

RICE UNIVERSITY

**Efficient Two-fluid Hydrodynamic Simulations of
Impulsively Heated Coronal Loops**

by

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ABSTRACT

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The coronal heating problem has puzzled physicists and astronomers for nearly a century. While the origin of these surprisingly high temperatures in the upper solar atmosphere is most certainly magnetic,

Contents

Abstract	ii
List of Illustrations	v
List of Tables	vii
1 Introduction	1
1.1 Structure of the Solar Atmosphere	3
1.2 The Solar Magnetic Field	6
1.2.1 The Solar Dynamo and Flux Emergence	6
1.2.2 Magnetic Reconnection	9
1.3 The Coronal Heating Problem	12
1.3.1 AC Heating	14
1.3.2 DC Heating	15
1.4 Summary	18
2 Coronal Loops	20
2.1 Observations	22
2.2 Modeling	24
2.2.1 Static versus Dynamic	24
2.2.2 Magnetohydrodynamics	24
2.2.3 Hydrodynamics	24
2.3 Summary	24
3 Emission Diagnostics	25
3.1 Spectral Line Intensity	26

3.2	Differential Emission Measure	29
3.2.1	Reconstruction Techniques	30
3.2.2	Scaling Laws	32
3.3	Forward Modeling	33
4	Numerical Modeling	36
4.1	One-dimensional Hydrodynamics	36
4.2	The EBTEL Model	38
4.3	The Two-fluid EBTEL Model	38
5	Results	40
6	Conclusions	41

Illustrations

1.1 Total eclipse as seen from Svalbard, Norway on March 20, 2015. Open and closed loops in the highly-structured solar corona are clearly visible. Photo courtesy of Miloslav Druckmüller.

1.2 (a) Schematic showing the layers of the solar atmosphere; the corona is the uppermost layer of the atmosphere. Courtesy of NASA (b) Temperature and density as a function of height above the solar surface; the corona is characterized by low densities and anomalously-high temperatures. Taken from Gary et al. (2007) . . .

1.3 (a) Flux emergence from twisted and tangled field below the photosphere yields loop structures that stretch high into the solar atmosphere. Taken from Gold (1964). (b) The coronal magnetic field topology is extremely complex, making determinations of coronal plasma dynamics difficult. Shown here is a magnetic field extrapolation constructed from an optical magnetogram. An image from the TRACE satellite is superimposed to show the loop structures. Taken from Reale (2010).

1.4 Evolution of β as a function of height in the solar atmosphere. In the chromosphere, the gas pressure dominates ($\beta > 1$) while in the corona, the opposite is true and $\beta < 1$, making cross-field motion negligible. Out in the solar wind, the plasma returns to $\beta > 1$. Taken from Gary (2001).

1.5	Cartoon showing basics of reconnection process. Image taken from Priest & Forbes (2000)	9
1.6	(a) Cartoon showing braided flux tubes rooted, or line-tied, to the solar surface. Braiding allows for the build-up and relaxation of the field, providing an energy release mechanism capable of heating the solar corona. Taken from Parker (1983b). (b) Image from the <i>Hi-C</i> sounding rocket mission showing spatially-resolved braided flux tubes in the solar corona. Image taken from Cirtain et al. (2013)	16
2.1	Image captured by the 171 Å channel of the Atmospheric Imaging Assembly (AIA) onboard the Solar Dynamics Observatory (SDO). An arcade of loops in an active region can be seen on-disk in the lower left corner. Several prominent loop structures also appear off-limb on the right side. Image courtesy of NASA/SDO.	21
3.1	(a) Simulated DEM curve and the resulting EM loci reconstruction. From the DEM curve, the plasma is approximately isothermal, being sharply peaked about 1 MK. Correspondingly, the EM loci curves intersect in a region centered on 1 MK, allowing for a well-constrained determination of T_C . Image taken from Landi & Klimchuk (2010). (b) Emission measure distribution reconstruction for an active region core. The colored curves show the emission measure calculation using the EM loci method. The black and red histograms show the $\text{DEM}(T)dT$ reconstructed using the MCMC method of Kashyap & Drake (1998). Here, the DEM is multiplied by the temperature bin for comparison with the EM loci calculation. Image taken from Warren et al. (2011)	35

Tables

2.1 Typical parameters for several different types of loops. Adapted from Reale (2010).	22
3.1 Summary of emission measure scalings from observational and modeling studies. Scalings are typically calculated coolward of the emission measure peak and typically range from 2 to 5. Adapted from Bradshaw et al. (2012)	34

Chapter 1

Introduction

The Sun is the most important celestial body to life on Earth. For the last five billion years, it has provided the light by which humans observe the world around them and the heat to save the planet from the frigid temperatures of interplanetary space. Because of its proximity, the Sun provides astronomers an exclusive and unique look into how stars behave. By observing and understanding the Sun, one can make conclusions about other types of stars in our galaxy and the universe.

Perhaps no other consistent celestial event has attracted as much attention as the solar eclipse. Solar eclipses have been observed and recorded for thousands of years, with some reports dating back to the fourteenth century BC (Golub & Pasachoff, 2010). The recordings of ancient eclipses have been heavily studied. Chinese rock drawings dating back to the Han dynasty (approximately 1900 years ago) appear to show the moon completely obscuring the Sun. Additionally, some have even suggested the Aubrey holes that surround the Stonehenge site were used to track and predict both solar and lunar eclipses (Golub & Pasachoff, 2010). Though some claims of ancient eclipse studies are controversial, it suffices to say that humans have long sought to study and explain the behavior of the nearest star to Earth.

In particular, solar eclipses have captured the attention of artists for centuries.



Figure 1.1 : Total eclipse as seen from Svalbard, Norway on March 20, 2015. Open and closed loops in the highly-structured solar corona are clearly visible. Photo courtesy of Miloslav Druckmüller.

Cosmas Damian Asam, a Bavarian painter and architect active in the early eighteenth century, used images of solar eclipses in many of his works, including several frescoes and an altarpiece. Olson & Pasachoff (2007) discuss how Asam, a deeply religious artist, was commissioned several times to depict the vision of St. Gregory the Great, a Benedictine monk, as described in his work *Dialogues*.

One of these depictions, an altarpiece at a Benedictine monastery in Kladruby, Czech Republic (see Fig. 7 of Olson & Pasachoff (2007)), shows “the visionary globe surrounded by a glowing halo of yellow light that more closely resembles the solar corona” (Olson & Pasachoff, 2007). An additional altarpiece at another monastery in Weltenburg, Germany shows a perhaps even more pronounced depiction of the solar corona during a total solar eclipse. Olson & Pasachoff (2007) note that, given his detailed depictions, Asam must have observed several solar eclipses as well as the

solar corona, with these astronomical events profoundly impacting his depictions of supernatural events in his works. This is but one example of how solar eclipses and their consequential insight into the highly structured solar atmosphere, have shaped scientific and artistic discourse throughout history. maybe here need some info on observing instruments, SDO/AIA, Hinode, Yohkoh, TRACE, IRIS, Hi-C; talk of early observations probably easily transitions into discussion of more modern observations, understanding of the Sun/corona need some sort of wrap-up for this section; maybe also add something here about practical applications/space weather?

1.1 Structure of the Solar Atmosphere

Though the Sun can be easily seen from Earth, its dynamic and highly structured atmosphere is not observable with the naked eye, with the one exception being brief glimpses of the corona during an eclipse. The interior of the Sun is of course very complex and constitutes a very different regime of physics than that seen in the solar atmosphere. Thus, this work will be primarily limited to the upper solar atmosphere with some discussion of the lower layers.

The solar atmosphere is often divided up into four separate regions: the photosphere, the chromosphere, the transition region, and the corona. Fig. 1.2(a) shows a cartoon of the different layers while Fig. 1.2(b) shows the density and temperature profiles of the atmosphere with each region labeled. The *photosphere* is what we typically refer to as the solar surface, with the actual surface located where the

optical depth, τ , is equal to 1. This region also contains the lowest temperature on the Sun, approximately 4400 K, located about 525 km above the surface (Carroll & Ostlie, 2007). The photosphere is where the majority of the visibly (with the naked eye) observable photons originate.

The temperature minimum defines the top of the photosphere above which lies the *chromosphere*. As can be seen from Fig. 1.2(b), the density in the chromosphere is many orders of magnitude less than that of the photosphere and the temperature has increased from the minimum up to about 1×10^4 K. Though not visible with the naked eye, the chromosphere is highly structured. Structures such as spicules, tall columns of gas that extend high into the solar atmosphere thought to heavily impact the behavior of plasma in the corona (De Pontieu et al., 2011), originate in the photosphere as well as filaments, essentially spicules observed on-disk and plage, bright regions surrounding sunspots.

Next is the *transition region*, so called because of the steep temperature and density gradients (see Fig. 1.2(b)) that mark the transition between the chromosphere and the corona. The transition region is extremely thin, only a few hundred kilometers as compared to the chromosphere which extends over many thousands of kilometers. However, in this very short change in altitude, the temperature in the solar atmosphere jumps from $\approx 1 \times 10^4$ K to temperatures exceeding 1×10^5 K.

Finally, the solar *corona*, or “crown”, is the highly-dynamic uppermost layer of the Sun’s atmosphere. Visible with the naked eye only during a total solar eclipse (see Fig.

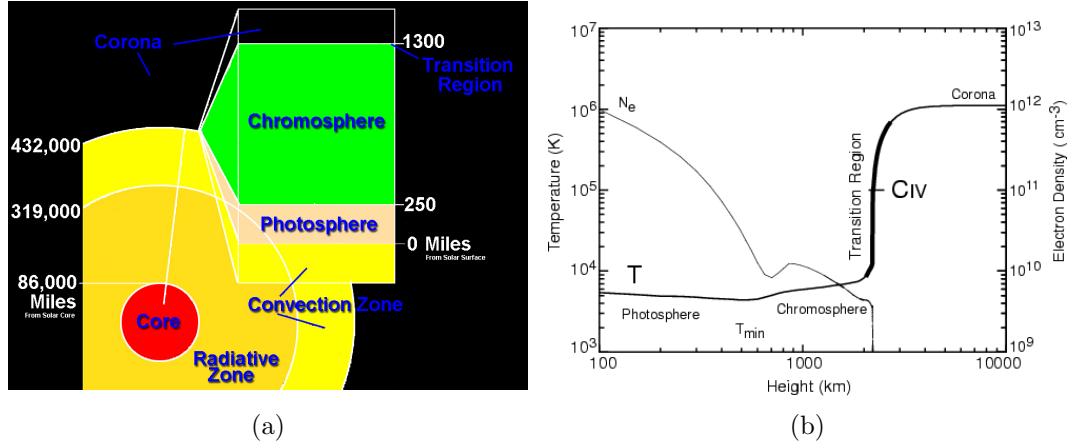


Figure 1.2 : (a) Schematic showing the layers of the solar atmosphere; the corona is the uppermost layer of the atmosphere. Courtesy of NASA (b) Temperature and density as a function of height above the solar surface; the corona is characterized by low densities and anomalously-high temperatures. Taken from Gary et al. (2007)

1.1), the corona is highly-structured and diffuse. Here, the temperature continues to increase, with typical coronal temperatures exceeding 1×10^6 K. Particularly high temperatures ($\approx 1 \times 10^7$ K) have also been observed in *active regions*, sites of intense magnetic activity associated with sunspots. These active regions can contain plasma as cool as 1×10^4 K as well and represent some of the most dynamic portions of the solar corona.

The work presented here will focus primarily on the plasma dynamics in the corona, in particular, in the cores of active regions where the intense magnetic field drives the motion of the plasma. In the following sections, we will discuss the origin of the solar magnetic field, its topology, how it impacts plasma in the solar corona, and finally how it connects to the anomalously high temperatures seen in the upper solar atmosphere.

1.2 The Solar Magnetic Field

Because of the high temperatures that characterize both the solar atmosphere and interior, much of the gas that makes up the Sun is ionized; that is, each atom has been stripped of at least one of its electrons. This means that the Sun is filled by a sea of charged particles, or plasma. In the corona where temperatures are often higher than 1×10^6 K, even heavy elements such as Mg, Ca, and Fe are stripped of their electrons. Because these particles are charged, their motion is strongly dictated by both the electric and magnetic fields via the Lorentz force law. Thus, to understand the dynamics of the coronal plasma, one must first try to understand the magnetic field that ultimately controls it.

1.2.1 The Solar Dynamo and Flux Emergence

The Sun, like the Earth, possesses an intrinsic magnetic field, though its origin and dynamics constitute a largely unsolved problem. Additionally, unlike Earth, the solar magnetic field is highly dynamic and comparatively quite strong (up to 1000 G in sunspots as compared to the < 1 G field on Earth) (Aschwanden, 2006). From the earliest studies of magnetism, it has been known that a conductor moving through a magnetic field produces a current. This current then in turn produces an additional magnetic field. Treating the convection zone (below the photosphere) plasma as the conductor and assuming some preexisting, but small magnetic field leads to what is commonly known as *dynamo theory* (Golub & Pasachoff, 2010).

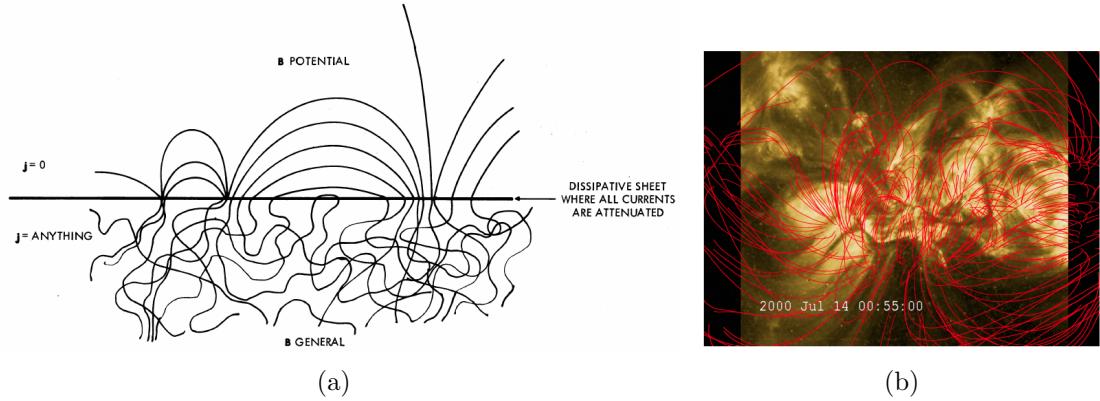


Figure 1.3 : (a) Flux emergence from twisted and tangled field below the photosphere yields loop structures that stretch high into the solar atmosphere. Taken from Gold (1964). (b) The coronal magnetic field topology is extremely complex, making determinations of coronal plasma dynamics difficult. Shown here is a magnetic field extrapolation constructed from an optical magnetogram. An image from the TRACE satellite is superimposed to show the loop structures. Taken from Reale (2010).

Dynamo theory seeks to show self-consistently how the interaction between the Sun’s hot plasma and some small initial field leads to the observed complicated topologies and field strengths. Modeling such a system is no easy task as the interactions between the field and the plasma are highly non-linear, meaning solving the so-called “dynamo equations” requires a significant amount of computational effort. But how does this field affect the plasma dynamics of the corona? The solar magnetic field is said to be *frozen in* to the solar plasma; that is, the field moves with the plasma. Since the Sun is not solid, it undergoes *differential rotation*, meaning that different latitudes are spinning at different rates. Since the field follows the motion of the plasma, this causes the normally dipolar (north-south aligned) field to have an east-west component. This amplified and twisted field is then carried up to the photosphere by magnetic buoyancy, an idea first proposed by (Parker, 1955), resulting

in dipolar loop-like structures poking through the solar atmosphere (see Fig. 1.3(a)).

This phenomenon, commonly referred to as magnetic flux emergence, is by no means a solved problem and is a continuing topic of research (see Cheung & Isobe, 2014).

Finding out how the solar magnetic field is generated and how it makes its way to the upper atmosphere will undoubtedly help to explain much of the observed plasma dynamics in the corona.

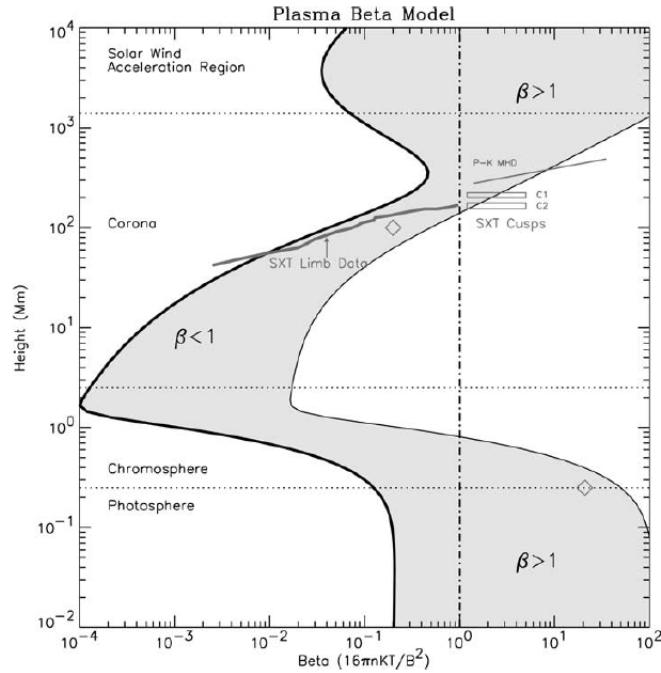


Figure 1.4 : Evolution of β as a function of height in the solar atmosphere. In the chromosphere, the gas pressure dominates ($\beta > 1$) while in the corona, the opposite is true and $\beta < 1$, making cross-field motion negligible. Out in the solar wind, the plasma returns to $\beta > 1$. Taken from Gary (2001).

1.2.2 Magnetic Reconnection

After the magnetic field is forced into the solar atmosphere by the buoyant motion of the convective zone, it remains rooted in the photosphere, whether it is an *open* (flux tube extends radially outward, possibly connecting with the interplanetary magnetic field) or *closed* (both ends attached to the solar surface) field line. Because the field is *line-tied*, or frozen into the photospheric plasma, the turbulent motion of the photosphere deforms and stresses the overlying field, leading to the storage of magnetic energy. This motion can eventually lead to a topological restructuring of the magnetic field as it relaxes from a stressed to an equilibrium state, a process popularly known as *magnetic reconnection*.

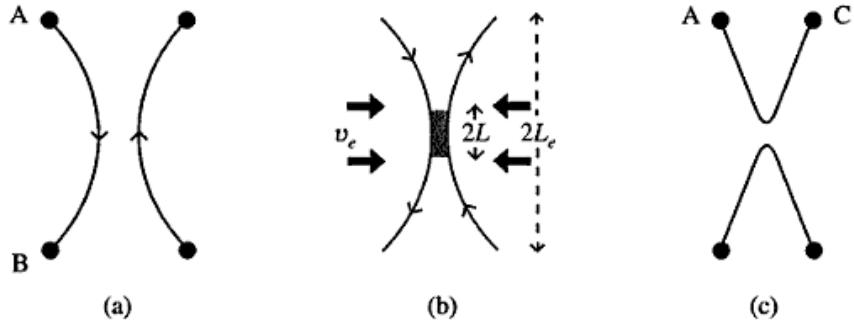


Figure 1.5 : Cartoon showing basics of reconnection process. Image taken from Priest & Forbes (2000).

Magnetic reconnection is thought to be a dominant process in a variety of space and astrophysical plasma environments, including Earth's magnetosphere, solar flares, and accretion disks. Additionally, reconnection has also been observed in laboratory settings such as the tokamak and the reversed field pinch (Priest & Forbes, 2000).

It is also thought to be the primary driver of some coronal heating mechanisms (see §1.3.2). The basic idea behind reconnection is illustrated in Fig. 1.5. When two oppositely-directed field lines are brought together in a conducting fluid (such as a plasma), a tangential discontinuity develops between them with current-carrying plasma squeezed into this area of discontinuity. Because the field lines are frozen into the plasma, a large magnetic gradient develops at the discontinuity in an area called the *diffusion region* (the gray box in Fig. 1.5). Because of these large gradients, the resistivity in this region becomes very high, allowing the magnetic field lines to diffuse through the plasma, reconnect, and relax to a topologically different, but more energetically favorable state. As the field lines reconnect and are pushed out of the diffusion region, they accelerate and heat the plasma. Reconnection is thus a non-ideal process as it allows for the conversion of stored magnetic energy to kinetic and thermal energy via dissipation (Priest & Forbes, 2000).

Research into magnetic reconnection has been continually driven by the desire to understand fundamental processes in astrophysical plasmas, primarily those that heat the corona and lead to solar flares. The first complete theory of reconnection was proposed by Sweet (1958) and Parker (1957, 1963). In the Sweet-Parker model, as it is popularly known, a diffusion region of length $2L$ and width 2ℓ is defined between two anti-parallel fields. The field lines are carried into the diffusion region with speed $v_i = \eta/\ell$, where η is the magnetic diffusivity. Using Maxwell's equations and the mass and momentum conservation equations from magnetohydrodynamic (MHD) theory,

the inflow velocity, or reconnection rate, can be rewritten as $v_i = v_{Ai}/\sqrt{R_{mi}}$, where v_{Ai} is the Alfvén speed and $R_{mi} = Lv_{Ai}/\eta$ is the Magnetic Reynolds number (Priest & Forbes, 2000). Unfortunately, the Sweet-Parker model predicts a reconnection rate far too slow to properly account for the energy release timescales observed in flare plasmas. In an effort to remedy the slow reconnection in the Sweet-Parker mechanism, Petschek (1964) suggested that magnetoacoustic shocks could provide an additional acceleration mechanism for the reconnection rate. Additionally, he proposed a smaller diffusion region, further shortening the reconnection timescale. For many years following Petschek’s work, it was thought that the problem of fast reconnection was solved. Today, however, thanks in part to increased computing power that makes three-dimensional and kinetic simulations possible, reconnection is now understood to be a far more subtle mechanism than previously thought, with the Petschek and Sweet-Parker models as only special cases (Priest & Forbes, 2000).

It goes without saying that the solar magnetic field is an extremely complex and dynamic system. Fortunately, one can still gain insight into the solar corona without completely solving the problems of the solar dynamo or magnetic reconnection. The corona constitutes what is commonly referred to as a “low-beta plasma” such that $\beta = p_{\text{thermal}}/(B^2/8\pi)$, the ratio between the gas pressure and magnetic pressure, is $\beta < 1$ (see Fig. 1.4). This means that the coronal magnetic field acts as a guide for the plasma, with the flow directed primarily in the field direction with negligible cross-field motion. A useful analogy is that of water flowing through a pipe: the

rigid pipe directs the flow of the water, but has no affect on the underlying fluid dynamics. Provided the field is reasonably stable and the relevant fluid equations are parameterized in terms of the field-aligned coordinate, our equations governing the plasma dynamics need not include the field at all!

Talk about drawbacks, mention MHD and hydrodynamic and hydrostatic; lead into coronal heating conversation; not sure if this is actually needed here...

1.3 The Coronal Heating Problem

Thus, far we have not addressed the important question of why the solar corona is so much hotter than the surface. From an intuitive, thermodynamic perspective, we would expect the atmospheric temperature to decrease as we moved away from the surface. However, nearly 70 years of observations have shown this is not the case and these anomalously high temperatures, dubbed the “coronal heating problem,” continue to baffle solar physicists today.

The discovery of the $> 1 \times 10^6$ K corona was made over the course of nearly fifty years through many laborious spectroscopic observations. Spectroscopic measurements during the late-nineteenth and early twentieth centuries of solar eclipses and images obtained from coronagraphs yielded a surprising result: an unknown line at 5303 Å , dubbed the “coronal green line” (Golub & Pasachoff, 2010). Because the line could not be associated with any known element, the line was identified as a new element, “coronium”. Later, more detailed examinations of the coronal spectrum re-

vealed several more unidentifiable spectral lines. Grotrian (1939) showed that many of these lines were due to highly ionized states of heavy elements such as iron and calcium. Additionally, Edlén (1943) identified four additional coronal lines, Fe X, Fe XI, Ca XII, and Ca XIII, providing more evidence for the presence of highly-ionized atoms in the corona and implying temperatures in excess of 1×10^6 K. Because of their work in coronal spectroscopy, Grotrian (1939) and Edlén (1943) are generally credited with the discovery of the million-degree solar corona. However, some (see Peter & Dwivedi, 2014) have argued that early coronal spectroscopists did not imply a million-degree corona and that it was in fact Nobel laureate Hannes Alfvén, in Alfvén (1941), who first proposed a hot corona, even arguing that it was the interaction between the magnetic field and the charged particles in the solar atmosphere that thwarted our usual thermodynamic intuition.

While these early spectroscopic observations are now better understood and a multitude of satellite observations have allowed a much clearer picture of the solar atmosphere, the exact mechanism responsible for the million-degree corona remains unknown. This is certainly not for lack of effort; over the last several decades, many different heating mechanisms have been proposed. These are traditionally divided into two-separate categories: *AC*, mechanisms that rely on waves to transfer energy from the lower atmosphere into the corona and *DC*, mechanisms that involve dissipation of energy stored in the stressed magnetic field. In order to be successful, a given mechanism must be able to explain the energy source of the heating, how the energy

is converted to heat, how the plasma responds and accurately predict any observables from the resulting plasma emission (see Fig. 1 in Klimchuk, 2006).

1.3.1 AC Heating

The same turbulent sub-photospheric motions that stress the magnetic field can also generate waves capable of propagating into the upper solar atmosphere. The AC- or DC-type heating classification depends on the timescale of the stressing motion: if it is longer than a characteristic crossing time of the coronal structure, it is classified as DC heating. If the timescale is shorter than a crossing time, it is classified as AC heating. A variety of wave modes have been observed in both the open and closed corona, including acoustic waves, Alfvén waves, and slow- and fast-mode MHD waves (Aschwanden, 2006). However, the simple existence of these waves is not enough to make them a viable candidate for coronal heating; they must be able to propagate into the corona with an adequate amount of energy and then efficiently dissipate this energy in order to heat the coronal plasma. For example, acoustic waves are capable of carrying an adequate amount of energy to heat the corona, but are almost entirely reflected by the steep density gradients in the transition region. Alfvén waves may also be capable of carrying enough energy to heat the corona, but specific frequencies are required to avoid reflection at the transition region. These particular modes have also been found to be non-dissipative under coronal conditions, making it hard for these waves to actually heat the corona even if they make it to the upper atmosphere

(Klimchuk, 2006). Wave modes generated in the corona would of course overcome the problem of crossing the transition region boundary. However, while oscillations in the corona have been observed (De Moortel et al., 2002b,a), the properties of these waves have not been measured precisely enough to say whether or not they are capable of heating the corona (Klimchuk, 2006).

1.3.2 DC Heating

As discussed in §1.2, the magnetic field that extends into the corona is rooted in the photosphere. Because of the underlying convection zone, the footpoints of these magnetic loops are subject to a turbulent velocity field and thus undergo a random walk across the surface. The turbulent motion of the footpoints causes the overlying loops to become twisted and braided around each other, leading to a highly stressed magnetic field above the solar surface. Gold (1964) first proposed the idea of a twisted field created by photospheric footpoint motions, suggesting that it was the relaxation of the stressed field to its equilibrium (or force-free) state that provided the energy release mechanism needed to power flares. Parker (1983a,b) applied this same concept to the corona and showed analytically that closely-packed magnetic flux tubes are subject to “dynamical nonequilibrium and reconnection” and that such a configuration easily leads to braided and twisted field lines with enough free energy to power the corona. Fig. 1.6(a) shows a cartoon of Parker’s braided field lines.

While Gold (1964); Parker (1983a,b) included few comparisons to observations

(mainly due to the lack of high-resolution instruments of the day), observations from the recently-launched High Resolution Coronal Imager (*Hi-C*) seem to support the idea of magnetic braiding. Cirtain et al. (2013) report two observations, in particular, from the *Hi-C* instrument that show braided magnetic fields, supporting the idea of a DC heating mechanism. Something here about Hi-C resolution maybe

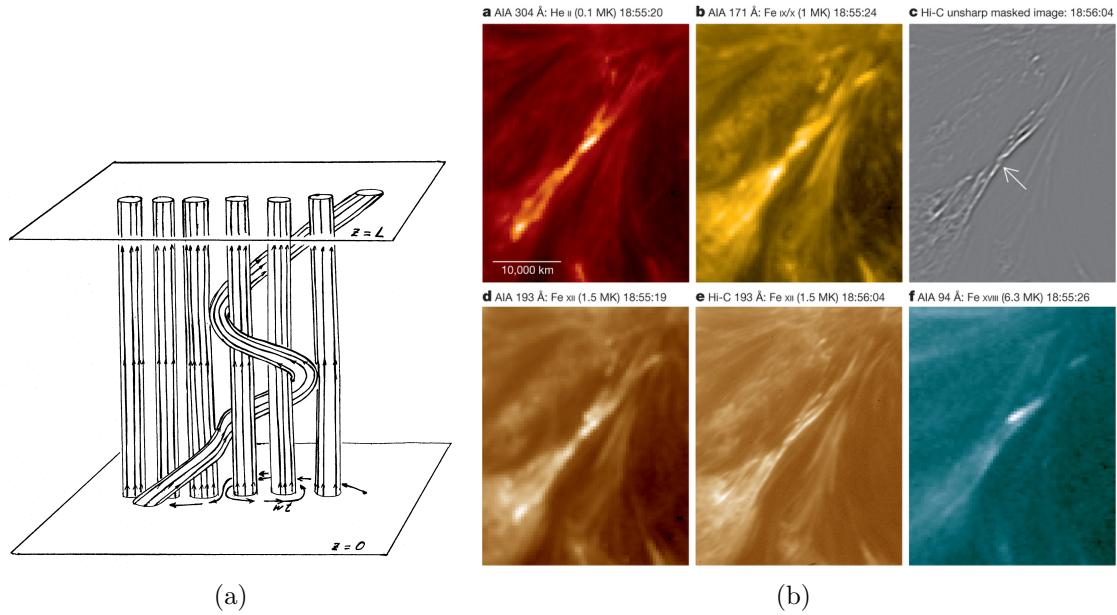


Figure 1.6 : (a) Cartoon showing braided flux tubes rooted, or line-tied, to the solar surface. Braiding allows for the build-up and relaxation of the field, providing an energy release mechanism capable of heating the solar corona. Taken from Parker (1983b). (b) Image from the *Hi-C* sounding rocket mission showing spatially-resolved braided flux tubes in the solar corona. Image taken from Cirtain et al. (2013).

This concept of braided and twisted fields illustrates the basic idea behind DC heating: energy is allowed to build up over time, most likely in the braided and twisted magnetic field as stressed by the turbulent footpoint motions, and then released, most likely through reconnection, as the field relaxes to a near-potential state. Parker

expanded on his early coronal heating theory with the now-seminal (Parker, 1988) in which he proposed that the 2-3 MK temperatures and 10^7 ergs cm^{-2} s^{-1} observed coronal energy input (Withbroe & Noyes, 1977) could be explained by many small, impulsive bursts of energy that he called *nanoflares*. These nanoflares are caused by the many tangential discontinuities that arise in the complex coronal magnetic field, leading to magnetic reconnection and the subsequent dissipation of the mechanical energy stored in the field by the random footpoint motions.

For the field, B , to become twisted and braided, the footpoints of the loop must have some transverse (with respect to the opposite footpoint, see Fig. 1.6(a)) velocity, v and thus a transverse field component, B_\perp . It can be shown that the work done on the field, W , by the footpoint motion is $W \propto v B_\perp B$, such that for $v = 0.5 \text{ km S}^{-1}$ and $B = 10^2 \text{ G}$, $W \approx 10^7 \text{ ergs cm}^{-2} \text{ s}^{-1}$, the observed value, when $\theta = \arctan B_\perp/B \approx 14^\circ$ (Parker, 1988). Using appropriate length scales, the energy per nanoflare event can be estimated as 10^{24} ergs, or about one-billionth of the energy of a typical flare. Thus, once $\theta = 14^\circ$, reconnection dissipates the stored energy and the field relaxes to equilibrium until it is wound up again by the footpoint motions, resulting in “bursty” energy release with the timescale and energy ultimately determined by the underlying velocity field and dissipation mechanism.

The Parker nanoflare concept has become one of the most favored and contentious coronal heating models (Cargill, 1994; Cargill & Klimchuk, 2004; Klimchuk, 2006). While many theoretical efforts (e.g. Bradshaw et al., 2012; Reep et al., 2013) have

shown the feasibility of nanoflares, the idea has long suffered from a lack observational evidence, though recent high-resolution and high-cadence observations (Brosius et al., 2014; Testa et al., 2013, 2014) have provided encouraging results. The term *nanoflare* has now become synonymous with impulsive heating in the energy range $10^{24} - 10^{27}$ ergs, with no specific assumption as to what underlying physical mechanism is responsible for this heating. Thus, in this work, the term nanoflare will refer to bursty energy release on a timescale of approximately 100 s or less.

1.4 Summary

The coronal heating problem is one of the most important unsolved problems in astrophysics. Gaining insight into the complex dynamics of the coronal magnetic field and plasma will allow us to better understand our own Sun and, in particular, how it affects our life here on Earth. Additionally, the field of solar physics has broader applications to fundamental astrophysical processes as studying our closest star will give us a better understanding of how stars throughout the universe behave and evolve. This thesis will focus on the study of coronal heating through the use of hydrodynamic models in which the underlying magnetic field is assumed to be static. In particular, this work will analyze how different heating models affect the emission measure and the resulting observables. This approach, commonly referred to as *forward modeling*, allows one to assess the physical validity of a number of free parameters when compared to observations. §2 will discuss loop structures in the

solar corona with an emphasis on recent observational findings and current modeling techniques. §3 will detail the emission measure diagnostics commonly used in coronal loop physics; in particular, how the line intensities are interpreted and how the differential emission measure (DEM) and other observables are inferred from observations. The details of the efficient hydrodynamic model used in this work will be discussed in §4, with particular emphasis given to the improvements made upon previous models. Finally, §5 and §6 will include the results and conclusions of our study, respectively, with comments regarding future work and improvements included in the latter.

Chapter 2

Coronal Loops

This chapter will discuss the discrete nature of corona in terms of coronal loop structures; Need a section on general plasma dynamics of loops to discuss energy transfer/loss/gain through heating/enthalpy/radiation/draining/filling; Also discuss general structure and how they are formed ;Give some general characteristics about them like length, temperature, density, through what layers they extend etc.; Show nice schematic

As discussed in §1.2.1, magnetic field lines emerge from below the photosphere due to magnetic buoyancy and differential rotation, extending high into the atmosphere with their footpoints rooted in the solar surface. Those field lines that are closed (both footpoints line-tied) form arch-like structures called *coronal loops*, the primary building blocks of the highly structured solar corona (Reale, 2010). Because $\beta < 1$ in the corona (see Fig. 1.4), the plasma is strongly confined such that the flow is directed primarily along the field with negligible cross-field motion. Thus, the enormously complicated magnetic field provides the scaffolding for the coronal plasma. As a consequence of this confinement, loops are individually isolated, leading to the observed inhomogeneity seen in the corona (Reale, 2010). Fig. 2.1 shows several distinct loop structures, both on-disk and off-limb, as observed in the 171 Å channel

of SDO/AIA.

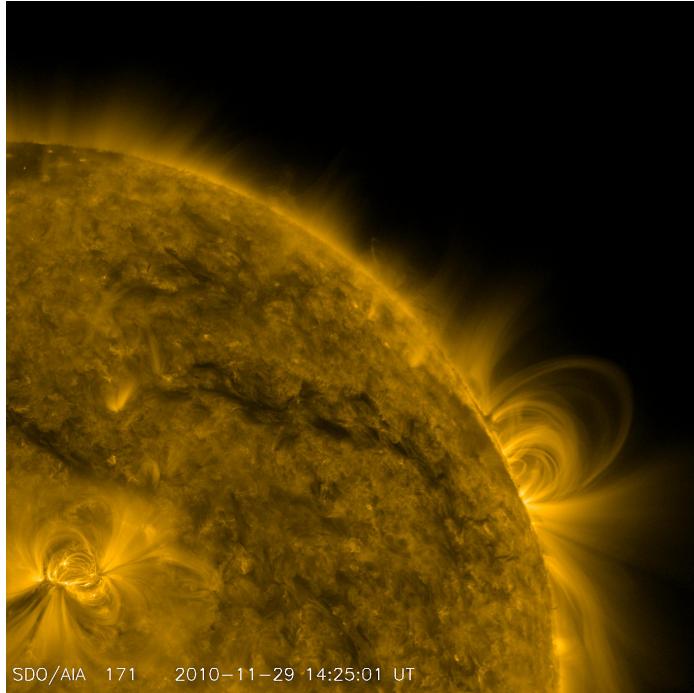


Figure 2.1 : Image captured by the 171 Å channel of the Atmospheric Imaging Assembly (AIA) onboard the Solar Dynamics Observatory (SDO). An arcade of loops in an active region can be seen on-disk in the lower left corner. Several prominent loop structures also appear off-limb on the right side. Image courtesy of NASA/SDO.

Coronal loops, because of their relatively high temperatures ($\sim 10^5 - 10^7$ K), are observed primarily in the extreme ultraviolet (EUV) and x-ray bands, with an associated wavelength range of $1 \sim 100$ Å . While most coronal loop temperatures exceed 10^5 K, loop plasma can span a range of temperatures and densities, due to both the complexity of the underlying field and the fact that individual loops are thermally isolated. As a result, loops are often categorized based on their thermal properties: *cool*, 0.1-1 MK, *warm*, 1-1.5 MK, and *hot*, ≥ 2 MK (Reale, 2010). Whether these thermal categories actually represent distinct classes of loops or if they are all

just transient states of the same type of loop is an open question in the coronal loops community, with the main debate revolving around whether different physical processes are at work in each type of loop. Additionally, loops have also been observed in several topologically different features on the Sun. See Table 2.1 for some typical parameters associated with these distinct regions. This work will focus primarily on simulating loops in active regions, where many hot, nonflaring loops have been observed and emission from loop plasma is unlikely to be contaminated by radiation from the underlying transition region (Tripathi et al., 2011; Warren et al., 2011, 2012; Winebarger et al., 2011).

Table 2.1 : Typical parameters for several different types of loops. Adapted from Reale (2010).

Region	L [10 ⁹ cm]	T [10 ⁶ K]	n [10 ⁹ cm ⁻³]
Bright points	0.1-1	2	5
Active regions	1-10	3	1-10
Giant arches	10-100	1-2	0.1-1
Flaring loops	1-10	> 10	> 50

2.1 Observations

Discuss some observations of loops and what has been learned about them, what constraints, multi-stranded versus single stranded Show some pretty pictures; if coronal loops are heated by nanoflares, what would be the observational signature?

The first evidence of magnetically confined loop structures came from soft x-ray observations by grazing-incidence telescopes aboard rocket missions in the late 1960s as reported by Vaiana et al. (1968). Analysis of data from these early missions

also allowed for classification of distinct topological features on the solar surface by Vaiana et al. (1973) (see Table 2.1). These early findings provided the first look at the x-ray-bright, highly-structured corona; however, the short observing times of these rocket missions prevented systematic studies of these newly-discovered structures. The first sustained coronal loop observations did not come until the Orbiting Solar Observatory IV, equipped with a grazing-incidence x-ray telescope, and the x-ray telescope aboard the *Skylab* space station (Krieger et al., 1972; Reale, 2010). Long observing times allowed for more accurate determinations of loop lifetimes, better comparisons between observations and loop models, and, perhaps most importantly, the finding that the coronal loop is the essential building block of all coronal structures (Rosner et al., 1978).

Need to introduce heating properties, frequency as the parameter we are most concerned about, review arguments for steady versus impulsive heating

In order to make accurate determinations regarding the physical processes and properties of coronal loops, large statistical surveys are needed. However, isolating and analyzing individual loops is incredibly difficult, due primarily to inadequate instrument resolution relative to coronal length scales. In general, loop structures resolved by current observing instruments are considered to be *multi-stranded*; that is, they are composed of many sub-resolution *strands*, where a strand is the smallest loop for which the cross-section is isothermal (Bradshaw et al., 2012). Need something here about active region cores specifically and why we choose to study these

over other parts of the AR

2.2 Modeling

Discuss modeling approaches, hydrodynamics versus magnetohydrodynamics, etc.

2.2.1 Static versus Dynamic

arguments for both; hydrostatic solutions briefly; scaling laws maybe combine MHD and hydrodynamics into single section, briefly mentioning MHD and why we discard the field

2.2.2 Magnetohydrodynamics

2.2.3 Hydrodynamics

2.3 Summary

Chapter 3

Emission Diagnostics

This chapter will discuss emission measure (EM), differential EM (DEM), line intensities etc. and how they are interpreted in observational and modeling contexts

Discuss how this is how we know anything about plasma in the solar atmosphere Can discuss forward modeling as well; this sets us up for the rest of the thesis

In solar physics (and astrophysics in general), all observational data must be collected through *remote sensing* techniques rather than *in situ* measurements due to the great distances and extreme environments inherent to the discipline. This means that atomic data and spectroscopy must be used to infer properties about the coronal plasma, including the temperature and density. To a good approximation, the hot, tenuous coronal plasma is *optically thin* to radiation emitted in the visible, ultraviolet (UV), and x-ray bands. This means that, between the observer and where the radiation was produced, the photons were not scattered or absorbed and reemitted. As a result, the observed radiation contains signatures of the plasma that produced it and has not been polluted by some intermediate process.

3.1 Spectral Line Intensity

As discussed in §1.3, spectroscopy has been a critically important tool in solar physics since the discovery of the million-degree corona. Modern observing instruments and techniques have allowed for the collection of an unprecedented amount of data at increasingly higher spatial resolution and temporal cadence. However, interpreting this spectroscopic data in terms of useful plasma parameters (e.g. density and temperature) continues to pose a challenge to both observers and modelers alike.

The dominant emission mechanism in the high-temperature, low-density solar corona is *bound-bound* emission,



where X represents the atomic species, m is the charge state, h is Planck's constant, and $\nu_{j,i}$ is the frequency of the emitted photon of energy $\Delta E_{j,i} = h\nu_{j,i} = hc/\lambda_{j,i}$, with j and i representing the bound and lower energy states, respectively (Mason & Fossi, 1994). In this process, ion X^{+m} spontaneously decays from excited state j to lower-energy state i , emitting a photon of frequency $\Delta E_{j,i}$. The associated emissivity (power per unit volume) for the transition can then be written as

$$P(\lambda_{j,i}) = N_j(X^{+m})A_{j,i}\Delta E_{j,i}, \quad (3.2)$$

where $N_j(X^{+m})$ is the number density of element X with charge state $+m$ in excited

state j and $A_{j,i}$ is the Einstein spontaneous emission coefficient. Thus, the power for a given spectral line is dependent on both the number of ions of a particular charge state and the number of ions in that particular charge state who are also in an excited state (Mason & Fossi, 1994; Bradshaw & Raymond, 2013). Finally, this volumetric power can be related to the observed line intensity as at Earth,

$$I(\lambda_{j,i}) = \frac{1}{4\pi R^2} \int_V dV P(\lambda_{j,i}), \quad (3.3)$$

where the integral is taken over the entire sphere of radius R , where R is the distance to the observer, and then normalized by the surface area of the sphere.

The next question is of course how these measured line intensities can be related to the properties of the coronal plasma. In Eq. 3.2, $A_{j,i}$ can be calculated from laboratory experiments for a given transition and $\Delta E_{j,i}$ is easily found provided $\lambda_{j,i}$ is known. The remaining quantity, $N_j(X^{+m})$ can be expressed as a series of ratios,

$$N_j(X^{+m}) = \frac{N_j(X^{+m})}{N(X^{+m})} \frac{N(X^{+m})}{N(X)} \frac{N(X)}{N(H)} \frac{N(H)}{N_e} N_e, \quad (3.4)$$

where $N(X^{+m})$ is the number density of element X in charge state $+m$, $N(X)$ is the number density of element X , $N(H)$ is the number density of hydrogen, and N_e is the electron number density (Mason & Fossi, 1994). Note that $N_j(X^{+m})$ has just been repeatedly multiplied by one in order to reexpress it in terms of the following ratios (from left to right in Eq. 3.4): fraction of $+m$ -ions in excited state j , fraction

of X atoms in charge state $+m$, relative abundance of X compared to hydrogen, and relative abundance of hydrogen compared to the number of electrons. The expression can be further simplified through the definition of the relative abundance $\text{Ab}_X = N(X)/N(H)$ and the common approximation $N(H)/N_e \approx 0.83$.

Eq. 3.2, and thus Eq. 3.3, can be further simplified by invoking the *coronal model* approximation. In optically-thin plasmas, it can be assumed that the collisional excitation and radiative decay occur from and to the ground state, respectively; in other words $i \rightarrow g$, with g representing the ground state. Additionally, the processes which determine the excitation level and those that determine the charge state are assumed to operate on disparate enough timescales such that the changes in energy level populations of the emitting ions, occurring on short timescales, can be decoupled from the changes in the charge state, occurring on longer timescales (Bradshaw & Raymond, 2013). Using these assumptions, statistical equilibrium between the spontaneous decay and collisional excitation processes demands

$$N_g(X^{+m})N_eC_{g,j}^e = N_j(X^{+m})A_{j,g}, \quad (3.5)$$

where $C_{g,j}^e$ is the electron collisional excitation rate between g and j (Bradshaw & Raymond, 2013). Plugging Eqs. 3.5 and 3.4 into Eq. 3.2 and letting $N_g(X^{+m}) \approx N(X^{+m})$ yields

$$P(\lambda_{j,g}) = (0.83)\text{Ab}_X\Delta E_{j,g}\frac{N(X^{+m})}{N(X)}C_{j,g}^eN_e^2. \quad (3.6)$$

Defining the *contribution function* $G(T, \lambda_{j,g}) = N(X^{+m})/N(X)C_{j,g}^e$, the intensity integral can be rewritten as

$$I(\lambda_{j,g}) = \frac{(0.83)\text{Ab}_X\Delta E_{j,g}}{4\pi R^2} \int_V dV N_e^2 G(T, \lambda_{j,g}). \quad (3.7)$$

If the plasma is isothermal over the emitting volume V , $G(T, \lambda_{j,g})$ comes outside the integral such that

$$I(\lambda_{j,g}) = \frac{(0.83)\text{Ab}_X\Delta E_{j,g}}{4\pi R^2} G(T, \lambda_{j,g}) \langle \text{EM} \rangle, \quad (3.8)$$

where $\langle \text{EM} \rangle = \int_V dV N_e^2$ is the average *emission measure*. However, in many cases, this isothermal approximation does not hold such that the intensity integral must be expressed as

$$I(\lambda_{j,g}) = \frac{(0.83)\text{Ab}_X\Delta E_{j,g}}{4\pi R^2} \int_V dT \phi(T) G(T, \lambda_{j,g}), \quad (3.9)$$

where $\phi(T) = N_e^2 dV/dT$ is the *differential emission measure* (DEM).

3.2 Differential Emission Measure

The DEM and EM provide information about the temperature distribution of the plasma. In particular, the DEM is a measure of the amount of emitting material at a particular temperature T in the plasma. Thus, the DEM is an observable that any viable theory of coronal heating should be able to predict with reasonable accuracy

(Golub & Pasachoff, 2010). Furthermore, because the DEM is a one-dimensional function, it provides a simple and powerful characterization of the multi-thermal coronal plasma. Unfortunately, calculating these quantities from observed line intensities is difficult because of both mathematical and data availability issues. Looking back to Eq. 3.8, one can see that, provided the line intensities and contribution functions are available over a significant range of temperatures, the emission measure can be calculated algebraically.

3.2.1 Reconstruction Techniques

If the plasma is assumed to be approximately isothermal over the emitting volume, say at temperature T_C , the contribution function can be taken outside of the integral as in Eq. 3.8. Furthermore, if the plasma is assumed to have a constant density over the emitting volume as well, then $\langle \text{EM} \rangle \approx N_e^2 V$. To determine the T_C , the temperature of the plasma, the emission measure for each spectral line is calculated over the appropriate temperature range,

$$\langle \text{EM}(T) \rangle = 4\pi R^2 \frac{I_j}{(0.83)\text{Ab}_X \Delta E_j G_j(T, N_e)}, \quad (3.10)$$

where the contribution function G for a particular spectral line j at temperature T and density N_e can be calculated using the CHIANTI atomic database (Dere et al., 1997; Landi et al., 2013). Typically, this procedure is done for a sufficiently large sample of observed spectral line intensities and the corresponding $\langle \text{EM}(T) \rangle$ curves

plotted on top of each other. Because of the isothermal assumption, these curves should intersect at a common point, $(T_C, \langle \text{EM}(T_C) \rangle)$ (Landi et al., 2002). Of course this isothermal condition is never actually met; thus, all curves tend to meet in a region of intersection. The more multithermal the plasma is, the broader the region of intersection will be and, consequentially, the value of T_C will be less constrained, leading to indeterminate measurements of the plasma temperature. An example for a simulated data set with a very narrow temperature distribution and thus a well-determined value of T_C is shown in Fig. 3.1(a).

Relaxing the isothermal assumption introduces a great deal more complexity. Unlike $\langle \text{EM} \rangle$, $\phi(T)$ cannot be determined algebraically; instead, Eq. 3.9, a Fredholm equation of the first kind, must be inverted to find $\phi(T)$, a task that poses well-known mathematical difficulties (Kashyap & Drake, 1998). One method for performing such an inversion is a discretization of Eq. 3.9

$$I_j = \frac{(0.83)\text{Ab}_X \Delta E_j}{4\pi R^2} \sum_{k=1}^N \phi_k \int_{T_k}^{T_{k+1}} dT G_j(T), \quad (3.11)$$

where j labels the spectral line and $[T_{k+1}, T_k]$ is the temperature interval over which ϕ_k is constant. For each measured line intensity I_j , ϕ_k for $1 \leq k \leq N$ are determined through some minimization procedure (Landi et al., 2012). To calculate the DEM curve, an arbitrary spline or polynomial is first assumed; subsequent corrections at $T_{eff,j}$, the temperature of maximum abundance for the line intensity I_j , are then calculated and a spline interpolation applied to these corrections for the en-

tire temperature range $[T_1, T_N]$. This correction is then applied to initial curve and the corrected curve is used as the initial DEM on the next iteration (Landi et al., 2012). This method of course has the disadvantage that a functional form for the DEM must be chosen *a priori*, biasing the temperature distribution. Additionally, a poor choice of temperature interval in Eq. 3.11 can oversmooth the DEM curve or cause convergence issues. quickly summarize the MCMC method, maybe need new paragraph.

3.2.2 Scaling Laws

From these emission measure distributions, simple observables can be calculated that any viable heating model should be able to produce. The most popular of these observables is the well-known scaling between the emission measure and the temperature,

$$\text{EM} \propto T^a. \quad (3.12)$$

This scaling, first proposed by Jordan (1980), has since been validated by both simulation and observation. A summary of model and observational results for the a parameter can be found in Table 3.1. Looking at the sample emission measure distributions in Fig. 3.1, it can be seen that, for isothermal (Fig. 3.1(a)) and multithermal (Fig. 3.1(b)) plasmas, the emission measure distribution tends to be strongly peaked. When determining the scaling parameter a , it is common perform a linear fit in $(\log T, \log \text{EM})$ coolward of the peak, typically between 6.0 and $\log T_{peak}$, where T_{peak}

is the temperature at which EM_{max} occurs. Observations of the inter-moss regions of active region cores, where the emission is less likely to be contaminated by transition region lines, indicate that $2 \lesssim a \lesssim 5$. Table 3.1 summarizes these findings. This large range of values is most likely due to both uncertainties in the atomic data as well as the variety of techniques used to reconstruct the emission measure distribution (see §3.2.1). These scalings have also been studied in the context of simulation. [mulu-moore,bradshaw,reep](#) summarize briefly. Additionally, Cargill (2014) showed that a can vary from 2 to greater than 5 depending on the heating frequency, whether the heating spectrum follows a power-law, and how the heating scales with the time between successive heating events.

[paragraph about scaling hotward of the peak and why this is important for nanoflare heating](#)

3.3 Forward Modeling

[Talk about forward modeling procedure, why it is important, how it is done, CHIANTI, hydrodynamics, some nice pictures](#)

Table 3.1 : Summary of emission measure scalings from observational and modeling studies. Scalings are typically calculated coolward of the emission measure peak and typically range from 2 to 5. Adapted from Bradshaw et al. (2012).

Slopes, a	Range of $\log T$	Type	Reference
3.26	6.00-6.60	observation	Warren et al. (2011)
2.17		model	
3.20	6.00-6.50	observation	Winebarger et al. (2011)
2.08-2.47 (background)	5.50-6.55	observation	Tripathi et al. (2011)
2.05-2.70 (background subtracted)			
1.60-2.00 (photospheric abundances)	6.00-[6.60-6.80]	model	Mulu-Moore et al. (2011)
2.00-2.30 (coronal abundances)			
1.70-4.50	6.00-6.60	observation	Warren et al. (2012)
1.91-5.17	6.00-[6.30,6.80]	observation	Schmelz & Pathak (2012)
0.58-2.24 (low-frequency nanoflares)	6.00- $\log T_{peak}$	model	Bradshaw et al. (2012)
0.79-3.65 (nanoflare trains)	6.00- $\log T_{peak}$	model	Reep et al. (2013)
$\sim 2 - \sim 7$ (low- to high-frequency heating)	[6.00-6.25]- $\log T_{peak}$	model	Cargill (2014)

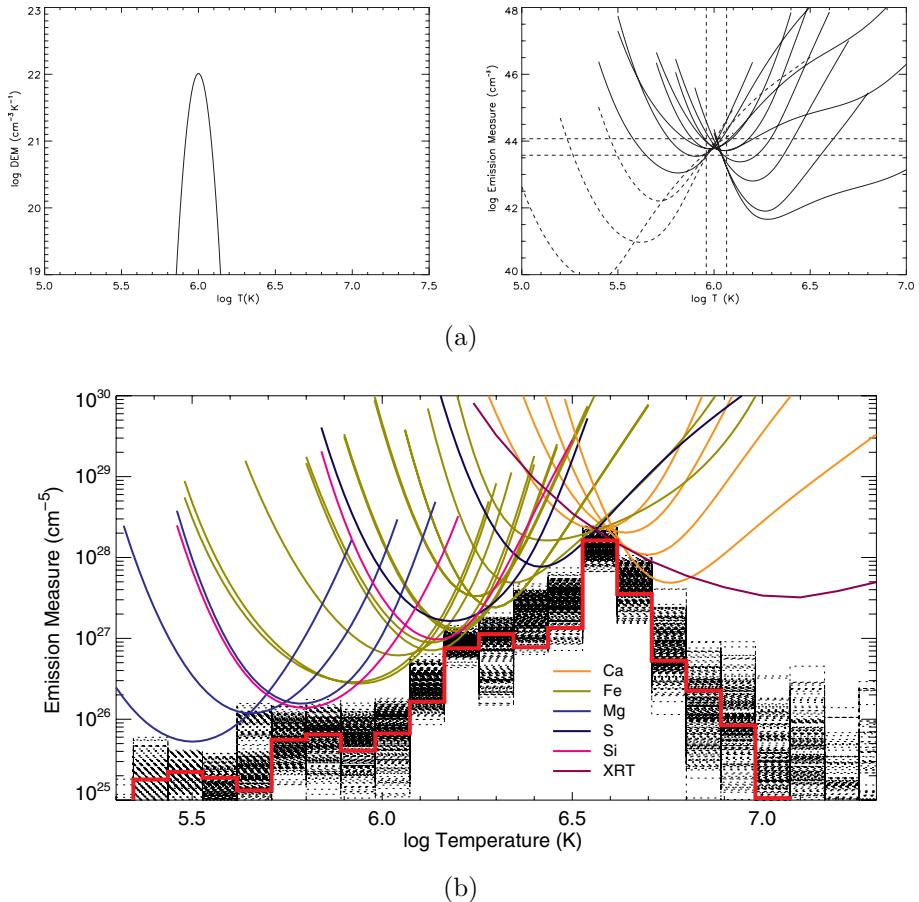


Figure 3.1 : **(a)** Simulated DEM curve and the resulting EM loci reconstruction. From the DEM curve, the plasma is approximately isothermal, being sharply peaked about 1 MK. Correspondingly, the EM loci curves intersect in a region centered on 1 MK, allowing for a well-constrained determination of T_C . Image taken from Landi & Klimchuk (2010). **(b)** Emission measure distribution reconstruction for an active region core. The colored curves show the emission measure calculation using the EM loci method. The black and red histograms show the DEM(T) dT reconstructed using the MCMC method of Kashyap & Drake (1998). Here, the DEM is multiplied by the temperature bin for comparison with the EM loci calculation. Image taken from Warren et al. (2011)

Chapter 4

Numerical Modeling

This chapter will outline our numerical approach to studying the solar corona

Often, the full 3D magnetohydrodynamic (MHD) equations are used when modeling loops in the solar corona. However, such treatments are computationally expensive and are prohibitive when studies of large parameter spaces are needed. Instead, the one-dimensional (1D) hydrodynamic equations are used to model plasma dynamics in coronal loops. These hydrodynamic equations do not include the magnetic field B . This is because the corona These equations are parameterized in terms of time, t , and the field-aligned coordinate, s .

4.1 One-dimensional Hydrodynamics

List the 1D hydrodynamic equations and briefly discuss their origin (i.e. how they were derived) and how they differ from MHD approach Include plasma β justification of why we don't include the magnetic field

The single-fluid 1D hydrodynamic equations are given by,

$$\frac{\partial \rho}{\partial t} = -\frac{\partial(\rho v)}{\partial s} \quad (4.1)$$

$$\frac{\partial(\rho v)}{\partial t} = -\frac{\partial(\rho v^2)}{\partial s} - \frac{\partial(p)}{\partial s} + \frac{\partial}{\partial s} \left(\frac{4}{3} \mu \frac{\partial v}{\partial s} \right) + \rho g_{\parallel} \quad (4.2)$$

$$\frac{\partial E}{\partial t} = -\frac{\partial}{\partial s} ((E + p)v) - \frac{\partial F}{\partial s} + E_H - E_R + \rho g_{\parallel} v, \quad (4.3)$$

where

$$E = \frac{1}{\gamma - 1} p + \frac{1}{2} \rho v^2. \quad (4.4)$$

s and t are the field-aligned spatial coordinate and time, respectively, $\rho = mn$ is the mass density, v is the bulk velocity, μ is the viscosity coefficient and g_{\parallel} is the component of gravitational acceleration parallel to s , E is the energy density and $p = 2nk_B T$ is the pressure subject to closure by the ideal gas law.

Additionally, we also have three terms that are specific to energy balance in a coronal loop: the heat flux $F = \kappa_0 T^{5/2} \partial T / \partial s$, typically given by the classical Spitzer-Harm expression, the ad-hoc heating term E_H , and the volumetric radiative loss term $E_R = n^2 \Lambda(T)$, where $\Lambda(T)$ is the radiative loss function. Emphasize terms that are “non-standard” hydrodynamic terms (e.g. radiation, heat flux, ad-hoc heating, enthalpy transfer) Bring up HYDRAD model and discuss speed limitations

4.2 The EBTEL Model

When using hydrodynamic simulations to study coronal loops, it is often best to compare computed observables with observations in order to make constraints on loop properties. To do this, one must explore a very large parameter space of initial conditions (e.g. loop length, heating frequency, power-law index, etc.,). Though the 1D hydrodynamic equations have been shown to accurately model coronal loops,

Derive the EBTEL equations from the 1D hydrodynamic equations and discuss the physics behind the EBTEL model; what physics is left out? what physical insight can we still gain from this model? Say why EBTEL is important (i.e. large parameter sweeps that can't be done with HYDRAD) Discuss 0D nature, maybe include discussion of previous 0D models and why EBTEL is better Show comparisons between EBTEL and HYDRAD

4.3 The Two-fluid EBTEL Model

Discuss importance of two-fluid effects in hydrodynamic models of the solar corona; include some quick calculations to show how electron and ion fluids can become decoupled Derive two-fluid EBTEL equations from 1D hydrodynamic equations Show several comparisons between HYDRAD and two-fluid EBTEL to justify its use in this study remember to show EM calculation, justification, how EBTEL treats DEM as opposed to HYDRAD Briefly detail solver(s) used, adaptive timestep, why this is necessary, etc. If there is time, discuss speed comparison between EBTEL and

HYDRAD

Chapter 5

Results

This chapter will include the results of our numerical study. Here we will also describe how the study was performed and what tools were used to perform the study. It is best not to introduce any new tools here; just pool from those that have already been discussed and show how they were applied. Show lots of plots and tables.

Chapter 6

Conclusions

This chapter will discuss the conclusions that we can draw based on the results in the results section. What do these results mean? What are the implications in the context of loops in active region cores? May include some topics for future work in this section as well or just wait and put it in a different chapter

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