with a neutral (black) and nonuniformly ionised (red) target. (b) Same data as (a), but flux is multiplied by $\epsilon^{\delta-1}$ in order to more clearly show the distinction between the two cases. (c) The spectral index of the HXR spectra shown in (a), calculated by using the derivative of log(Flux) with respect to log(Photon Energy).
tion is qualitatively similar to that presented by Su et al. (2009), but in terms of the magnitude of the variation in index, is far smaller than the observed average rise of $\sim 0.3$–1.4. This may suggest that local ionisation within the chromosphere contributes some small amount to the known break in RHESSI spectra, but the major cause of the break which can be attributed to NUI would still be the variation in ionisation fraction across the transition region.

It should, however, be noted that there are methods of producing a break in the HXR spectrum which serve as an alternative to NUI. For example, incorrect background subtraction of the solar spectrum could result in a photon spectrum which does not relate to the accelerated electron spectrum, or pulse pile-up may not be corrected for adequately (Smith et al., 2002). Beyond instrumental effects, HXR albedo (Kontar & Jeffrey, 2010; Kontar et al., 2006), return currents (Holman, 2012; Knight & Sturrock, 1977), or the presence of a low-energy cutoff in the injected spectrum (Gan et al., 2002; Sui et al., 2007) may all also contribute to a break in the usual power-law HXR spectrum.

5.5 Conclusion

Here we have shown that the predicted local ionisation and heating of chromospheric plasma on the onset of a flare can produce an increase in the vertical HXR source extent by a factor of up to $\sim 3$ for commonly-imaged energies. This increase results from a combination of localised ionisation and a density enhancement in the upper chromosphere, but does not require any increase in coronal density, and therefore results in no extreme upward shift of source centroids. This helps to explain the observation of RHESSI source sizes that appear too large when com-
pared to predictions based on simple interpretations of the CTTM. In particular, the red curve in Figure 5.2 (b) can be directly compared with the measured values from Kontar et al. (2010), shown in blue. In both the model and observed cases, the HXR source size decreases with increasing energy, however model source sizes were shown to be a factor of 3–6 smaller in vertical extent than observations. By taking into account local NUI in addition to a chromospheric density enhancement, we have been able to account for most of this discrepancy which, while resulting in some increase in source height, does not encounter the problem of dramatically relocating HXR sources to the corona.

Previously, the disagreement between observed and modelled source sizes has been addressed by a number of mechanisms. Kontar et al. (2010) point out that treating the flare loop as a single monolithic loop is an oversimplification, and perhaps strands of different density structure could contribute to an apparent vertical lengthening of the source. Various physical processes expected to take place during flares, such as magnetic mirroring and pitch-angle scattering, have also been put forward as other contributing factors. However, Fokker-Planck modelling of the nonthermal electron distribution in a flare has demonstrated that these processes would have only minor effects on source size (Battaglia et al., 2012). The resulting conclusion has been that a threaded loop structure, wherein multiple smaller flare strands exist with different density profiles, remains the best explanation. While these processes are certainly still expected to take place, they may no longer be required in order to explain the observed discrepancy in vertical HXR source sizes.

It is important to clarify here the mechanism by which the NUI component of this model, aside from the density enhancement, produces a more vertically
extended source. In addition to causing an enhancement in the wings of a HXR source profile such as that shown in Figure 5.1 (c), the ionisation of the plasma also has the effect of reducing emission near the peak. By reducing the efficiency of emission primarily near this peak, the overall source becomes broader. Because, for lower photon energies, peak emission and peak ionisation both occur at the same position along the loop, it is expected that the reduction in HXR flux will always occur at the location of brightest HXR flux, and therefore always produce some extension. This suggests that the effects described here should be common, and possibly applicable to all flares.
5. HARD X-RAY SOURCE SIZES IN A BEAM-HEATED AND IONISED CHROMOSPHERE
Magnetic reconnection is believed to be the primary mechanism by which the non-potential magnetic energy stored in a coronal magnetic field is transferred into the energy rapidly released during solar eruptive events. Unfortunately, due to the fine spatial scales at which reconnection is thought to occur, it is not directly observable in the solar corona. However, larger-scale processes such as associated inflow and outflow, and signatures of accelerated particles have been put forward as evidence of reconnection. In this Chapter, we outline the analysis of such flows observed by the Atmospheric Imaging Assembly, which appear during the apparent collapse of a transequatorial coronal field structure. We also make an interpretation of the acceleration model based on associated Type I radio emission detected by the Nançay Radioheliograph. We identify a gradual (1 km/s) and fast (5 km/s) inflow phase, as well as a fast (30 km/s) and rapid (80–100 km/s) outflow phase, resulting in an estimated reconnection rate \( \frac{v_{in}}{v_A} \) of 0.05. We interpret
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the associated brightening and dimming of the radio emission as evidence for accelerated particles at the reconnection region responding to a gradual fall and rapid rise in electric drift velocity, in response to the inflowing and outflowing field lines. These results are presented as a comprehensive example of 3D null–point reconnection.

This work has been submitted to the Astrophysical Journal Letters (O’Flannagain et al. 2014b, in review). The analysis outlined in this Chapter was completed by the first author, while background on magnetic modelling and solar radio bursts was provided by the co–authors.
6.1 Introduction

Solar eruptive events (SEEs) are believed to be triggered by the release of non-potential magnetic energy in twisted coronal magnetic field structures. The primary mechanism proposed for the transfer of this stored energy into the accelerated particles, heated plasma and ejected matter that make up the SEE is thought to be magnetic reconnection. Proposed in the 1940s, and built upon by Sweet (1958), Parker (1957), and Petschek (1964), it has been demonstrated that, with a suitably small-scale diffusion region, reconnection can release enough energy to drive the SEE within a standard flare timeframe (Priest & Forbes 2000, see Section 2.1.5). Theory has been further developed to include breakup of the reconnecting current sheet into smaller magnetic islands, potentially allowing even greater rates (Kliem et al. 2000; Shibata & Tanuma 2001). Reconnection has been reproduced in laboratories (Yamada et al. 2010; Zhong et al. 2010) and been measured in the Earth’s magnetosphere (Deng & Matsumoto 2001; Øieroset et al. 2001).

However, as fast reconnection requires a diffusion region with a spatial scale on the order of the electron gyro-radius (Daughton & Roytershteyn 2012), it is not directly observable in the solar corona by modern instrumentation. Indirect observations are possible though as reconnection is thought to have certain signatures visible at larger spatial scales. For example, active region loops which appear to be flowing in towards and outwards from apparent X-points on the solar limb have been interpreted as reconnection inflow and outflow (Savage et al. 2012a; Su et al. 2013). Inflows have also been observed at the limb along with microwave emission, which were presented as evidence for particle acceleration at
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an X–point (Narukage et al. 2014). Separately, Type I radio noise storms have been explored as a signature of accelerated particles at potential reconnection sites near the interface between active regions (Willson 2002, 2005). While these storms are known to occur over very long timescales (e.g., Del Zanna et al. 2011), they have also been observed to react rapidly to flare and CME activity (Aurass et al. 1990; Iwai et al. 2012a; Kathiravan et al. 2007).

In order to address these disparate forms of both non–flaring and impulsive acceleration, we must take into account models which address the 3D nature of reconnection. 3D reconnection has been divided into a number of different categories, such as spine and fan reconnection (Dalla & Browning 2005; Priest & Titov 1996), null–point reconnection (e.g., Masson et al. 2009), and reconnection along quasi–separatrix layers (Demoulin et al. 1996, 1997). Simulations of these regions have shown to produce the accelerated electron distributions expected from flares (Baumann et al. 2013; Browning et al. 2010; Priest & Pontin 2009; Stanier et al. 2012). In particular, it has been shown that reconnection and acceleration are likely along 3D separators – the lines of intersection between separatrices, or surfaces of different magnetic connectivity (Gorbachev & Somov, 1989; Longcope et al. 2005; Parnell et al. 2010).

In this Chapter we present observations of a coronal X–shaped structure which appears near disk–centre on 6 July 2013. This structure persists for the full passage of the active region across the disk, and is closely accompanied by a persistent radio Type I noise storm. In particular, observations of a partial collapse of this structure are discussed, and the implications for coronal magnetic reconnection are explored.
6.2 Observations

The focus of this work is the evolution of a quadrupolar X-shaped coronal structure located to the north or north-west NOAA active region AR11785, identifiable in SDO/AIA’s coronal channels (94, 131, 171, 193, 211, and 335 Å) for its full passage across the solar disk, from 2 to 12 July 2013. For the majority of this time, a Type I radio noise storm is also observable by the Nançay Radioheliograph (NRH: Kerdraon & Delouis, 1997), consistently located above the X-shaped structure.

An overview of the approach of AR11785 is given in Figure 6.1. Two AIA

![Figure 6.1: Summary of the evolution of active region AR11785 during its approach towards disk centre. AIA 171 Å images are shown at times when the active region is (a) near the eastern limb, and (b) near disk centre. The location and polarity of the underlying line-of-sight magnetic field as detected by HMI are shown in black for negative polarity, and white for positive polarity. These represent contours of 80 G in the associated HMI images. Overlaid as blue circles are the location of Type I noise storm source centroids at all available frequencies from 150 to 445 MHz.](image)
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171 Å images are shown, at two different times during the passage of the active region towards disk centre, covering a total of three days. The X–shaped structure appears to begin forming in panel (a), but is most clear at disk centre (b). The centroids of the NRH radio noise storm sources – defined as all emission above 50% of the maximum for each image – are denoted by blue circles, with lighter shades of blue corresponding to higher frequencies of radio emission. As shown, higher frequency emission is consistently located closer to the apparent X–point.

As Type I noise storms are believed to be plasma emission (e.g., Iwai et al., 2012a; Willson & Groff, 2008), this would indicate that the higher frequency sources are originating from regions of higher density, and are therefore likely to be lower in altitude above the solar surface. This scenario fits the interpretation of Type I radio noise storms as a column of radially stratified dense plasma above an active region containing some population of accelerated particles (McLean, 1981). The linear nature of the centroid locations indicates a ‘line’ of such plasma, which appears to move from north to south over the passage of the active region, whilst still remaining ‘rooted’ at the X–point.

On 6 July, as the active region and accompanying coronal loop structure approached disk–centre, it appeared to undergo a partial collapse beginning at ~09:29 UT. This collapse was concurrent with an enhancement of NRH emission and a series of small C–class flares, detected by both the Ramaty High Energy Solar Spectroscopic Imager (RHESSI: Lin et al., 2002) and the Geostationary Operational Environmental Satellite (GOES), which measures spatially integrated solar SXRs in 1–12, and 3–25 keV energy bands. A summary of this collapse is shown in Figure 6.2. Prior to the collapse, four loop structures, connecting each opposite–polarity footpoint pair, are visible in the AIA 171 Å channel at
6.2 Observations

Figure 6.2: Overview of the collapse of the X–shaped coronal structure. Top: AIA 171 Å images covering the 90 minutes surrounding the collapse. Overlaid on the earliest image are the ‘slices’ used to generate the time–distance plots shown in Figures 6.4 and 6.5. Slices A and B are 67 arcseconds and 96 arcseconds long, respectively, with the arrow denoting the reference direction for the time–distance plots. Middle: NRH 430 MHz images showing the Type I storm source at the times of the above images. Overlaid is the field of view of the AIA images. Bottom: GOES 1–12 keV and RHESSI 3–6 keV lightcurves, showing the flares which occurred during the collapse.
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~09:29 UT. One hour later, after the collapse and flare have occurred, the northern and western segments of the X-shaped structure are no longer visible in the 171 Å image, and partially diminished in 193 Å.

The brightening NRH source (seen in 430 MHz images in Figure 6.2 third row) roughly maintained its overall position to the upper-left of the coronal X-point. Meanwhile, preliminary RHESSI imaging demonstrates that HXRs originated from the apparent lower-right ‘leg’ of the quadrupolar structure visible in AIA 171 Å. This indicates that energy release occurred during this time, perhaps as a result of the destabilisation of the coronal structure. However, as the purpose of this work is to identify the reconnection event itself, hereafter we will focus on the coronal location of the collapse and radio noise storm, first by estimating the magnetic field.

6.3 Data Analysis

6.3.1 Potential Field Extrapolation

In order to interpret these observations in the context of magnetic reconnection, the structure of the coronal magnetic field at the location of the active region must first be estimated. To do this we obtained potential field line extrapolations using the mpole suite of Interactive Data Language (IDL) programs (Longcope, 2001; Longcope & Klapper, 2002; Longcope & Magara, 2004). This piece of software simplifies the line-of-sight magnetic field, usually produced by a magnetogram such as those provided by HMI, into a number of positive and negative poles. Treating each region of radial field as magnetic point charges
enables us to define the 3D surfaces separating magnetic field lines of differing connectivity (separatrices) and the 3D lines of intersection between these surfaces (separators) \citep{Longcope1996}. The presence of the strong current sheets expected at these separators identify them as important locations of 3D reconnection \citep{Parnell2010}, occurring, for example, between an emerging and
an existing active region (Longcope et al. 2005). Therefore it could be expected that signatures of accelerated particles, such as the NRH emission observed in this work, could be located along these separators.

The magnetic field topology produced by these extrapolations is shown in Figure 6.3. As shown, the extrapolation indeed produces an X–shaped system of magnetic field, although of different detailed structure from that observed in AIA 171 Å. This topology appears to remain roughly constant throughout the collapse phase, despite the clear outflow and dimming of the observed western loops shown in Figure 6.2. During the extrapolation process, a large number of separators were produced (≈400–500 per time interval) and so for comparison only the separator which most closely passes the NRH sources (overplotted in blue) is shown for each time interval. However, the majority of separators did originate at the apparent X–point.

The relationship between the separator and the ‘line’ of NRH sources is of particular interest in this result. As shown, prior to the collapse, a separator does appear to pass through the area occupied by the NRH sources, although almost perpendicularly to their apparently linear formation. At the second time interval, shortly after the last X–ray flare occurs, the sources are gathered closer to the X–point, and are marginally better aligned with the separator. However, by the third time interval, 30 minutes later, the sources, with increasing frequency, are pointed directly towards the X–point, well–aligned with the separator originating from it.

As these extrapolations are produced using the assumption of a potential force–free field, it is expected that they will not produce an accurate model of magnetic field line geometry for solar active regions shortly before flares, as this
is when a large amount of non–potential energy is stored (e.g., Murray et al. 2013; Wiegelmann et al., 2005). It follows that if this non–potential energy is released to drive the observed flares, the coronal field structure should become more potential after energy release. This adds weight to the idea that the NRH sources lie along the ‘real’ 3D separator. As non–potential magnetic energy is lost from the active region, these NRH sources come to lie along the separator produced from the potential extrapolation as this is, following the flare, a more accurate model of the coronal field.

6.3.2 EUV and Radio Comparison

In order to characterise the inflow and outflow at the X–point, AIA 171 and 193 Å flux values were recorded along two curved ‘slices’, as shown in Figure 6.2. The flux along these slices is shown against time for the AIA 171 and 193 Å channels in the first two panels of Figures 6.4 and 6.5 respectively. As shown, there is evidence in both channels of a gradual inflow along slice A from ∼08:30 to 10:00 UT, followed in the 171 Å channel by a faster apparent inflow. Notably, the inflow seems to occur earlier in the 193 Å channel. From 09:40 to 10:10 UT, a much more rapid outflow is apparent along slice B, with a faster outflow period of finer loop strands appearing shortly after 10:00 UT. In both channels, the gradual inflow exhibited a velocity along the slice of ∼1 km/s, with the faster inflow in 171 Å of ∼5 km/s. The fast outflow shortly before 10:00 UT had a velocity of ∼30 km/s in both channels, with the rapid outflow reaching ∼80–100 km/s.

These velocities can be used to measure the magnetic reconnection rate in
this event. If we take the faster inflow as $V_{in}$ and the rapid outflow of finer loop strands as our $V_{out}$, we can estimate the rate as $M_A = V_{in}/V_A \approx V_{in}/V_{out}$, where $V_A$ is the Alfvén speed. Using this approximation, we have a reconnection rate of $M_A = 0.06$. This agrees well with the lower rate value reached in the event studied in [Su et al. (2013)].

Of particular interest here is the timing of the brightening and motion of the NRH radio sources during the time of this collapse. Shown in panel (c) of Figures 6.4 and 6.5 are the smoothed and normalized brightness temperatures for all of the NRH observable frequencies, excluding the two lowest. This is because these low–frequency sources were not localised to one point for the full duration of the collapse. The given brightness temperatures were acquired by averaging the pixel values over the source area, defined as the 50% contour of the maximum value for each image. The lightcurves are smoothed over 6 points for clarity, and normalised to their own maxima, which reached values of $\sim 10, 0.9, 0.4, 0.2, 0.07, 0.04, \text{and } 0.03 \times 10^9 \text{ K}$ for frequencies 228, 270, 298, 327, 408, 432 and 445 MHz, respectively. The extent of the NRH sources is shown in the final panel (d) of Figures 6.4 and 6.5. This property was produced in an attempt to estimate the length of the NRH source, across all frequencies, and so was simply produced by taking the distance between the centroid of the 432 and 298 MHz emission. Local sharp peaks in the plot are produced when images corresponding to one or both of these frequencies are briefly dominated by noise or instrumental artifacts.

As shown, during the gradual inflow phase of the collapse, all of the presented NRH channels exhibit a gradual decrease in relative brightness temperature, reaching a minimum around 9:40 UT. This is concurrent with a peak in the NRH extent, which represents a spreading of the NRH source frequencies further.
6.3 Data Analysis

**Figure 6.4:** Time–distance plots demonstrating the apparent plane–of–sky motion of brightening in the 171 Å channel along slice A (panel (a)) and slice B (panel (b)), which are shown in Figure 6.2. The horizontal dotted line denotes the location of the apparent coronal X–point, and vertical dashed lines correspond to the three time intervals used to produce the three panels in Figure 6.2. (c) NRH brightness temperature, averaged over the area within a 50% contour of each NRH image, for its 7 highest recorded frequencies, smoothed over 6 points and normalised to clarify local peaks. (d) NRH extent, measured as the linear distance between the centroid of the highest– and lowest– frequency stable sources, which were 432 and 298 MHz, respectively.
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Figure 6.5: Time–distance plots demonstrating the apparent plane–of–sky motion of brightening in the 193 Å channel along slice A (panel (a)) and slice B (panel (b)), which are shown in Figure 6.2. In contrast with the 171 Å emission, this higher energy emission appears to brighten earlier and closer to the apparent X–point. (c) and (d) NRH brightness temperature and NRH extent, as also shown in figure 6.4.
apart, and further to the solar north–east. This is followed by the simultaneous occurrence of the rapid AIA inflow/outflow, a sudden brightening in all NRH frequencies, and a decrease in the radio source extent. In particular, all NRH channels appear to peak in brightness sharply at 10:00 UT, at the same time as the minimum in source extent. This is followed by an overall peak in the NRH high-frequency channels, and a subsequent peak in the lower channels roughly 10 minutes later. The NRH emission then decays gradually to background levels, still localised to the same location near the X–point.

6.4 Interpretation

These observations can be better understood by examining the numerical results of the reconnection model outlined in Browning et al. (2010). In this work, 2D reconnection theories are built upon by examining particle acceleration at a 3D null point. The primary purpose is to characterise the acceleration for different particle species, and for varying electric and magnetic field strengths. An important result of this model is that for stronger magnetic fields, efficiency of particle acceleration was diminished. It is put forward that this is an effect of the electric drift speed having an inverse dependence on magnetic field – for stronger magnetic fields, a lower drift rate prevented ambient electrons from entering the acceleration region.

While it is also important to consider the effect of variation in the electric fields in the active region, our observational results provide strong evidence supporting the above interpretation. During the gradual inflow phase prior to 09:40 UT, the magnetic field is gathered around the null point with no corresponding outflow.
According to the above model, this increase in field strength would reduce the rate of null point acceleration, resulting in fewer accelerated particles available to produce nonthermal Type I emission. Conversely, the rapid outflow after 09:40 UT rapidly reduces the magnetic field strength around the null point, allowing for an equivalently rapid increase in the number of accelerated particles, and so a rapid increase in radio emission across all observed frequencies.

Indeed, the behaviour of the NRH source length also supports this picture. The ‘lengthening’ of the series of NRH sources before 09:40 UT could result from an overall increase of the density along the separator. For a power–law or exponential density structure, a uniform increase in density would shift sources at each frequency upwards, and further apart from one another. The rapid decrease in ambient density associated with the reconnection outflow after 09:40 UT would then have the opposite effect, rapidly decreasing the extent of the NRH emission.

It should of course be noted that these measurements were taken from plane–of–sky NRH centroids, and so could also have a contribution from 3D motion of the acceleration region – perhaps the separator itself.

\section*{6.5 Conclusion}

Here we have identified a solar radio noise storm which exhibits close consistent association with a quadrupolar transequatorial loop structure for its full passage across the disk. The interpretation of this noise storm as plasma emission produced indirectly by accelerated electrons at a 3D magnetic null point supports the idea that boundaries between interacting or trans–equatorial active regions are ideal locations for non–flaring reconnection due to their opposite magnetic po-
larity orientations, following Hale’s law \cite{Hale,1919,Pevtsov,2000,Sheeley,et,al.,1975,Tsuneta,1996}. This idea of non–flaring reconnection has previously been put forward as a solution to the coronal heating problem \cite{Parnell,2007,Priest,2003}.

Beyond this long–term relationship, we have investigated the behaviour of the NRH sources during a rapid collapse of the X–point structure, as observed in AIA. We have interpreted the variation in radio brightness temperature as a fall and subsequent rise in accelerated electron population caused by the change in electric drift rate as magnetic fields are swept into and then out of the acceleration region. Our interpretation builds a complete picture of both non–flaring and impulsive reconnection by combining observations of coronal inflows and outflows at X–points with an associated signature of accelerated particles. As such, this work strongly supports the 3D reconnection models outlined in Dalla & Browning \cite{Dalla&Browning,2005} and Browning \textit{et al.} \cite{Browning,et,al.,2010}.

This work could be built upon by taking into account the non–potential nature of the interacting active region during the time coming up to the collapse and flare. In this work, a form of potential field extrapolation was used, while it has been shown on many occasions that linear or non–linear force–free extrapolations are better suited to active region magnetic fields \cite{Scharmer,et,al.,2008,Wiegelmann,et,al.,2005}. With these types of extrapolations, it could be determined with confidence whether the origin of the Type I radio noise storm was indeed consistently outlined by a 3D separator.
6. OBSERVING MAGNETIC RECONNECTION IN A COLLAPSING CORONAL NULL-POINT
The first two goals of the research outlined here were to determine whether the HXR motion and size observed by RHESSI could be explained by the collisional thick–target model. In the first case, the downward–motion of sources was consistent with electron spectral hardening, while in the latter case the implementation of a locally ionised region in the chromosphere was found to be sufficient to explain large HXR source sizes. Finally, an observation was made of the potential source of the nonthermal particles responsible for HXR emission – a magnetic reconnection event high in the corona. This event was analysed and found to potentially be the origin of Type I radio noise storms. In this Chapter we review these results, and indicate how they can be built upon in future work.
7. DISCUSSION AND FUTURE WORK

7.1 Coronal Hard X-ray Source Response to Electron Spectral Hardening

In this work we demonstrated that the location of HXR source emission descends down the legs of the flare loop during the period of spectral hardening (see Chapter 4). The primary result of this work was that this behaviour, based on a prediction of the collisional thick target model, was verified in a unique class of solar flare – an early impulsive event. While this result supports the standard flare model of accelerated electrons in the corona propagating downwards to chromospheric footpoints, it does not exclude alternative models entirely. However, it does provide a strong constraint; any model which invokes global re-acceleration along the loop, or local re-acceleration within the footpoints, would need to also produce descending sources as observed in this event.

7.1.1 Expanded Dataset

Given that this study included only one event, it is of course essential to show that this behaviour is generally apparent in early impulsive flares. This will help to reinforce the interpretation that this descent should occur in all flares with standard loop densities, but is not seen due to contamination by thermal X-rays. In order to expand the analysis to more flares, the first step is to determine a subset of RHESSI–observed events which can be considered early impulsive.

The first step in searching for early impulsive in the RHESSI database is to define a measurable parameter that should work to isolate them. One such example is shown in Figure 7.1 which includes a histogram of ‘delay times’ between
7.1 Coronal Hard X-ray Source Response to Electron Spectral Hardening

**Figure 7.1:** Histogram of SXR delay times for 70,000 RHESSI flares. \( t_{\text{sxr}} \) is defined as the time of the peak of 3–6 keV emission, while \( t_{\text{hxr}} \) is defined as the time of the peak of the highest emission detected, as recorded in NASA’s RHESSI flare list (NASA, 2014). The vertical dash–dotted line shows the delay time for the 28 November 2002 flare, which was studied in detail in Chapter 4.

The peak in HXRs and the peak in SXRs. For this purpose, HXRs are defined as the emission in the highest observed energy band, as recorded in NASA’s RHESSI flare list (NASA, 2014). As shown, a large number of the 70,000 events included in the search exhibit a delay time of 0–10 seconds, which are mostly made up of short thermal events with no significant HXR peak, or where the HXR peak is followed by a thermal peak in the same channel. A small local peak around the time of the 28 November 2002 event indicates that perhaps events which exhibit a delay time of \( \sim 40 \) seconds have a correlation with early impulsive events. This alone is not enough to define the full class of events, so we now describe a more involved approach which makes use of RHESSI’s spectral capabilities.
7. DISCUSSION AND FUTURE WORK

In order to more accurately determine the early impulsive nature of a large number of events, a procedural spectral fitting method was implemented. Selections were made from a large number of RHESSI–observed flares, based on the criteria that the highest energy value recorded was above a certain threshold (to avoid small, primarily thermal events), and the radial distance of the flare from the solar center was above 850 arcsec (so that accurate source heights could be measured in the later stage of analysis). The earliest time interval of this subset of events was chosen such that the total counts summed over the interval were above a certain threshold (decided to be 30,000 counts/detector). This was done by increasing the length of the interval until the threshold was reached, and ensured the spectrum had sufficient counts for reliable spectral fits.

The HXR spectra produced by integrating over these time intervals were then automatically fit by a full variable thermal and a thick–target (broken power–law) component. The important parameter extracted from these fits was the photon energy at which the two components were equal, i.e. the energy below which was dominated by thermal emission, and above which was dominated by nonthermal emission. As the fits were automated, it was presumed that these values were rough estimates only. The results of this analysis are shown in Figure 7.2. Shown is a scatter plot of events against this ‘switch’ energy on the vertical axis, and peak count rate on the horizontal axis. There is a clear separation between a distribution of events which exhibit a low switch energy, below ~20 keV, including the 28 November 2002 event which is shown in red, perhaps indicating a population of early impulsives. However, as these events are known to be rare, it is important to determine if this method accurately isolates them by individually analysing the events.
7.1 Coronal Hard X-ray Source Response to Electron Spectral Hardening

Figure 7.2: Scatter plot of ‘switch’ energy versus peak flux for 52,000 RHESSI events. The ‘switch energy’ is defined here as the energy at which the fitted thermal and nonthermal components of the HXR spectrum are equal, i.e., where one begins to dominate over the other. In early impulsive events, it would be expected that this transition happens at a low energy, as the thermal component is comparatively weak. The red cross represents the switch energy and peak flux for the 28 November 2002 flare, which was studied in detail in Chapter 4.

Upon detailed inspection, a number of early impulsive events were revealed. An example which demonstrates similar behaviour to the event of 28 November 2002 is shown in Figure 7.3, where the lightcurve at the top shows a localised HXR peak before the rise in SXR. Further, the descent of HXR sources, shown by the points in the bottom portion of the Figure, correlates well with the hardening of the derived electron spectrum, shown as a solid line. Quantitatively, this supports the findings outlined in Chapter 4, but in order to build on this, the remaining early impulsive events need to be examined rigorously. This is intended to be done as future work.
7. DISCUSSION AND FUTURE WORK

Figure 7.3: Initial analysis of an early impulsive event. This event was found using the methods outlined above. **Top:** Lightcurve of RHESSI emissions at 3–12, 12–50, and 50–300 keV, including a nearly isolated peak in the harder bands shortly before 03:04 UT, highlighted in gray. **Bottom:** Results of an early study intended for comparison with the HXR source descent and spectral hardening observed in the 28 November 2002 event. The solid line shows the HXR spectral index based on power–law fits to the RHESSI spectrum, while the points indicate the radial height of the source centroids above the solar photosphere.
7.1 Coronal Hard X-ray Source Response to Electron Spectral Hardening

7.1.2 Fokker-Planck Modelling

The CTTM was the model used to produce the predictions of HXR source size in this analysis. It relies on the assumption that energy loss experienced by the accelerated distribution of electrons travelling along the flare loop are a result of Coulomb collisions only, and does not take into account any other mechanisms such as magnetic mirroring or pitch–angle scattering. In order to account for these effects, a common approach is to produce solutions of the Fokker–Planck equation, a method developed by [Hamilton et al. (1990)] which has been used recently in modelling RHESSI emission in solar flares (e.g., Battaglia et al., 2012).

The Fokker–Planck equation describes the evolution of an initial electron flux distribution $F_0(E_0, z)$ in the presence of a magnetic field $B$:

$$\frac{\partial F}{\partial z} - \frac{1 - \mu^2}{2\mu} \frac{d(ln B)}{dz} \frac{\partial F}{\partial \mu} - \frac{Kn(z) \partial F}{\mu E} + \frac{n(z)K}{2E^2\mu} \frac{\partial}{\partial \mu} \left( 1 - \mu^2 \right) \frac{\partial F}{\partial \mu} = -\frac{Kn(z)}{\mu E^2} F$$

(7.1)

where $\mu$ is the pitch angle term equal to the cosine of the angle between the particle trajectory and the magnetic field ($\mu = \cos \theta$). In brief, this equation follows from the Vlasov Equation introduced in Section 2.1.2, adding terms to take into account further individual physical mechanisms such as those mentioned above.

Shown in Figure 7.4 is an early sample output of the simulation which produces a numerical solution of Equation 7.1 to produce an electron distribution function over energy, pitch angle, and position along a flare loop. Here, a height of zero represents the point of acceleration, or the looptop in this model. Two clear local peaks in electron flux exist a few Mm away from the point of acceleration.
Figure 7.4: Sample Fokker–Planck result. Top: Image of nonthermal electron flux against height on the vertical axis and electron energy on the horizontal axis. Here, a height of zero represents the looptop, or point of acceleration, so positive and negative heights represent distance from this point. A sample ‘slice’ through energy near the looptop gives a spectrum (bottom–left), while a slice through height at 100 keV produces a distribution with height (bottom–right).
one first needs to model the bremsstrahlung emission using Equation 2.54.

The first step using this model is to determine the conditions, such as the initial pitch–angle, magnetic field strength and density distributions, under which the observations outlined in Chapter 4 can be reproduced through the same variation in spectral index. It can then be determined whether these conditions are physically reasonable based on previous observations. Following this, it would be useful then to explore the parameter space to determine if other effects, such as a temporal variation from primarily perpendicular to parallel pitch–angle distribution, could produce the observed descent. If so, an alternative and more complete interpretation of HXR source descent could be presented. Alternatively the argument for the close relationship between HXR descent and spectral hardening would be strengthened.

7.2 Hard X-ray Source Sizes in a Beam-Heated and Ionised Chromosphere

In this work, as outlined in Chapter 5, model HXR source sizes were produced for the case where a local peak in ionisation and temperature existed within the chromosphere, as expected to be caused by the accelerated electrons themselves. Theses sizes were shown to be 3–6 times larger than those for the standard CTTM in a neutral chromosphere, for commonly–observed energy ranges (∼30–70 keV). This has largely accounted for the difference between CTTM–based model HXR source sizes, and those observed by high–resolution RHESSI imaging, first highlighted in Kontar et al. (2010). While this work does not completely remove the
7. DISCUSSION AND FUTURE WORK

requirement for other energy loss mechanisms or density redistributions to take
place in flares, it vastly reduces the requirement for them to occur ubiquitously.

As with the work outlined in Chapter 4, the modelling performed here to
address HXR source sizes will benefit from the supplementation of Fokker–Planck
simulations, as they would provide a more complete description of the response of
electron beams in the chromosphere to a local peak in temperature and ionisation.

7.2.1 HXR Source Asymmetry

While the modelling work outlined in Chapter 5 was presented primarily as a
solution to the ‘HXR size problem’, there is an additional interesting property
of the spatial structure of HXR emission which may be closely related to non–
uniform ionisation (NUI). The simple implementation of a locally ionised region
in the chromosphere results in a dual–peaked or at least asymmetric gaussian
in HXR intensity along the 1D loop. This is demonstrated in Figure 7.5. The
density and ionisation fraction for the target atmosphere are shown at the top for
the neutral case (in black) and the the local NUI case (in red). Similarly, the HXR
emission profile along a vertical distance through the chromosphere is shown for
increasing energy below. Notably, for increasing emitted photon energy, the HXR
structure gradually transitions from this asymmetric source to a single symmetric
Gaussian, as emission is produced in the region below the peak in ionisation.

This general behaviour of the transition from a dual–peaked or asymmetric
HXR source at low energies to a simple single Gaussian source at higher energies
could be observable in RHESSI observations of nonthermal emission at the limb.
In particular, it could be argued that this behaviour would be more visible at
7.2 Hard X-ray Source Sizes in a Beam-Heated and Ionised Chromosphere

Figure 7.5: HXR source asymmetry in a locally ionised chromosphere. (a) Density and ion fraction for a simple model solar atmosphere for the case of a neutral chromosphere (black) and a chromosphere with a local peak in ionisation fraction (red). b-d) Simulated HXR emission from the chromosphere at 20, 40 and 81 keV, calculated based on the collisional thick target model, again for the fully neutral (black) and locally ionised (red) cases presented above.
lower energies, where the emission is expected to be located near the peak in ionisation, and therefore early impulsives are again an ideal subset of flares to search. A number of events were extracted from the list produced in Section 7.1.1 and images were reconstructed with the goal of identifying this behaviour.

A number of these flares were seen to exhibit this behaviour, with an early observation of one case shown in Figure 7.6. The HXR profiles shown on the right indicate the emission observed along the slices of the images shown on the left, where the dashed line shows emission along the central arc, and the solid line shows emission integrated across the distance between the two outer arcs. As zero in this case is the position at the top of the arcs, the right-hand side of the profiles can be seen as nearer the photosphere. As shown, with increasing energy, emission evolves from a dual-peaked structure to a single Gaussian. Images at even higher energies show a completely localised Gaussian.

An important caveat here is that observations of a transition from an extended, potentially asymmetric HXR source to a localised Gaussian can also occur as a result of transition from thermal emission in the loop to nonthermal emission at the footpoint. For this reason it is crucial to ensure that all images used in this type of analysis are dominated by nonthermal effects.

Future work involves examining a larger number of flares, as well as by determining the feasibility of detecting HXR source asymmetry by direct analysis of the RHESSI visibilities rather than in reconstructed images. If this can be done reliably, an automated routine to detect this feature can then be developed.
7.2 Hard X-ray Source Sizes in a Beam-Heated and Ionised Chromosphere

Figure 7.6: Observations of HXR asymmetry in a large solar flare. Left: RHESSI HXR PIXON images, overlaid with three curved slices. Right: HXR profiles of emission along the central slice (dashed line) and integrated across the distance between the two outer slices (solid line). In this case, zero is the position at the top of the slice in the images on the left. Energy of photon emission increases from top to bottom.

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7.3 Observing Magnetic Reconnection in a Collapsing Coronal Null–Point

The work outlined in Chapter 6 serves as a first observation of the close association between a Type I radio noise storm and a coronal X–point observed in EUV emission. The clear connection between the high–frequency noise storm emission and the location of the EUV X–point was interpreted as evidence of a brand of acceleration and reconnection that has, in recent years, become a topic of great theoretical study, namely 3D separator reconnection. Further work is intended to be done in order to quantify the relationship between these two types of emission, as well as gain a greater understanding of the detailed radio spectra.

7.3.1 AIA Image Processing

The time–distance plots presented in Figures 6.4 and 6.5 of Chapter 6 were produced using the unprocessed AIA 171 and 193 Å images available at the time. In order to enhance any observation of inflow and outflow, it is intended to repeat this analysis on the same images after they have been processed to highlight the fine structure of solar loops. An example of a piece of software designed to do this is the multi–scale Gaussian normalisation (MGN) routine presented by Morgan & Druckmüller (2014).

The MGN algorithm produces an output image $I_{out}$ from an input image $I_{in}$ by the following relation:

$$I_{out} = hC_g' + \frac{1 - h}{n} \sum_{i=1}^{n} g_i C_i'. \quad (7.2)$$
7.3 Observing Magnetic Reconnection in a Collapsing Coronal Null–Point

Here, $C'_i$ is the arctangent of $kC$, where $k$ is a variable scaling factor, and $h$ is the scaling factor of the sum of $C'_g$ and $C'_{ni}$. $C'_{ni}$ is a renormalisation function which takes an image $I_{in}$ and produces a new image that is locally normalised to a standard deviation of one and a mean of zero, by:

$$ C = \frac{I_{in} - I_{in} \ast k_w}{\sigma_w}. \quad (7.3) $$

Here, $k_w$ is a 2D Gaussian kernel of width $w$ in $x$ and $y$, and $\sigma_w$ is given by

$$ \sigma_w = \sqrt{((I_{in} - I_{in} \ast k_w)^2) \ast k_w}. \quad (7.4) $$

This normalisation is performed to produce $n$ locally normalised images, each for a different value of $w_i$, the one–sigma width of the Gaussian kernel. The normalised sum of these images is then added to an image produced using the standard method of a gamma–transform (taking the original image to the power of $1/\gamma$, where $\gamma > 1$), to arrive at Equation 7.2. For further details on the motivation for this analysis, see Morgan & Druckmüller (2014).

A demonstration of the MGN processing is shown in Figure 7.7. The solar image on the left has been gamma–transformed, while the image on the right was created from the same base image using the MGN code. As shown, the fine details of the off–limb loops, as well of the on–disk active region, are revealed in the image on the right. It is intended to perform this same analysis on the AIA images of the X–shaped structure presented in Chapter 6. Following this, analysis of the inflows and outflows will be repeated, perhaps resulting in a clearer observation of reconnection flows.
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Figure 7.7: A demonstration of the multi-scale Gaussian normalisation algorithm. 
Left: An AIA 171 Å image of the western solar limb during a flare and filament eruption. This image was produced by the gamma-transform (taking the original image to some power between 0 and 1). Right: The same image, produced using the MGN code, clearly highlighting the fine details of the coronal loop structure, especially off-disk \citep{MorganDruckmuller2014}.

7.3.2 Detailed radio analysis

The X-point collapse described in Chapter 6 was of particular interest because it occurred during an observing campaign of the LOw-Frequency ARray (LOFAR; \citep{vanHaarlem2013}). As shown in Figure 7.8, high-resolution radio dynamic spectra from 110 to 190 MHz were recorded for the time coming up to 10:00 UT, when the collapse occurred. In fact, observations are available for 8:00 UT to
7.3 Observing Magnetic Reconnection in a Collapsing Coronal Null–Point

**Figure 7.8:** LOFAR dynamic spectrum of radio emission during the X–point collapse outlined in Chapter 6. As shown, the Type I noise storm which persists throughout the day is well–observed by LOFAR. In particular, a drifting bursty structure is apparent in the 20 minutes prior to the flares and collapse of the active region loops studied previously. The emission appears to drift to lower frequencies, which for plasma emission can be interpreted as a drift to lower densities.

16:00 UT on the same day. As shown, the radio noise storm was well–observed by LOFAR, which was sensitive enough to detect emission for the full day of observations. The drifting bursty structure in the latter half of the shown dynamic spectrum might indicate drift or motion of the acceleration region to lower densities – perhaps higher in the corona – during the collapse phase of the coronal X–point. Detailed analysis of these structures and how they correlate temporally to the periods of inflow and outflow described in Chapter 6 is intended as future work.

In addition to spectral analysis, the observations at this time were made in the tied–array beam mode, meaning that imaging of the emission at any time
7. DISCUSSION AND FUTURE WORK

and frequency is possible (see Morosan et al. (2014) for details). If the images produced at the frequencies shown in Figure 7.8 are not noise–dominated, it may be possible to localise the noise storm at frequencies below the 150 MHz lower limit of NRH. This could potentially reveal the nature of the emitting structure to greater heights or lower densities in the corona.

7.4 High–Resolution Observations of Ribbon Formation

A final series of observations which were completed shortly before the completion of the work outlined here were done using the CRisp Imaging Spectro–Polarimeter (CRISP) camera at the Swedish Solar Tower (SST; Scharmer et al., 2003) on the island of La Palma, Spain. The SST, as shown schematically in Figure 7.9, includes a 1 m primary lens which passes light through a vacuum tube to an optical bench in an observing station below ground. The light can then be passed to various cameras, including CRISP.

The CRISP apparatus includes a dual FabryPerot interferometer, which provides the capability of fast wavelength tuning within the 5100 to 8600 Å range of the visible to near-IR portion of the spectrum (Ortiz & van der Voort, 2010; Scharmer et al., 2008). CRISP is therefore capable of high–resolution Hα imaging spectro–polarimetry and can rapidly scan through numerous available positions along the 6563 Å line. As with any ground–based instrument, CRISP images are susceptible to the distortion caused by atmospheric turbulence. These effects, referred to as ‘seeing’, are largely removed during the time of observation.
Figure 7.9: Schematic diagram of the Swedish Solar Tower. The $\sim 1$ m primary fused silica lens is located at the top of the tower in order to avoid the refraction caused by convective air flows near the ground. The lens is mounted on a turret which tracks the Sun throughout the day. The light then reflects off the two shown 1.4 m flat mirrors to be passed to the mirror setup labelled as $c$, which passes the light to an optical bench which includes the desired beam splitters and cameras. $a$ and $b$ show the Schupmann corrector system, which is intended to cancel out the effects of the 1 m lens. (Scharmer et al. 2003).
by the adaptive optics (AO) system in the SST. Furthermore, images exhibiting a reduction in resolution due to seeing can be processed by the Multi-Object Multi-Frame Blind Deconvolution reconstruction technique \cite{vanNoort2005MOMFBD}.

During the observing campaign of June 2014, a rare observation was made of ribbon formation in an X-class flare during its very early onset, during which RHESSI was also making observations. Two sample MOMFBD images produced during this time is shown in Figure 7.10. From left to right, the H\(\alpha\) line core (6563 Å), H\(\alpha\) blue wing (6563 - 1.29 Å), and the AIA 1700 Å images are presented. The CRISP images show a clear pair of ribbons which separate with time, but images were in fact produced from their formation at \(\sim 12:40:00\) UT. An interesting feature is the evolution of the fine spiked structures along the ribbon, which is a topic of future work.

In this observation, despite the potential seeing effects, every image in the X-class flare observation is close to the theoretical diffraction limit for the SST (130 km). Data processing was implemented in order to deal with higher frequency distortions (beyond the detectable range of the AO) with the use of standard procedures in the reduction pipeline for CRISP data \cite{delacruzrodriguez2014} and this includes a post-MOMFBD correction for differential stretching \cite{Henriques2012} resulting from isoplanatic patches within the Earth’s atmosphere.

The co–temporal RHESSI observations provide a unique opportunity to investigate the nature and timing of the relationship between H\(\alpha\) and HXR emission in the very early stages of a large solar flare. A number of previous studies have investigated the time delay between H\(\alpha\) and HXR or microwave flare emissions (e.g.,}
Figure 7.10: CRISP images of ribbon formation in an X-class flare. Shown are the Hα blue wing (top–left), and line core (bottom–left) images produced by CRISP at the SST on 10 June 2014. For comparison with UV emission, the 1700 Å AIA image is also shown (right). This also provides a reference for the spatial resolution of the CRISP images.
Studies such as these can be done in order to narrow down the nature of the energy transport mechanism from corona to chromosphere. For example, Radziszewski et al. (2011) recently concluded that some events show evidence for energy transfer by nonthermal beams, while others are more consistent with transfer by thermal conduction fronts. The event described here is unique in that it is an X–class flare observed from the very early onset, such that detailed RHESSI imaging and spectroscopy can be performed during ribbon formation with good counting statistics.

Preliminary RHESSI analysis of the event has been performed, as summarised in Figure 7.11. As shown, RHESSI HXR sources appear primarily from two locations at the north and south end of the developing flare ribbons. During this time, the HXR spectrum exhibits two to three periods of softening and hardening. The imaging can be improved by the use of the VIS FWDFIT method, which will produce high–resolution centroid locations, which can be compared with the Hα ribbon jet–like features and brightenings to investigate any correlation. In addition, the de–modulated high–resolution RHESSI lightcurves can be compared with the integrated intensity of Hα emission at various points along the ribbons as a second method of gaining spatial information on the origin of the HXR emission. These tasks are intended to be completed in future work.
7.4 High-Resolution Observations of Ribbon Formation

Figure 7.11: Summary of RHESSI source correlation with CRISP ribbon formation. *Top:* CRISP Hα -1.29 Å images, revealing formation of flare ribbons in an the (second) X-class flare of 10 June 2014. Overlaid as solid black lines are the RHESSI 50% and 90% contours of CLEAN images produced at the same time. *Bottom:* RHESSI spectra produced at the same time as the associated images above. Crosses represent the background-subtracted RHESSI X-ray spectra, while the red and green lines are the thermal and thick-target power-law fits, respectively.
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