

ABSTRACT

HENDRICK, SEAN P. Thermal and Non-Thermal Contributions to the X-ray Spectrum of Large Magellanic Cloud Supernova Remnants. (Under the direction of Associate Professor Stephen P. Reynolds).

The study of supernovae and the expanding remnants that form after the explosion is important to understanding star formation and the distribution of elements in the interstellar medium (ISM). In supernova explosions all elements beyond iron are synthesized and then distributed into the ISM, where new star formation is begun in the wake of the explosion. Examining the galactic sample of supernova remnants (SNRs) is limited by absorption and distance uncertainties. The Large Magellanic Cloud (LMC) contains over thirty SNRs at a known distance with little absorption. By examining the X-ray spectrum of a large sample of supernova remnants we can learn how they evolve and how the elements are distributed. This work examines the X-ray spectrum of 14 LMC remnants: three new *Chandra* observations, and eleven archival ASCA observations. The plasma that is responsible for the X-ray emission is heated by a shock wave created by the supernova explosion that expands into the ISM. The picture is further complicated by a reverse shock that is created at the interface between the ISM and the material ejected from the star by the explosion. The ejecta are comprised of elements made by stellar fusion, while the ISM is mostly hydrogen and helium. After the shock passes, the ions contain most of the thermal energy while the electrons are still cold. Coulomb interaction between the ions and electrons behind the shock heat the electrons and continue to ionize the heavy elements, until equilibrium is reached. Non-equilibrium ionization (NEI) models are used to describe the thermal contributions to the spectrum. Existing models assume that hydrogen and helium are the sole source of electrons in the plasma. A new NEI model is introduced which accounts for the overabundance of heavy elements in the remnant ejecta and their contribution to the electron density. At each step in the ionization history of the plasma, the ionization fractions of ten elements are determined. This is a plane-parallel shock model that can account for non-Coulomb heating at the shock front and uses shock velocity as a parameter rather than shock temperature. The heavy element model is tested against the current models and used to fit the *Chandra* observations. Hard, non-thermal tails that are well described by a power law have been observed in the X-ray spectra of several remnants. Synchrotron radiation observed in the radio regime for low energy electrons can be extrap-

olated to X-ray energies to account for high energy (> 1 TeV) electrons. By combining X-ray data with radio observation parameters we can constrain the maximum energies of the shock-accelerated electrons in the supernovae above which the electron spectrum must steepen. None of the remnants in this study have a maximum energy above 100 TeV. The heavy element NEI model has proven a valuable tool that will be released to the community. It determined the forward shock velocities of the *Chandra* observations to be 1000 km/s (0534-69.9), 1100 km/s (0548-70.4) and 1200 km/s (0453-69.5). SNR 0534-69.9 and SNR 0548-70.4 spectra show middle-aged remnants with significant contributions from the ejecta inside the outer shell. The ejecta detection, a suprising result due to the remnants ages, allowed the classification of Type Ia explosions for both. The SNR 0453-68.5 observation indicates a core-collapse explosion due to the presence of a pulsar wind nebula.

**Thermal and Non-Thermal Contributions to the X-ray Spectrum
of Large Magellanic Cloud Supernova Remnants**

by

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Biography

I was born in Philadelphia, PA on June 14, 1973, son to John Francis and Kathleen Elizabeth, and younger brother to 14 month old Kevin Michael. We moved to the suburbs shortly there after, growing up in Oreland, PA, just outside the city limits. Starting at a very young age, my brother and I participated in organized baseball, basketball, and soccer learning the lessons sportsmanship and teamwork along the way (usually with our father as the coach and our mother in the crowd). A common site in my house would be my father sitting and reading science-fiction novels, a trait I reflect to this day. When I was in the second grade I asked him which of his books I should read; he handed me Isaac Asimov's "Foundation Trilogy," and from that moment on I was hooked. In those early years we would pack up for the summer, and move down to Ship Bottom, NJ. I'd spend my days on the beach and my nights looking at the stars. Later, I would spend the summers from 1989–1997 working as a lifeguard for the town of Ship Bottom getting paid to sit on the beach that I loved so much.

After attending public school in Springfield Township (Montgomery County), I attended La Salle College High School in Wyndmoor, PA under the expert tutelage of the Christian Brothers. I was able to academically excel in my years there, as well as participate in a wide range of activities from "Math-letes" to the football team. Besides any academic achievements, I am quite proud of my years playing football and lifeguarding. The reason for this is that I do not walk as most men do; nerve damage caused my right leg not to grow with the rest of me. As a result it is an inch or two shorter than the left leg, and the foot, three shoe sizes smaller, continues to resist my brain's attempts to wiggle my toes or move my ankle beyond a small angle. This causes me to walk with a limp that becomes quite painful after an extended period of time. However, I am quite aware of how lucky I am compared to some, and I never let it hold me back.

Studying astronomy is something I have wanted to do since the night when I first saw Saturn in a telescope. Choosing a college involved simply making a list of schools that offered astronomy as a major. I discovered the University of Virginia, and was enrolled in the Fall of 1991. I immediately began course work in a joint Physics and Astronomy program that would prepare me for graduate school study in astrophysics. Entering UVA with a nearly a year's worth of credit from my high school's great array of AP courses, I was able to complete my studies in the Fall of 1994, with a Bachelor of Arts degree.

From Charlottesville, VA, I continued south to find myself here in Raleigh, NC attending North Carolina State University in the Fall of 1995. It was here that I discovered teaching, working in astronomy labs and the Physics Tutorial Center from 1996 until 2002. I truly enjoy standing before a class trying to impart knowledge, and seeing the light go in the student's eyes that tells you you have succeeded. I hope to continue teaching in the future, and have missed it this past year as I have prepared the document now before you.

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Chapter 1

Supernovae and Their Remnants:

An Overview

1.1 What Is a Supernova?

When a main sequence star runs out of hydrogen fuel in its core, it starts on the path to its final state. Our Sun will expand to a red giant (to the orbit of Venus, about 200 times its current radius) as the helium core shrinks due to gravity and a thin hydrogen burning shell about the core expands. Helium burning will begin in the core creating carbon and oxygen. When the helium burning in the core is exhausted the star will eject its outer H and He burning layers and leave behind a white dwarf (about the size of the Earth). Solitary white dwarf stars slowly cool from 10^5 K to thermodynamic equilibrium with the surrounding space. More massive stars die in violent explosions known as supernovae. It is in these supernova explosions that the heavy elements of the periodic table are formed and dispersed throughout the interstellar medium (ISM). The stellar lifecycle is continued as new star formation begins due to the injection of energy and material into the ISM.

History has observed many supernova explosions over the centuries. Tycho Brahe and Johann Kepler observed explosions in 1572 and 1604, and Chinese observers recorded several events including SN1006 and SN1054 (resulting in the Crab Nebula). The flash

of the explosion appeared as a bright ‘guest star’ that disappeared after a few months, with some so bright they were visible during the day. The optical spectra observed for supernovae distinguish between the types of explosion by the presence of hydrogen lines. Type II explosions have H-lines, Type I explosions do not. The energy source of the bright optical signal is the radioactive decay of nickel to cobalt (6.1 day half-life) and then cobalt to iron (77.1 day half-life). Once that is complete, the optical signal fades away. Type I explosions are further classified as Ia, Ib, and Ic based on their spectra. Type Ia explosions contain silicon lines, Ib and Ic do not. Type Ib explosions are helium rich, while type Ic supernovae are helium poor. For each type of explosion, about 10^{51} ergs of kinetic energy is inferred to be carried into the ISM. A review of supernova explosions and their optical signals can be found in Kirshner (1988).

1.1.1 Type II

Massive stars ($> 8M_{\odot}$) will end their lives via gravitational collapse. When the initial hydrogen burning is complete in the core, a thin shell around the core continues burning hydrogen into helium. This shell creates enough pressure that the core contracts and starts helium burning into carbon. After helium is exhausted in the core a new shell is formed and core carbon burning begins. This process continues until the inner region of the star resembles an onion, as seen in Figure 1.1 (from Kaler, 1992) with shells of different burning processes and an iron core. Since energy cannot be gained in the fusion of iron, the core cannot balance its radiation against the gravity of the rest of the star. When the iron core has a mass of $1.4M_{\odot}$, gravity causes it to decay into neutrons and collapse, for less massive stars forming a neutron star 10 km in diameter. Extremely massive stars ($> 15M_{\odot}$) will further compress the mass of a neutron star to a single point, a black hole. The infalling material from the inner layers rebounds off the core and collides with the still infalling outer layers. The explosion sends the material blasting into space at $\sim 5,000$ km/s, ejecting $5 - 20M_{\odot}$. The ejecta contain the products of the nuclear burning, oxygen, silicon, sulfur, iron, and small amounts of elements heavier than iron produced in the explosion itself. This gravitational collapse releases 10^{53} ergs of energy, 99% of which is carried off by neutrinos. Although many neutrinos are created in the fusion of electrons and protons into neutrons during neutron star formation, about ten times as many are thermally emitted as the neutron star cools from a temperature (kT) of 3-5 MeV.

1.1.2 Types Ib and Ic

The explosion mechanism in Types Ib and Ic is the same as for Type II, the gravitational collapse of a massive star as seen in Figure 1.1. The difference is that the outermost layers of the star have been lost prior to the explosion. For Type Ib, the star has released its hydrogen layer in its stellar wind. Type Ic explosions lack both H and He lines indicating the loss of the two outermost layers. As above, neutron stars, and possibly black holes, are formed as a result of Type Ib and Ic explosions.

1.1.3 Type Ia

Type Ia explosions are from white dwarf stars in binary pairs. White dwarfs are very dense, almost the mass of the Sun condensed to the size of the Earth. When their companions expand to become giants the white dwarfs can tug on their outer layers. Mass transfer occurs through an accretion disk to the white dwarf from its companion, as seen in Figure 6.2 (image taken from <http://www.astro.psu.edu/users/rbc/a1/lec12a.html>). This will eventually push the star over the $1.4M_{\odot}$ Chandrasekhar mass limit. The quantum effect of electron degeneracy pressure can no longer maintain the star, and deflagration (subsonic burning) of the carbon and oxygen results in an explosion, totally destroying the star. $1.4M_{\odot}$ of material, larger iron, is ejected at 10,000 km/s, releasing 10^{51} ergs in kinetic energy, but there is no corresponding neutrino flux.

1.2 Astrophysical Shocks

Each explosion creates a shock wave, a disturbance moving into the surrounding medium highly supersonically. These blast waves propagate into the ISM heating particles, sweeping them along, compressing magnetic fields, destroying dust, crushing interstellar clouds, and accelerating cosmic rays. A shock can be approximated as a discontinuous jump when the mean free path of the particle is much less than the macroscopic length scale. At the jump, the Rankine-Hugoniot jump conditions apply (Shu, 1992, equations 15.31-15.33):

$$\rho_2 u_2 = \rho_1 u_1 \quad (1.1)$$

$$\rho_2 u_2^2 + P_2 = \rho_1 u_1^2 + P_1 \quad (1.2)$$

$$\frac{1}{2}u_2^2 + h_2 = \frac{1}{2}u_1^2 + h_1 \quad (1.3)$$

$$h = \frac{\gamma}{\gamma - 1} \frac{P}{\rho} \quad (1.4)$$

where the subscripts indicate upstream (1) and downstream (2) in the shock frame, and h is specific enthalpy. We shall assume a monatomic gas with $\gamma = 5/3$. A basic picture of a shock front is found in Figure 1.3. Cold, rarefied material is upstream from the shock, and as the shock passes it becomes hot and dense. The ideal gas law can relate temperature to pressure.

$$P = \frac{3}{4}\rho_2 v_s^2 = k(n_e + n_i)T_s \quad (1.5)$$

The pressure term is the ram pressure generated by a shock with speed v_s in the downstream region. A shock at 10,000 km/s heats the plasma to temperatures above 10^7 K. The thermal radiation of the particles extends to the X-ray regime at these temperatures. We assume a shock to be ‘collisionless,’ with no Coulomb collisions at the shock front and magnetic fields mediating the particle interactions. This is a good approximation when the Coulomb mean free path is much greater than the gyroradius of the thermal particles.

The acceleration of particles can occur at the shock front via first-order Fermi acceleration (Blandford & Ostriker, 1978; Jones & Ellison, 1991). Particle diffusion caused by Alfen waves allows some particles to repeatedly cross the shock. The particles increase in energy with each crossing, and some particles escape the shock becoming cosmic rays. The shock accelerates both electrons and ions in the same way, once all particles are extremely relativistic. Particle acceleration is detected in all SNRs by radio emission which indicates acceleration up to GeV energies. X-ray observations can determine if acceleration is occurring up to TeV energies associated with cosmic rays. Chapter 4 of this work uses the X-ray observations to constrain the maximum energy of electron acceleration. If particle acceleration to extremely high energies is common for supernova remnants, then a significant percentage of the 10^{51} ergs of explosion energy could be carried away by those particles.

Another important issue is the creation of a reverse shock. The blast wave travels out to the ISM, driven by the material ejected from the progenitor star. At the interface between the ISM and the ejecta, a reverse shock can form once the outer shock has slowed significantly. This shock will travel backwards, in the Lagrangian frame, through the ejecta that have enhanced abundances of heavy elements from stellar fusion processes. Dwarkadas & Chevalier (1998) calculate how long the reverse shock will travel through the ejecta before reaching the center and dissipating. In Chapters 5 and 6 we will examine how to distinguish between the ejecta heated by the reverse shock and the ISM material heated by the forward shock.

1.3 Supernova Remnants

Possible stellar remnants of a supernova are a black hole, a neutron star, or nothing. In each case, a supernova remnant (SNR) is formed by the expanding shock wave after the explosion, once it has had time to interact with the ISM and ejecta. As more material is swept up by the shock, it decelerates, eventually dissipating over a few $10^5 - 10^6$ years. SNR evolution consists of three stages.

1.3.1 Ejecta Dominated

In the early stage, the freely expanding ejecta dominate the dynamics of the remnant. The reverse shock is prominent, and the ejected mass is greater than or comparable to the ISM mass swept up by the forward shock. The X-ray spectrum will indicate the elemental abundances of the ejecta as well as the ISM. This stage lasts for the first few thousand years of the remnant's life. Examples of ejecta dominated SNRs are Galactic remnant Cassiopeia A, Small Magellanic Cloud (SMC) remnant 1E0102-72.3, and the very young Large Magellanic Cloud (LMC) remnant SN1987A.

1.3.2 Sedov Phase

Sometimes called the adiabatic phase (however, the ejecta dominated stage is also adiabatic), this begins when the mass swept up by the forward shock greatly exceeds the ejected mass. The dynamics of this stage are described by Sedov's self-similar solution

(Sedov, 1959). Figure 1.4 is the picture of a strong shock wave expanding in a uniform medium. Now the X-ray spectrum is indicative of the shocked ISM abundances. The Sedov solution assumes an explosion in a uniform medium of density ρ with energy E . The following equations describe the evolution of the SNR radius R and shock speed v_s as functions of time (assuming the adiabatic index $\gamma = 5/3$):

$$U_{sh} = \frac{2}{5} \frac{R}{t} \quad (1.6)$$

$$R = 1.15 \left(\frac{E}{\rho} \right)^{1/5} t^{2/5} \quad (1.7)$$

where U_{sh} is the speed of the forward shock propagating into the ISM. General age ranges for the Sedov phase are 10,000 to 100,000 years after the explosion. Examples of Sedov phase remnant include the LMC remnants 0548-70.4, 0534-69.9, and 0453-68.5 discussed in this work.

1.3.3 Radiative Phase

Known as the ‘snowplow’ phase of SNR evolution, this begins when the shock temperature drops below 10^6 K. Radiative cooling becomes important, and the shocked material becomes too cool to be observed in X-rays. The forward shock will continue to sweep up ISM material as momentum conservation takes over the dynamics. Eventually the material will become fully mixed into the ISM, and will be incorporated into the next generation of stars.

These three stages of SNR evolution are oversimplified. Transitions between the stages are gradual, and parts of a remnant can be in radiative phase even when global remnant dynamics are still based on the Sedov phase. Likewise, remnants entering the Sedov phase can still contain significant emission from ejecta in the SNR interior.

1.4 Interstellar Medium

What is the composition of the material between stars? The Sedov solution assumes that the ambient medium is of uniform density, which is not the case. There are several different components of the ISM distinguished by temperature and density. The

conditions of the ISM will affect the morphology of the SNR since the shock front will slow down in denser material. Understanding the composition of the ISM near a remnant is an important issue. Spitzer (1978) Chapter 1 gives an excellent review of the components of the ISM. The state of interstellar gas can be defined in terms of temperature:

- Molecular gas – at very low temperatures, $T \sim 10$ K, simple molecules such as H_2 and CO can form. Such clouds are nearly opaque.
- Atomic gas – hydrogen is the most abundant element in the universe, and large amounts exist between the stars. At temperatures of $T \sim 100$ K neutral hydrogen and helium make up a majority of the ISM mass.
- Ionized gas – stars that are strong sources of ultraviolet radiation can provide the 13.6 eV necessary to ionize hydrogen. Ionized nitrogen and oxygen are also observed around the young O and B type stars with high surface temperatures of $T \sim 10,000$ K that photoionize the gas. Higher levels of ionization occur in collisionally ionized gas where temperatures of 10^7 K are generated by a shock wave.

Besides gas, we must consider two other important constituents: dust and magnetic fields. Dust grains with radii smaller than 10^{-4} cm have a strong effect on observations. They are composed of graphite, iron, silicon carbide, silicates, and frozen gas. Dust scatters and absorbs light of all wavelength, with increasing effects against decreasing wavelengths. Interstellar magnetic fields align these grains and polarize the light from objects we attempt to observe through them. The field strength in the ISM fluctuates over both large and small scales, with a mean value of $\mathbf{B} \sim 3 \mu\text{G}$. Magnetic fields are also responsible for synchrotron radiation when interacting with relativistic electrons in the galactic plane.

The different components of the ISM are organized into several different structures, which are regulated by supernova explosions as described in Ikeuchi et al. (1984) and McKee & Ostriker (1977). The hot phase of the ISM exists within the expanding bubbles of old SNRs, each surrounded by a thin dense shell of swept-up material. At temperatures of 10^6 K, the density within these shells is $< 10^{-2} \text{ cm}^{-3}$, and they can expand over ~ 100 pc. These ‘superbubbles’ expand into the warm intercloud medium (WIM). The WIM has temperatures of a few thousand degrees, with a density of 0.01 cm^{-3} . This material exists throughout the Galaxy, filling in the space between the cold dense clouds of the ISM and the expanding hot shells. The dense clouds exist in two main forms. HI clouds are clouds of

neutral hydrogen that are very common. They have temperatures of $T \sim 100$ K, and contain $10\text{-}100 M_{\odot}$ at densities of $\sim 20 \text{ cm}^{-3}$. When these clouds exist near, or surrounding, hot O and B type stars, parts of the HI cloud become photoionized HII regions as described above. Molecular clouds make up the other form of the cold ISM. With low temperatures of $T \sim 10$ K, and densities ranging from $10\text{-}1000 \text{ cm}^{-3}$, or higher, simple molecules can form. Within giant molecular clouds, $R \sim 100$ pc, gravity and magnetic fields mediate new star formation.

Evolution of the ISM is dependent on supernova explosion. New explosions send energy and material to the ISM. Existing clouds are heated and crushed, and new clouds form as the thin dense shells of the hot phase cool down. The components are in rough pressure balance with one another ($n_1 T_1 \sim n_2 T_2$), and they mix and cool to create new stars. It is the inhomogeneity of the ISM within the Galaxy that makes observing local supernova explosions difficult. A general estimate of SN explosion rates predict 1 event each 20 years (Heiles, 1987), and yet the explosion observed by Kepler in 1604 is the last recorded in our Galaxy. This discrepancy is most likely a result of the ISM obscuring events from our view. There is little material between galaxies, so we can turn to the Magellanic Clouds for a full sample of supernovas. There we can observe SNRs interacting with the local ISM with little absorption along our line of sight.

1.5 Broadband Detection of Supernova Remnants

After the initial light from the explosion, the expanding remnant radiates throughout the entire electromagnetic spectrum. In the radio regime, synchrotron radiation from relativistic electrons (energies $\sim 1 - 10$ GeV) spiraling in magnetic fields is prominent. The combination of high-energy particles and compressed magnetic fields make most supernova remnants very bright in radio. The possible presence of a neutron star acting as a pulsar can also be determined by radio observations.

Infrared observations indicate the thermal emission from dust that is heated and destroyed by the shock and newly created dust in the wake of the shock wave. The visible spectrum contains emission from collisionally ionized atoms (SII, NII, OIII, etc) and $H\alpha$ (Balmer series) recombination lines. The presence of Balmer series lines seems to be an indication of a type Ia supernova (Smith et al., 1991). Although no H is present in the

explosion, the shock wave ionizes the neutral H in the ISM to create the Balmer emission. Ultraviolet observations of SNRs show emission from higher ionized states than in the optical, indicating hotter material that is cooling down to optical levels.

This work will concentrate on X-ray observations of SNRs. The highly ionized states created by the shock include hydrogen-like ions of C, N, and O, and helium-like ions of Si, S, and Fe. L-shell and K-shell line emission are observed for heavy elements. There is also a continuum component in the X-ray regime that contributes to the spectrum. Chapter 2 reviews the details of each type of X-ray emission. The ionization stages observed in ultraviolet and optical occur as the metals cool from these X-ray states with post-shock temperatures of 10^7 K. Besides the emission from hot thermal gas, there is some indication of SNRs with non-thermal tails in the hard X-ray regime. With the X-ray observatories placed in orbit over the past 30 years, our understanding of supernova remnants has increased with each new detector. Data sets from both the ASCA and *Chandra* satellites are analyzed herein.

1.6 Summary

In Chapter 2, X-ray observations will be discussed: the physical processes of X-ray emission, the satellites, and the data. A discussion of the current state of spectral modeling will follow in Chapter 3. This will also serve to introduce and test a new model that is appropriate for heavy element dominated shocks. Chapter 4, previously published as Hendrick & Reynolds (2001), is a study of 11 LMC supernova remnants that constrains the maximum energy, E_{max} , of electron acceleration of SNRs. Radio observations of the remnants are combined with archival ASCA observations to find E_{max} . *Chandra* observations of SNR 0548-70.4 and SNR 0534-69.9 are presented in Chapters 5 and 6, respectively. There we will examine the question of ejecta versus ISM emission in those mature remnants. Chapter 7 describes the *Chandra* observation of SNR 0453-68.5, the largest and presumably oldest of our *Chandra* targets. An analysis of these data indicates the presence of a pulsar wind nebula within the remnant. A follow up radio observation is also included to attempt to detect the presence of a pulsar (Gaensler et al., 2003). Chapter 8 contains the discussion and conclusions from this study with recommendations of future work.

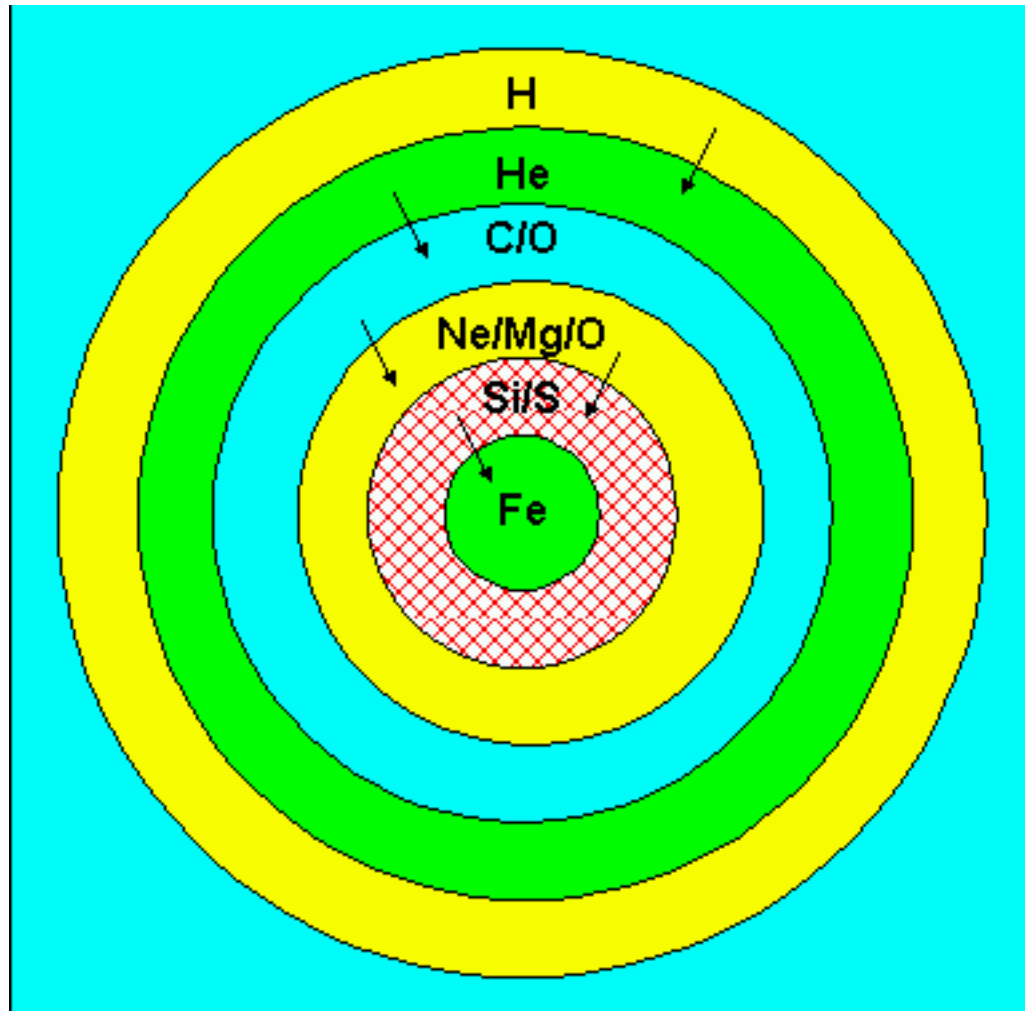


Figure 1.1: Structure of a massive star prior to explosion(from Kaler, 1992).

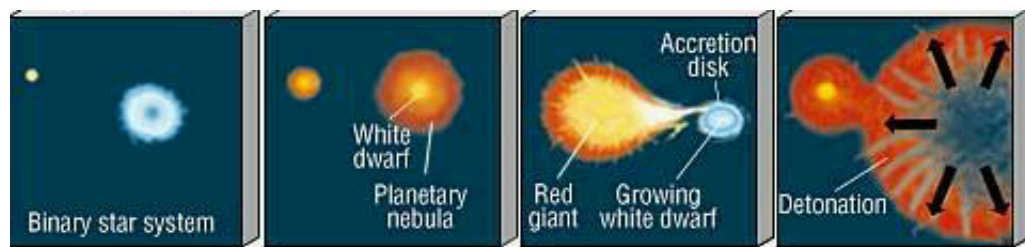


Figure 1.2: Mass transfer between a white dwarf and its companion that results in a Type Ia explosion (image taken from <http://www.astro.psu.edu/users/rbc/a1/lec12a.html>).

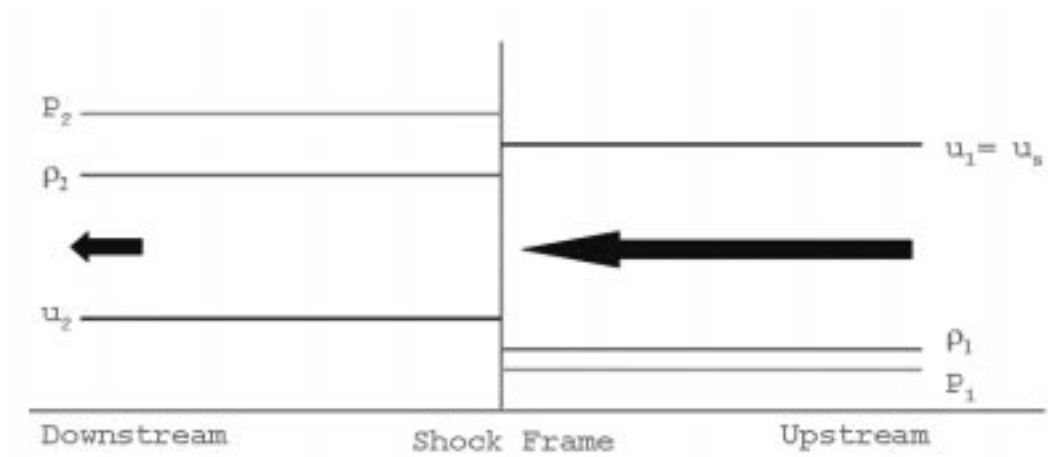


Figure 1.3: Depiction of the RH jump conditions at the shock front.

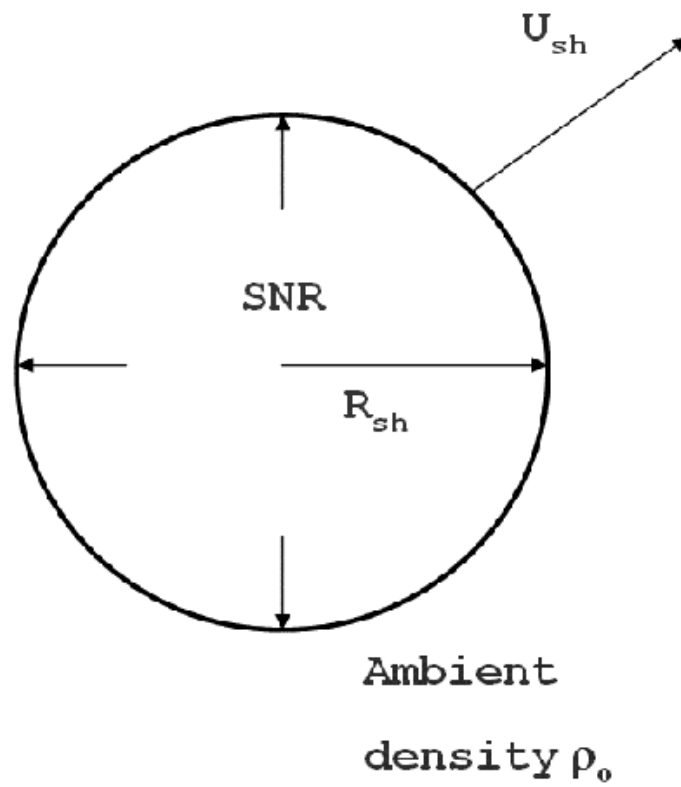


Figure 1.4: Sedov phase SNR, an explosion expanding in a uniform medium.

Chapter 2

X-ray Spectrum: Emission, Detection, and Analysis

2.1 Introduction

X-ray emission from the hot plasma in a supernova remnant result from a combination of several different physical processes. Each process must be accounted for in the data analysis of the X-ray spectrum. Before getting into the analysis of SNR data, we shall review the radiative processes of bremsstrahlung, synchrotron emission, line emission, and inverse Compton scattering in Section 2.2. Photons created by these processes can be highly energetic, with energies from 0.01 keV up to a few hundred keV. X-ray photons must be collected in space since they are absorbed in the Earth's atmosphere. Over the past 30 years, many generations of X-ray telescopes have been put in orbit, each improving on the abilities of the previous missions. Section 2.3 reviews the important historical satellites and the evolution of spatial and spectral resolution to the current generation of observatories. With the data collected and an understanding of the emission processes, we need to process and filter the data before extracting the images and spectra for analysis. Section 2.4 discusses the basic techniques needed to reduce the data to usable form.

2.2 Radiative Processes

2.2.1 Bremsstrahlung

Bremsstrahlung, or braking radiation, is the energy loss of electrons accelerated in the electric fields of passing nuclei. This form of radiation was originally discovered as the explanation for decelerating beams of electrons. This process is identical to free-free emission of an electron making the transition between unbound states of the nucleus. Similar nonrelativistic particles (electron-electron, or proton-proton) exhibit no radiation of this form. It is in the interaction between electrons and ions that bremsstrahlung occurs. Thermal bremsstrahlung indicates a Maxwellian distribution of electron velocity, while non-thermal bremsstrahlung implies a non-Maxwellian distribution, and depends on knowing the actual velocity distributions. There is also the case of relativistic bremsstrahlung. Electrons with energy below 1 MeV are non-relativistic and ionization losses are important. Above 1 MeV is the relativistic regime when relativistic bremsstrahlung losses dominate. Excellent summaries on bremsstrahlung can be found in Longair (1992) chapter 3, Rybicki & Lightman (1979) chapter 5, and Shu (1991) chapter 15.

Figure 2.1 shows an electron moving at velocity v past an ion of charge Ze , where b is the impact parameter. The electrostatic force of $F = Ze^2/4\pi\epsilon_0 b^2$ is experienced between the electron and nucleus, for a duration of $2b/v$. The momentum impulse of this force translates into kinetic energy transfer between the particles. Impulse perpendicular to the electron's velocity is the dominant contributor to the radiation. We must determine the number of collisions within a range of impact parameters to examine the radiation spectrum. From the electron's frame of reference, we get the following equation (Longair, 1992, equation 3.35):

$$I(\omega) = \frac{Z^2 e^6 N}{12\pi^3 \epsilon_0^3 c^3 m_e^2} \frac{1}{v} \ln \frac{b_{max}}{b_{min}} \quad (2.1)$$

where $I(\omega)$ is the total radiation emitted per unit frequency, $\omega = \nu/2\pi$, and N is the number density of the nuclei. Integration over all possible impact parameters would diverge since this treatment does not include the acceleration of the electron from the collision nor does it include a full quantum mechanical description of the interaction. Limits on the range are dependent on the fact that the electron cannot lose all of its energy in the process and that for large values of b other ions would be more likely to interact with the electron. b_{max} is

defined as $\gamma v/\omega$, the limit when radiation is in the exponential tail and doesn't contribute to the intensity. $b_{\min}$ can be calculated in the classical limit for low velocity electrons ($Ze^2m_ev^2/8\pi\epsilon_0$), or in the quantum limit for fast electrons ($\hbar/2m_ev$). The logarithm of $b_{\max}/b_{\min}$ is also known as the Coulomb logarithm, $\ln \Lambda$.

The equation above is for a single electron. We must include a Maxwellian distribution of velocities for the electrons as appropriate for a thermal plasma. For thermal bremsstrahlung the spectral emissivity (radiated power per unit volume and frequency) is (Longair, 1992, equation 3.43):

$$K_\nu = 6.8 \times 10^{-51} Z^2 T^{-1/2} N N_e \exp(-\frac{h\nu}{kT}) \text{ Wm}^{-3}\text{Hz}^{-1} \quad (2.2)$$

This equation demonstrates the exponential drop off of bremsstrahlung radiation due to the low numbers of electrons available in the high velocity portion of the Maxwellian distribution. The energy loss rate for bremsstrahlung is (Longair, 1992, equation 2.45):

$$-\frac{dE}{dt} = 1.435 \times 10^{-40} T^{1/2} Z^2 N N_e \bar{g} \text{ Wm}^{-3} \quad (2.3)$$

The $\bar{g}$ term is the Gaunt factor, defined as the Coulomb logarithm averaged over velocity and frequency (for $h\nu \ll kT$, $\bar{g} \propto \nu^{-0.1}$). For a Maxwellian distribution, this factor ranges in value from 1.1-1.5 for most plasma conditions. The spectrum of the radiation is roughly flat up to $h\nu \sim kT$, above which it begins to drop off. Non-thermal bremsstrahlung needs a new calculation of gaunt factors as well as knowledge of the velocity distribution. For relativistic bremsstrahlung, a correction factor must be added to the equations to account for the new values of $b_{\max}$ and $b_{\min}$ calculated in the relativistic limit. This work will employ non-relativistic thermal bremsstrahlung to define the continuum emission found in SNRs, since most electrons in SNRs have energies far below 1 MeV.

2.2.2 Synchrotron

Electrons accelerated in a magnetic field will radiate as they spiral around the field lines. Figure 2.2 (taken from <http://chandra.harvard.edu/resources/illustrations/>) depicts the physical interaction that creates this emission. The particle's motion is uniform circular motion around the magnetic field. For non-relativistic particles, cyclotron radiation occurs. Cyclotron radiation is a single spectral line of radiation at the frequency $\omega_B = eB/mc$. As the electrons move into the relativistic regime, higher harmonics of this frequency contribute

and $mc \rightarrow \gamma mc$ in ω_B . The electric field of the electrons, which varied sinusoidally with the same frequency ω_B as the gyration in the magnetic field, becomes a series of sharp pulses, repeating for time intervals of $2\pi/\omega_B$, as velocity $v \rightarrow c$. The spectrum now involves many harmonics at integer multiples of ω_B/γ , up to a maximum frequency of $\gamma^2\omega_B/2\pi$. A power law distribution in electron energy creates the power law spectrum observed. This spectrum is the Fourier transform of the sharp pulses in the electric field. Relativistic electrons create time dilation effects in arrival time of the radiation from different points in its orbit, and we must also consider the Doppler shift in frequency. For a single electron, the power per unit frequency is found to be (Rybicki & Lightman, 1979, equation 6.18):

$$P(\omega) = \frac{\sqrt{3}}{2\pi} \frac{e^3 B \sin \alpha}{mc^2} F\left(\frac{\omega}{\omega_c}\right) \quad (2.4)$$

where $F(\omega/\omega_c)$ is a dimensionless function defined in Longair (1994) Table 18.1. The critical frequency above is defined as $\omega_c = (3/2)\gamma^3\omega_B \sin \alpha$, where α is the pitch angle between the magnetic field and the velocity of the electron. At low frequencies this term rises as $(\omega/\omega_c)^{1/3}$, peaking at $\omega/\omega_c = 0.25$, and then drops off exponentially for high frequencies $\omega \gg \omega_c$.

With multiple electrons, distribution in position will create different angles with the B field. Each electron will therefore have slightly different harmonic values. Distribution in electron energy will add further broadening of the lines that creates the continuum spectrum known as synchrotron radiation. Polarization lets us examine the direction of the magnetic field generating the emission. The motion of the relativistic electrons causes the signal to be beamed in the direction of motion. From a distribution of electrons, we see mainly partial linear polarization, in the direction perpendicular to the projection of the magnetic field on the plane of the sky. For power-law energy distributions of electrons in a uniform magnetic field, the maximum polarized fraction is about 70%, and less in more realistic inhomogeneous situations. The fractional circular polarization is of order $1/\gamma$ which is normally negligible for radio and higher frequency synchrotron emission in which we infer γ of order 10^4 or higher.

Combining equation 2.4 with a power law particle distribution of relativistic electrons and integrating over all energies gives a power law spectrum Rybicki & Lightman (1979) equation 6.22a gives the total power as:

$$P_{tot}(\omega) \propto \omega^{-(p-1)/2} \quad (2.5)$$

where p is the particle distribution index, $N(E) \propto E^{-p}$, and the spectral index of synchrotron radiation is $\alpha = (p - 1)/2$. The spectrum rises as $\omega \rightarrow 0.29\omega_c$, and drops off exponentially above that value. Synchrotron emission is dominate in the radio regime, but the power law aspect of the spectrum means the effects should continue into the X-ray regime. This depends on how far the electron distribution extends in energy, which constitutes a possible test of electron acceleration. In the next chapters, we shall examine how different loss mechanisms cause synchrotron emission to eventually drop off to the levels observed in the X-ray spectrum of the supernova remnants that accelerate the electrons.

2.2.3 Line Emission

Line emission occurs when an electron bound to an atom drops from a high energy level down to a lower energy level and a photon is released. This is the type of emission that is common in neon signs, although much more energy is required for the line to appear as X-ray photons. Each ionization state of each element has unique energy levels that create unique spectra. Line emission observed in the visible regime is caused by transitions between the lower electron levels in nearly neutral atoms. Upper level electrons that drop down to the lowest levels are responsible for X-ray line emission. In a hot plasma, most elements are highly ionized, to H- or He-like states. When these ions collide with an electron, a bound electron can be excited out of the lowest energy levels in atom. This creates a hole in the lower level, usually the K-shell or L-shell, that is filled in by an electron from an upper level. Table 2.1 lists the major lines seen in X-ray spectra for some of the most abundant elements (Bearden, 1967). The strength of the line in a spectrum is proportional to the abundance of that element in the plasma. SNRs have varying amounts of this line emission that can be used to determine the chemical composition of the radiating plasma.

2.2.4 Inverse Compton Scattering

Compton scattering occurs when a photon scatters off a stationary electron and the wavelength of the photon increases. Energy and momentum of the photon are transferred to the electron. The inverse is when a low energy photon is scattered by an ultrarelativistic

electron to high energies. Now it is the electron that transfers energy and momentum to the photon. The energy loss rate for a single electron is given by Longair (1992) equation 4.31:

$$-\frac{dE}{dt} = \frac{4}{3}\sigma_T c U_{rad} \left(\frac{v^2}{c^2}\right) \gamma^2 \quad (2.6)$$

U_{rad} is the energy density of the photons incident on the electron and σ_T is the Thomson cross-section. This is similar to the energy loss rate for synchrotron radiation

$$-\frac{dE}{dt} = \frac{4}{3}\sigma_T c U_B \left(\frac{v^2}{c^2}\right) \gamma^2 \quad (2.7)$$

with U_B as the magnetic energy density. This is an interesting parallel between radiative processes. With a power law distribution of electrons, we get a power law photon spectrum (no matter what the original photon spectrum was), with the same spectral index as for synchrotron radiation, $\alpha = (p - 1)/2$.

2.3 X-ray Satellites

Observation in the X-ray regime is a purely satellite based field due to photoelectric absorption of X-ray photons in the atmosphere by molecules and atoms. Even with the satellite in orbit, the issue of capturing the high-energy X-ray photons with a telescope remains. Lenses and mirrors used for optical observation are useless for capturing such high energy photons. However, for sufficiently large angles of incidence, an X-ray photon will reflect off a conducting material. A paraboloid surface is effective, but for 1 keV photons only a small range of angles is focused, with the range decreasing as energy increases. Adding a hyperboloid surface after the paraboloid shortens the focal length and focuses photons in an annulus defined by the surfaces. Nesting paraboloid-hyperboloid surfaces within each other increases the effective area of the telescope. Once we have collected the signal, unwanted events such as cosmic rays, gamma rays, and high-energy electrons released from the walls of the detector by a cosmic ray must be discerned from the X-ray photons. The isotropic cosmic X-ray background adds even more noise to any observation.

There are two main types of detectors used in X-ray astronomy: gas filled proportional counters and solid-state devices. Proportional counters consist of a gas filled capacitor connected to the counter that detects current. The gas chamber, which is sensitive to both

photons and high energy particles, is usually surrounded by another detector that is only sensitive to particles. Therefore, any event that activates both detectors is flagged for removal in processing as a particle event. Within the gas, electron-ion pairs are created. A single pair is created through photoionization by an X-ray photon, while many pairs are created through ionization losses of an energetic particle. With the proper voltage across the gas filled counter, the single electron created by photoionization ionizes other atoms, creating one additional pair for each 30 eV of the incident photon energy. Therefore, the strength of the pulse created is proportional to the energy of the incident photon. Position can be determined by the ratio of the charge at each end of the anode and by using multiple anode wires.

Solid-state detectors improve on the statistical accuracy of gas filled detectors. In this case, electron-hole pairs are created in the material (usually silicon) which needs only ~ 3 eV to create a hole. A detector chip consists of sensitive strips separated by thin channel stops. Those strips are further broken up into pixels, each connected to three read-out electrodes. A detector is an array of several hundred of these pixels that allow spatial position of the photons to be recorded. By creating high resistance across a pixel, the holes are pulled into the substrate material while the electrons liberated by photoionization are trapped in the potential well of the pixel. Charge coupled devices (CCDs) are the most widely used detectors on the current missions. Charge coupling refers to the method of reading out the charge by changing the voltage on the read-out electrodes that shifts the charge to the next channel. This process continues across the device until the charge reaches an amplifier which records the charge of each pixel into memory. The efficiency at which the charge is transferred is a very important issue that must be addressed when working with CCD data. The charge contained in each pixel is proportional to the X-ray flux density at that position. Longair (1992) contains an excellent review of all these telescope related issues.

2.3.1 Historical Notes

The UHURU satellite, the first orbital X-ray observatory, was launched in 1970. It consisted of two proportional counters sensitive between 2 and 10 keV. Although it had poor angular resolution, UHURU was able to map the sky. It detected many variable X-ray sources that are now known to be accretion disks around neutron stars. The next important

mission was the Einstein X-ray Observatory in 1978 (Giacconi et al., 1979), which had five instruments. The most notable of these was the High Resolution Imaging Telescope (HRI). This instrument employed grazing incident optics to increase angular resolution to 4 arcsec with a field of view 25 arcmin in diameter. Deep surveys of small areas led to the discovery of discrete X-ray sources in the Andromeda Galaxy (M31) and the Magellanic Clouds. In 1990, ROSAT (Truemper, 1982) was launched to perform a survey of the entire sky in soft X-rays. The catalog created by the ROSAT All-Sky Survey (RASS) is still used for the planning of new observations. It was 1000 times more sensitive than UHURU and discovered 150,000 objects. It also discovered the detailed morphology of SNRs in X-rays.

2.3.2 ASCA

The Advanced Satellite for Cosmology and Astrophysics (ASCA) was launched in 1993 and operated for seven and a half years (see Tanaka et al., 1994). Archival ASCA datasets were used in Chapter 4 of this work to examine 11 supernova remnants. It was the first X-ray telescope to use CCDs among its four instruments. ASCA contained two gas imaging spectrometers (GIS2 and GIS3) that contained imaging gas scintillation proportional counters, and two solid-state imaging spectrometers (SIS0 and SIS1). 120 nested gold-coated aluminum foil surfaces in the telescope were able to attain an angular resolution of 0.5 arcmin in each detector. Both ROSAT and Einstein HRI detectors had higher spatial resolution, but the strength of ASCA was the high spectral resolution. The CCD devices were sensitive between 0.4-12.0 keV with an energy resolution of about 2% at 5.9 keV, while the GIS could detect between 0.8-12.0 keV, with slightly poorer energy resolution of 8%. ASCA discovered non-thermal X-rays from the Galactic remnant SN 1006, providing a link between the cosmic rays accelerated in that remnant (Koyama et al., 1995) and non-thermal X-rays that is explored in Chapter 4. It could also probe the chemical abundances within distant objects, determining that clusters of galaxies contained material consistent with origin from Type II supernova explosions (Loewenstein et al., 1994).

2.3.3 *Chandra*

Originally called the Advanced X-ray Astrophysics Facility (AXAF), the satellite was renamed in honor of Subrahmanyan Chandrasekhar whose work was mentioned in

Chapter 1. Launched in 1999, the *Chandra* X-ray Observatory has the highest level of spatial and spectral resolution yet. The improvement of the previous generations of telescopes can be seen in Figures 2.3 and 2.4 that compare an image of LMC SNR 0534-69.9 made by *Chandra*, with a spatial resolution of 0.5 arcsec, to an archival image from ROSAT. The instruments include the High Resolution Camera (HRC) and the AXAF Charge-Coupled Imaging Spectrometer (ACIS), as well as a low and a high energy transmission grating. Both contain arrays to be used for imaging, with fields of view of 30' and 16' respectively, and longer arrays, 7'x97' and 8'x48', to be used as readout for the gratings. A grazing incidence iridium-coated imaging telescope focuses the X-ray photons. Our observations with *Chandra* employ the ACIS array, which contains 8 front illuminated chips and 2 back illuminated chips. The focal point of the ACIS-S array is within the back illuminated S3 chip. The resolution falls off with distance from the focus, but the LMC remnants observed and reported on in this work fall completely within that chip. One problem with *Chandra* is the issue of charge transfer efficiency within the CCDs; it is inefficient due to radiation damage suffered early in the mission. It is less of a problem with the S3 chip and a correction for it has yet to be standardized. The issue of charge transfer inefficiency (CTI) is discussed in detail on the *Chandra* X-ray Center's webpage, <http://cxc.harvard.edu/>, but this work will not consider it.

2.4 X-ray Data

Data from X-ray telescopes are often stored as events lists. Each photon or particle event is recorded with its energy, time of arrival, and location. Each event is given a grade that helps discern a particle from a photon event. The ground based teams that run the telescopes process the raw data from the satellite and make them available to the observers. Besides events lists, secondary files are generated for each observation to assist in the data reduction. When cleaning the data we must account for the effects of the Earth's magnetic field (i.e the South Atlantic Anomaly), the Sun, and the telescope itself. Good time intervals are created to eliminate times in the satellite's orbit when the Earth or Sun may interfere. Another important product is a list of bad pixels which were known on launch or have degraded over time. Constant calibration is performed for active missions to measure and account for any changes over the lifetime of the telescope. When an observation is performed

and the data files received, the first task is to clean and filter the raw data. Bad pixels, bad times, and bad grades must all be removed. Appendix A summarizes the data processing and reduction techniques for ASCA and *Chandra* event files.

With the cleaned events file created, we can now extract images and spectra from the data. Different filters can be applied (energy, time, region) depending on the nature of the investigation. Here we are mainly concerned with spectral analysis. We need to create an image and select the regions of interest. By examining the images in Figure 2.3 we can see that the ASCA telescope failed to resolve the remnants, requiring us to use the entire SNR emission as our region, while the resolution of *Chandra* lets us choose from several parts of the remnant. We also must define a source-free background region to account for the isotropic X-ray background radiation as well as charged particles and other instrumental background. We can extract spectra from both the source and background regions, but we're still not able to analyze the data. Response matrices must be generated. A Redistribution Matrix File (RMF) relates detector channels to true photon energies, calibrating the energy axis of the spectrum. We also need an Auxiliary Response File (ARF) which relates the strength of an event to the actual received photon intensity, calibrating the intensity axis. By multiplying the data by these matrices we get a distribution of counts that would occur in a detector of infinite energy resolution. Next we need to rebin the spectrum to a minimum of 15 counts per bin for statistical reasons and subtract the background spectrum. Now analysis of the X-ray spectrum can begin utilizing the XSPEC data analysis software package (Arnaud, 1996).

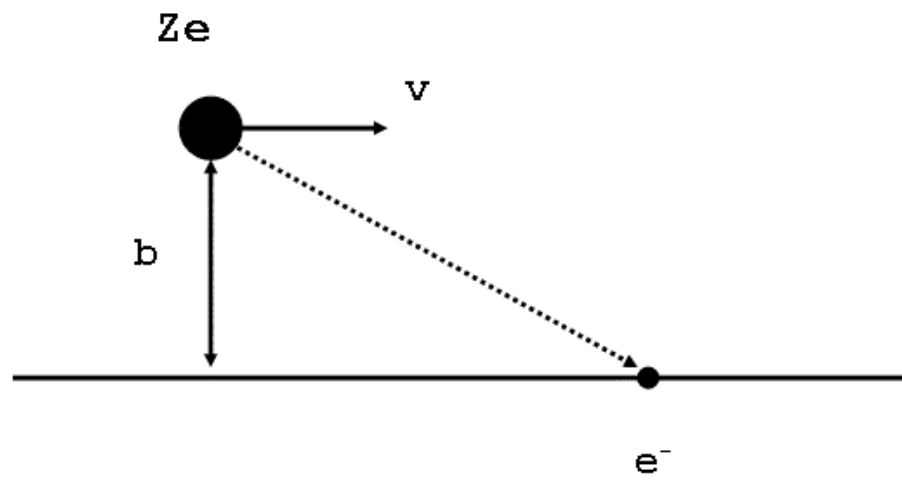


Figure 2.1: Physical process of bremsstrahlung radiation, from the electron's frame of reference.



Figure 2.2: Physical process that generates synchrotron radiation (taken from <http://chandra.harvard.edu/resources/illustrations/>).

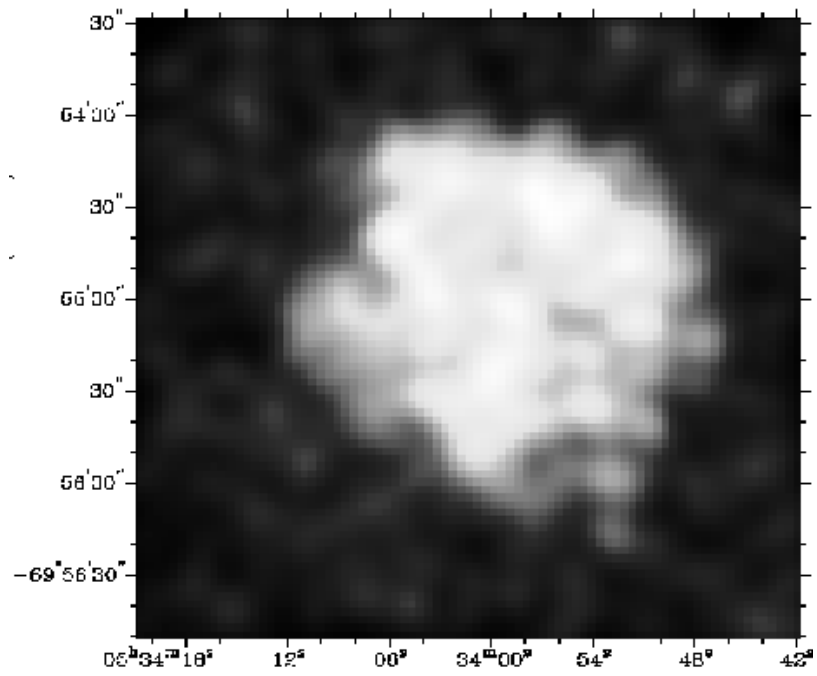


Figure 2.3: ROSAT image of SNR 0534-69.9 (from Williams et al., 1999).

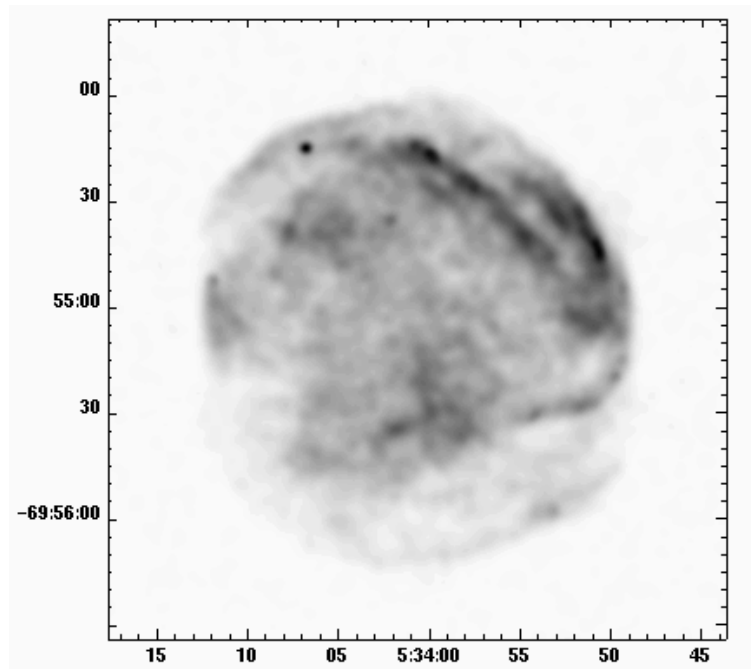


Figure 2.4: *Chandra* image of SNR 0534-69.9.

Table 2.1: X-ray line emission in keV (Bearden, 1967) for neutral atoms

Element	$K\alpha$	$K\beta$	$L\alpha$	$L\beta$
C	0.277			
N	0.392			
O	0.525			
Ne	0.849			
Mg	1.25	1.30		
Si	1.74	1.84		
S	2.31	2.46		
Ca	3.69	4.01	0.341	0.345
Fe	6.40	7.05	0.705	0.718
Ni	7.48	8.26	0.851	0.869

Chapter 3

Spectral Modeling

3.1 Introduction

With the current generation of X-ray satellites (*Chandra* and XMM) we are seeing the detailed structure of hot plasma present in supernova remnants. Spatially resolved spectroscopy that began with the ASCA mission is becoming a crucial tool in understanding the structure and chemical composition of the emitting material. It is the electrons in the plasma that are responsible for the emission of X-ray photons and the observed spectrum either through free-free radiation (bremsstrahlung), or by collisional excitation of atoms. Knowing the electron temperature is important for X-ray emission calculations. Hamilton et al. (1983) performed these calculations for full Sedov models and found that the shape of the X-ray spectrum depends on the shock speed v_s , the ionization timescale τ (defined as $n_e t$, the postshock electron density times the remnant's age) and the amount of electron heating at the shock front. Also important are the abundances of heavy elements for the determination of thermal continua and the strengths of line emission. Sedov models assume that swept up ISM material causes the emission, where hydrogen and helium are the only contributors to the electron density. Young, ejecta-dominated SNRs such as 1E0102-72.3 and Cassiopeia A, have X-ray spectra in their interiors that show abundance ratios of the elements far from the solar ratios found in the ISM of either the Galaxy or the LMC. These elevated levels of heavy elements are currently unaccounted for by the current generation of

models. A new set of emission calculations is performed that accounts for the contributions of electrons from heavy elements to the electron density and temperature as would occur in SNR ejecta emission.

The equilibration of both temperature and ionization is extremely important in understanding X-ray spectra of SNRs. Ionization equilibrium occurs when the processes that create an ion and destroy it are balanced. This would simplify spectral calculations since ionization states would be constant. In a shock heated plasma this is not the case as the newly shocked ions are collisionally ionized to higher states. It will take a very long time for ionization equilibrium to be achieved. Therefore, a state of nonequilibrium ionization (NEI) exists in supernova remnant shocks (Itoh, 1978, 1979; Gronenschild & Mewe, 1982). To accurately account for emission in SNRs we must track the evolution of ionization states in the regions behind the shock. For this reason, an ionization timescale parameter $\tau = n_e t$ is used to define the time since a fluid element has been shocked. The electron density is important since dense material means more electron-ion collisions that would achieve ionization equilibrium faster. Thermal equilibrium assumes that ion and electron temperatures are the same. As the shock heats the particles, ions are initially heated more than electrons, unless some form of collisionless electron heating exists at the shock front. Ions and electrons start moving towards equilibrium via collisions behind the shock. Achieving equilibrium in ionization states and in temperature takes time, and in the presence of astrophysical shocks it is unlikely that either condition exists in a young supernova remnant.

X-ray spectroscopy of SNRs provides information about the condition of the emitting plasma including temperature, chemical composition, distribution, and density. Shock speeds, age, and explosion energy can also be determined from X-ray observations. We can use spatially resolved spectroscopy to investigate different regions of a remnant to examine spectral differences from the outer limb to the interior. We not only find information about the SNR, but also about the local ISM in which it is expanding. Several tools have been developed, ranging from the simple to the very complex, to help extract this information from X-ray data. Equilibrium models assume the thermal and ionization equilibrium discussed above. By assuming a single temperature and ionization equilibrium, such models cannot accurately describe the condition of material behind the shock. The most common NEI models current available assume a constant temperature and a single ionization timescale, which is accurate for a fluid element within the plasma but cannot fully describe the

emission. Non-temperature equilibration models assume $T_e \neq T_i$, but still have a single ionization timescale. Plane-parallel shock models improve over the NEI models by including a full range of ionization timescales from $\tau = 0$ directly behind the shock up to its ionization age $\tau = n_e t$. Some plane shock models are characterized by a constant electron temperature, but there are also plane shock models that include the Coulomb heating of electrons at the shock front. The next step in complexity is to add the spherical symmetry seen in SNRs. The Sedov-Taylor self-similar solution is an idealized structure that is a good choice for spherical NEI calculations. Emission calculations are done for a full range of ionization timescales, and the expansion of the fluid elements behind the shock is included. Finally, to fully understand the structure and X-ray emission of SNRs, we would need hydrodynamical modeling, possibly including shocks modified by cosmic ray acceleration (Jones & Ellison, 1991), and X-ray emission calculations. This is impractical in most cases due to the complexity of the calculation.

One thing the current generation of models have in common is the assumption that the sole contribution to the electron density is from the full ionization of hydrogen and helium. When investigating the emission of supernova remnant ejecta emission, this may not be a safe assumption. The forward blast wave sweeps up the ISM material that is dominated by H and He, clearing the region around the former star. The stellar material created by fusion processes prior to the explosion, and in the explosion itself, is much more abundant in C, O, Mg, Si, S, and Fe. This material, ejected in the supernova, will be shock heated by a reverse shock travelling backwards (in the Lagrangian frame). Electrons from the heavy elements are injected into the plasma by successive ionizations from collisions with the electrons already present. The injection of cold electrons from these elements is important for understanding the evolution of electron temperature behind the shock. We present a new plane-parallel shock model that includes heavy element electron contributions to X-ray spectra of SNRs. At each step over a full range of ionization timescales, this model calculates the ionization states for each element and the electron temperature, and determines the contributions to the X-ray spectrum. This model is tested against existing NEI and plane-parallel shock models. The full implementation of the heavy element shock model with the latest version of the spectral analysis software XSPEC should provide observers with a powerful new tool to investigate supernova remnant ejecta emission.

3.2 Current State of Modeling

3.2.1 Absorption

The X-ray spectral analysis package XSPEC contains many models useful for analysis, and allows for external models to be linked to it. Before reviewing models of thermal emission found in XSPEC, we need to discuss some other model components necessary for a complete analysis. Whenever we are observing a distant object, we must account for the amount of interstellar absorption along the line of sight. This absorption is primarily due to photoionization of trace heavy elements, but is customarily described in terms of a column density of hydrogen N_H , with heavy elements assumed present in solar ratios. In the Galaxy, there are large amounts of dust and hydrogen clouds that absorb X-ray photons. A standard absorption model is the Wisconsin absorption model (Morrison & McCammon, 1983) *wabs* (this assumes solar abundances along the entire line of sight). It was used in Chapter 4 for the absorption to the LMC, but that was not completely accurate. The LMC has abundances 0.3-0.4 solar, and *wabs* will overpredict the absorption. A more accurate estimate of absorption is employed in Chapters 5, 6, & 7, where we used a two component absorption model. One component (XSPEC model *phabs*) is for galactic absorption determined by observations of diffuse neutral hydrogen in the Galaxy by Dickey & Lockman (1990). From their work, we find that the galactic absorption column density towards the LMC is $N_H = 7 \times 10^{20} \text{ cm}^{-2}$. We then used a similar absorption model, *vphabs*, with its variable abundances set to LMC levels, to fit the absorption in the LMC.

3.2.2 Possible Non-Thermal Component

The most important non-thermal emission component is synchrotron radiation. The spectrum generated is well described by a power law until frequencies are reached that correspond to the maximum electron energies present. The maximum energy that electrons can be accelerated to by the shock depends on age and escape. The time for particle acceleration must be less than the remnant's age or the electron's radiative loss time. Both factors can cut off the spectrum by limiting the maximum electron energy. The first-order Fermi acceleration at the shock depends on particles crossing the shock several times, and some particles will escape this process entirely in the absence of magnetohydrodynamic

waves that scatter the electrons back towards the shock. Reynolds (1998) describes two models, *srcut* and *sresc*, that take information of radio observations of synchrotron emission, and extrapolate across the electromagnetic spectrum to the X-ray regime. The radio spectral index and the flux at 1 GHz are used to estimate the level of synchrotron emission in X-rays. By comparing the expected emission levels of the power law extrapolation to the observed X-ray emission we can determine the frequency at which the cut-off began. The models are modified power laws that include an exponential cutoff. These issues are addressed for 11 LMC remnants in Chapter 4, where the *srcut* model is used to determine maximum electron energy.

3.2.3 Non-Equilibrium Ionization

Now let us discuss the ionization states of the elements heated by the shock wave. Masai (1984) and Hughes & Helfand (1985) employ a matrix solution to the problem of nonequilibrium ionization. They assume that electrons are removed one at a time, ignoring the possibility of several ionizations occurring at once. By including the ionization and recombination rates, and assuming a constant electron temperature, a tridiagonal matrix $\mathbf{A}$ is generated. This matrix describes the state of each ion of each element at that temperature and a specific ionization timescale. Borkowski et al. (1994) has an excellent review of these calculations in Appendix A of that paper. The basic equation to be solved is:

$$\frac{d\mathbf{f}}{dt} = n_e \mathbf{A} \cdot \mathbf{f} \quad (3.1)$$

where $\mathbf{A}$ is the tridiagonal matrix defined as:

$$A_{ij} = \begin{cases} -[C_i(T_e) + \alpha_{i-1}(T_e)], & j = i; \\ +C_{i-1}(T_e), & j = i - 1; \\ +\alpha_i(T_e), & j = i + 1; \\ 0, & |j - i| > 1. \end{cases} \quad (3.2)$$

For each step in ionization timescale, the values for the ionization fractions $\mathbf{f}$ are determined based on the ionization rate of electron collisions $\alpha_i(T_e)$ and recombination rates $C_i(T_e)$. Within the current generation of models, these rates are based on a single electron temperature, and are only used to provide measures of line emission from the elements. A

constant electron density of n_e is assumed to exist within the plasma that is not affected by the number of electrons added by these ionizations.

3.2.4 Thermal Equilibrium

The lack of thermal equilibrium in the plasma complicates the analysis of X-ray spectra from supernova remnants. In a plasma heated by a ‘collisionless’ shock, the kinetic energy of the ions is dissipated into random thermal motions by interactions mediated by the magnetic field. The question of electron heating, however, remains open. In the case of a high Mach number shock, electrons and ions both have their kinetic energy randomized. This results in a ratio of electron to ion temperature that is far from equilibrium ($kT_e \sim m_e v_s^2$, and the same for ions, means that $T_e/T_i \sim m_e/m_i \sim 1/1836$). There remains the possibility of some form of collisionless heating of electrons at the shock front that could raise the temperature ratio all the way to equilibration. Two extreme cases are that full equilibration occurs at the shock front, or that no heating of electrons occurs and the electrons are heated solely by Coulomb collisions in the postshock region. Hamilton et al. (1983) examined both these cases in their theoretical calculations, while Hughes et al. (1998) applied these extremes to LMC SNR data and found better agreement with the models that have no collisionless heating of electrons. Real shocks, however, may fall between these extremes. There are indications that collisionless heating is more important in slow shocks and less important in fast shocks. To model the spectrum generated by a shock heated plasma, it is necessary to take into account a range of possible values for the initial T_e/T_i ratio.

3.2.5 Shock Models

A good description of emission from mature remnants comes from models based on Sedov-Taylor dynamics. The Sedov solution begins with an explosion in a uniform medium that expands with spherical symmetry. This only applies for remnants in which swept up ISM material greatly outweighs the ejected matter. Calculations of emission for these objects have been done many times, with the most extensive being Hamilton et al. (1983). Other calculations include Kaastra & Jansen (1993); Borkowski et al. (2001). Updated atomic data for H-like ions of each element (Callaway, 1994) and L- and K- shell emission

from Fe ions are included (Liedahl et al., 1995). The focus of this work is on emission from ejecta, but Sedov models are still very useful in determining the emission from the limbs of a remnant where H and He are still the dominant elements. The X-ray spectrum of ejecta must be viewed through the outer shock wave that is heating the ISM material around the ejecta. Sedov models are parameterized by the ionization timescale τ (defined as post shock electron density n_e times the remnant's age t_o), elemental abundances, post-shock temperature T_s , and post-shock electron temperature T_e . The Sedov model employed by Borkowski et al. (2001) consists of intensive calculations that are very time consuming. To make this model more efficient, a table of models can be pre-calculated over a grid of the parameters mentioned above, with fitting done by interpolation among these models. Sedov models account for the evolution of each fluid element in the plasma as it expand over time resulting in decreasing mean temperature.

Full Sedov models, which are quite complex and descriptive of SNR dynamics, are the most detailed solution to the question of nonequilibrium ionization. The simplest solution is found in single electron temperature, single ionization timescale NEI models which are used quite commonly. A single ionization timescale model will not include line emission from under-ionized material near the shock (at low τ). Contributions from those early ionization states are ignored. The shock wave is still expanding and shocking new material in a SNR, so a full range of ionization timescales is a necessity for a full description of the X-ray spectrum. A single NEI model can effectively define the emission of a single fluid element within a remnant, but should not be used to describe the integrated emission from all elements. Some current authors combine several NEI models to get a range of ionization timescales and electron temperature. Borkowski et al. (2001) find that 10 such zones are required to describe low temperature Sedov models, with fewer needed at high temperatures. The addition of more zones will improve this agreement, with absolute agreement occurring as the number of zones approaches infinity. This is an impractical situation; although the spectrum may be well defined, we learn little about the SNR. We need a physically reasonable model that is computationally efficient and descriptive of the plasma.

Plane-parallel shock models simplify the calculations from the details of a spherical blast wave in Sedov model. With the advent of spatially resolved spectroscopy this approximation is very useful. By examining only a small section of a remnant, we do not need to assume spherical structure. A range of ionization timescales from zero behind the

shock to the shock's ionization age $\tau = n_e t_s$ (t_s is the shock age, which differs slightly from the remnant age in Sedov models, $\tau_{shock} = 0.404\tau_{Sedov}$) is used in these models, with the ability to set the lower limit of τ other than zero. A lower limit allows investigation of spectral differences between fluid elements in different regions behind the shock. As with each model discussed so far, a constant value of electron temperature is assumed in the entire post shock region. These models perform well when compared with full Sedov models, with differences only seen at low energies. Versions of plane-parallel shock models can include electron heating at the shock front with the inclusion of a mean shock temperature parameter T_s . Borkowski et al. (2001) find good agreement with Sedov model values of ionization age and $\langle T_e \rangle$ using the multitemperature plane shock model. High electron temperature Sedov models can be described well, but the models differ for low velocity shocks. The electron density in plane shock models is based on the full ionization of H and He at the shock front. Even in the case when H and He abundances are set to zero, these plane shock models assume that some n_e exists in the plasma. A new plane shock model with detailed counting of the electron density at each timestep, based on the evolution of the ionization fractions, and a range of electron temperatures is introduced in the next section to improve the accuracy of plane-parallel shock models.

3.3 Heavy Element Dominated Shocks

The limitations in the current class of NEI models are the single electron temperature values as well as the assumption that hydrogen and helium are the sole source of electrons in the plasma. In order to address these problems we present a new plane-parallel shock model appropriate for heavy element dominated plasmas. Contributions to the electron density from heavy elements are added to the electrons from hydrogen and helium to get a more complete view of the evolving plasma. This requires calculating the electron temperature at each time interval, determining the ionization state of each element at that temperature, and counting the number of cold electrons added by the new ionizations. Knowing the electron temperature and ionization states of the heavy elements lets us generate the X-ray spectrum of that component of the plasma. The process is repeated for each time step, so that the final spectrum is the sum of the spectra from each different temperature region. Below we shall review the theoretical calculations behind the model,

and the details of the model itself. We believe this will be an effective new tool in the analysis of young SNRs or any other shocked plasmas that are dominated by elements other than hydrogen and helium.

3.3.1 Theory

The calculations begin with the ion and electron temperatures, and the equations that describe the interchange of energy between ions and electrons (similar to Spitzer (1962) equation 5-30):

$$\frac{dT_i}{dt} = -\frac{T_i - T_e}{t_{eq}} - \frac{2}{3} \frac{T_i}{\rho} \frac{d\rho}{dt} \quad (3.3)$$

$$\frac{d}{dt} \left(\frac{n_e T_e}{n_i} \right) = \frac{T_i - T_e}{t_{eq}} + \frac{2}{3} \frac{1}{\rho} \frac{n_e}{n_i} T_e \frac{d\rho}{dt} \quad (3.4)$$

These equations are valid for a gas without energy losses or electron thermal conductions with only adiabatic heating and cooling. They are not valid across very strong shocks.

The relationship between the ion and electron temperature is obvious from these equations, but X-ray data only give us a measurement of the electron temperature. We need to reduce these equations to isolate the electron temperature in terms of quantities that describe the plasma and that can be calculated with some assumptions. The first step is to add the above equations and move some factors around to arrive at:

$$\frac{n_i}{n_i T_i + n_e T_e} \frac{d}{dt} \left(\frac{n_i T_i + n_e T_e}{n_i} \right) = \frac{2}{3} \frac{1}{\rho} \frac{d\rho}{dt} \quad (3.5)$$

At this point we can perform a check for an adiabatic gas by noting that $n_i T_i + n_e T_e \propto P \propto \rho^\gamma \propto \mu^\gamma n_i^\gamma$. μ is the mean mass per particle, and γ is the adiabatic index which we assume to be 5/3. Returning to equation 3.4, we can substitute in the above equation and multiply through by $n_i/n_e T_e$,

$$\frac{n_i}{n_e T_e} \frac{T_i - T_e}{t_{eq}} = \frac{n_i}{n_i T_i + n_e T_e} \frac{d}{dt} \left(\frac{n_i T_i + n_e T_e}{n_i} \right) - \frac{n_i}{n_e T_e} \frac{d}{dt} \left(\frac{n_e}{n_i} T_e \right) \quad (3.6)$$

The form of the right hand side of the equation can be reduced noting that:

$$\frac{1}{a} \frac{da}{dt} - \frac{1}{b} \frac{db}{dt} = \frac{b}{a} \frac{d}{dt} \left(\frac{a}{b} \right) \quad (3.7)$$

Our new equation becomes:

$$\frac{n_i}{n_e T_e} \frac{T_i - T_e}{t_{eq}} = \frac{n_i T_i + n_e T_e}{n_e T_e} \frac{d}{dt} \left(\frac{n_e T_e}{n_i T_i + n_e T_e} \right) \quad (3.8)$$

Solving equation 3.8 is complicated since both electron and ion temperatures evolve over time. We need to remove them in favor of another quantity and extract the electron temperature once the equation is solved. We introduce χ_e as the percent of energy in electrons, defined as:

$$\chi_e = \frac{n_e T_e}{n_i T_i + n_e T_e} \quad (3.9)$$

so that equation 3.8 becomes:

$$\frac{n_i}{n_e T_e} \frac{T_i - T_e}{t_{eq}} = \frac{1}{\chi_e} \frac{d\chi_e}{dt} \quad (3.10)$$

At this point the ion temperature is still prominent in the equation and in the definition of χ_e . However, we must first bring in the definition of the equilibration timescale from Spitzer (1962) equation 5-31:

$$\frac{1}{t_{eq}} = g(f, A) \frac{n_e}{T_e^{3/2}} \quad (3.11)$$

where $g(f, A)$ is a function of the ionization fractions and chemical abundances that is defined as:

$$g(f, A) = \frac{8(2\pi)^{1/2} e^4 \ln \Lambda m_e^{1/2}}{3k^{3/2}} \sum \frac{Z_j^2 n_j f_j}{n_{tot} m_j} \quad (3.12)$$

The ion densities in the sum are based on the elemental abundances. The log term is the Coulomb logarithm which we assume to have a constant value, $\ln \Lambda = 30$. Spitzer (1962) describes this term in detail and shows it is not strongly dependent on changes in T_e and n_e , so a constant value assumption is a safe one to make. Substituting in the definition of equilibration timescale, and multiplying both sides of equation 3.10 by χ_e yields the following:

$$\frac{d\chi_e}{dt} = g(f, A) \frac{1}{T_e^{3/2}} \frac{n_i n_e (T_i - T_e)}{n_e T_e - n_i T_i} \quad (3.13)$$

The right hand side can be rewritten in terms of χ_e to become:

$$\frac{d\chi_e}{dt} = g(f, A) \frac{n_i}{T_e^{3/2}} \left[\frac{n_e}{n_i} (1 - \chi_e) - \chi_e \right] \quad (3.14)$$

We next eliminate the $T_e^{3/2}$ term by using our definition of χ_e which results in:

$$\frac{d\chi_e}{dt} = g(f, A) \frac{n_i n_e^{3/2}}{(n_e T_e + n_i T_i)^{3/2}} \frac{\left[\frac{n_e}{n_i} (1 - \chi_e) - \chi_e \right]}{\chi_e^{3/2}} \quad (3.15)$$

The ideal gas law can help to further eliminate the temperatures (besides those within χ_e) from the equation. The upstream ram pressure ρv_2^2 is equal to the thermal pressure created by the ions and electrons. We bring in the 1/4 drop in density from the Rankine-Hugoniot jump conditions (assuming a strong shock and $\gamma=5/3$), to determine the upstream ion density. We introduce an overall shock temperature to the ideal gas law to get:

$$\frac{3}{16k} \mu n_i v_s^2 = n_e T_e + n_i T_i = (n_e + n_i) T_s \quad (3.16)$$

or:

$$k T_s = \frac{3}{16} \frac{\mu}{\frac{n_e}{n_i} + 1} v_s^2 \quad (3.17)$$

Equation 3.16 lets us eliminate the remaining temperature terms in equation 3.15. Equation 3.17 is important for determining the expected shock temperatures for a given shock speed in a plasma with known chemical composition. Our final equation is arrived at by substituting 3.16 in 3.15 and noting that $\tau = n_e t$:

$$\frac{d\chi_e}{d\tau} = \left(\frac{16k}{3} \right)^{3/2} \frac{1}{\mu^{3/2} v_s^3} g(f, A) \left(\frac{n_e}{n_i} \right)^{1/2} \frac{\left[\frac{n_e}{n_i} (1 - \chi_e) - \chi_e \right]}{\chi_e^{3/2}} \quad (3.18)$$

This differential equation depends on the abundances of elements and their ionization fractions, as well as factors that can be calculated from these numbers such as the mean mass per particle μ , and the density ratio of electrons to ions. We bring in the shock velocity, v_s , rather than the average shock temperature, T_s , used in some current models. Solving this equation can give us the time evolution of the electron temperature in a shock-heated plasma. To simplify the above equation, we define two values assumed constant over small steps in τ , c_1 and c_2 , where $c_2 = n_e/n_i$ and c_1 is the rest of the leading factors:

$$\frac{d\chi_e}{d\tau} = c_1 \frac{[c_2(1 - \chi_e) - \chi_e]}{\chi_e^{3/2}} \quad (3.19)$$

3.3.2 The Model

The parameters for the model are the amount of electron heating at the shock front β (defined as T_e/T_i), the shock speed v_s , and the range of ionization timescale τ_{min} and τ_{max} . The code calculates emission from $\tau = 0$, but specifying a lower limit means the values below τ_{min} are not included in the spectral calculations. Also important are the abundances of the elements H, He, C, N, O, Ne, Mg, Si, S, Ca, Fe, and Ni. We have departed from the usual convention of using solar abundance ratio scales in favor of absolute abundances. A redshift parameter has also been included for completeness. Notice that no temperature parameter exists. The increase of the electron temperature in the postshock region means that a single electron temperature is not an accurate definition of the emitting material. The shock speed parameter replaces the shock temperature parameter found in Sedov models, and is more useful in understanding the overall shock structure. Integrating the model into the XSPEC data analysis package gives the code access to the data libraries for line emission and ionization fractions released with the software. We can also compare our calculations with existing models in XSPEC. Here we shall review the calculations within the model that solve the equation for χ_e above. (Please see Appendix B for a complete copy of the heavy element (HE) shock model installed in XSPEC version 10.)

The calculations begin directly behind the shock at $\tau = 0$ and step over small ranges of τ until the maximum value. We assume that at the shock front all the hydrogen and helium are fully ionized, while the heavy elements are singly ionized. From these values the initial value of the density ratio n_e/n_i is determined. β gives the initial electron temperature. If it is non-zero then we determine the initial electron temperature from the shock speed (assuming $T_i \sim T_s$) and β as:

$$T_e = \beta \frac{3}{16k} \frac{\mu}{\frac{n_e}{n_i} + 1} v_s^2 \quad (3.20)$$

We then choose a small value of $\tau_{init} = 5 \times 10^5$ s/cm³ to begin our calculations. A zero value of β indicates that no heating of electrons occurs at the shock front. We turn to an analytical approximation of equation 3.19 appropriate for small ionization timescales. If

the first step is over the values of χ_e from 0 to 0.0001 we can solve for the value of τ_{init} that we will begin our steps from:

$$\tau_{init} = \frac{2}{5} \frac{1}{c_1 c_2} \chi_e^{5/2} \quad (3.21)$$

In either case, we begin the main body of the calculations to determine the electron temperature history from the initial value of τ up to τ_{max} .

For each step in τ , we take the values of χ_e at the start of the step and use a 2nd order Runge-Kutta method with adaptive stepsizes to solve equation 3.19. The most important assumption is that ionization fractions are constant over a small step in timescale. The fractions are calculated with a modified version of the eigenvalue method described in Borkowski et al. (1994). The modification is that instead of calculating the fractions over the entire range of τ from $T_e = 0$ to the single temperature assumed in plane shock model, the fractions are calculated in each step assuming an average electron temperature over $\Delta\tau$, between τ_L and τ_H . The fractions from the previous step are the starting point for the fractions calculated in the current step. With the assumption of constant ionization fractions over a step, our density ratio n_e/n_i is also constant over that step. A fully complete solution would be able to consider the evolution of the ionization fractions over time, but that is computationally impractical. Our method estimates the time evolution of the ionization fractions as a step function. If the stepsize determined by the 2nd order Runge-Kutta stepper is larger than $10^{10} \text{ cm}^{-3}\text{s}$ in ionization timescale, this assumption may not be correct. When that occurs, the stepsize is reduced and the calculations are repeated. The question of stepsize becomes a problem for large values of ionization timescale, but testing indicated that each element reached equilibrium prior to $\tau = 5 \times 10^{13} \text{ cm}^{-3}\text{s}$. Due to instability in the calculations for such large values of τ (when the logarithmic stepsizes become too large), we cut off at this value. With equilibrium reached, further calculations are not required and we can assume the electron temperature remains constant beyond $\tau = 5 \times 10^{13} \text{ cm}^{-3}\text{s}$ and record the current value of T_e and ionization fractions for the rest of the timesteps.

The spectrum building is done in the usual fashion for XSPEC models. For each step between τ_{min} and τ_{max} , a temperature index is calculated from the log of the average T_e in the step. The index, required to be between 1 and 1251, is used to record the emission intensity of each ion of each element (159 ions total) at that temperature. The emission for

each ion at a particular temperature is the element’s abundance times the ionization fraction of that ion times $\Delta\tau$. The contributions of each ion at each temperature are then used to build the components of the X-ray spectrum. Continuum emission and line emission of the elements are determined from that information. The only departure for standard spectrum building is the question of normalization, which relates to the emission measure of the plasma. For a heavy element dominated shock n_i replaces the hydrogen density n_H in the emission measure equation so that:

$$EM = \int n_e n_i dV \quad (3.22)$$

For use in XSPEC, the parameter “norm” is defined as $EM \times 10^{-14} / 4\pi D^2$ (D is the distance to the source).

3.4 Testing the Code Versus Models

Figure 3.1 is a plot of electron temperature versus the logarithm of ionization time-scale for three different plasmas: solar abundance, pure oxygen, and pure iron. The solar abundance plasma reaches the lowest maximum value of electron temperature. The curves for pure oxygen and pure iron reach successively higher temperature at their equilibrium plateaus due to the increase in mean mass per particle for each plasma. Although these curves show the different histories, Figure 3.1 does not help in understanding observed X-ray spectra. Observers must start by observing the T_e values and then have to work backwards to determine abundances. If the pure O plasma is scaled to the same maximum T_e as the solar plasma, then we get Figure 3.2, the observer’s perspective. Figure 3.3 is the same for the Fe plasma. We can clearly see the composition of the emitting material greatly affects the shape of the electron temperature history. This indicates that the assumption of solar abundances will not give the correct spectrum when the composition is far from solar. Examining plots of electron temperature versus ionization timescale can also be used to examine the effects of electron heating at the shock front. Figure 3.4 shows the electron temperature history for several values of β (0.0, 0.1, 0.5, and 1.0). This example used a pure oxygen plasma, with a shock speed of 1000 km s^{-1} and $\tau_{max} = 1 \times 10^{13} \text{ cm}^{-3}\text{s}$. For large β , the high temperatures at the shock decrease rapidly with successive ionizations that eventually bring the curve back to the $\beta = 0$ case, but with significant differences in

the histories.

As discussed above, single-component NEI models assume a single temperature and single ionization timescale for the emitting plasma. The HE model should be able to reproduce the behavior of NEI models for large τ . With an ionization timescale of $10^{14} \text{ cm}^{-3}\text{s}$, the plasma is in equilibrium. We can calculate the shock temperature from a shock speed of 1000 km s^{-1} using equation 3.17. This shock temperature is equal to the electron temperature of the NEI model in equilibrium. Fake datasets were generated with the *vnei* model in XSPEC with *Chandra* ACIS-S response for this timescale for pure plasmas of each heavy element and for a solar abundance plasma. We then applied the HE shock model to these generated datasets and compared the results. Figures 3.5, 3.6, and 3.7 show the results of this comparison for three example plasmas: pure oxygen, pure iron, and a solar abundance plasma respectively. These fits are rather poor, but the reason is very descriptive of the differences in the models. The HE model calculates from $\tau = 0 \rightarrow 10^{14} \text{ cm}^{-3}\text{s}$, and the contributions from low ionization timescales are present in these figures. With multiple ionization timescales these model cannot agree. Agreement comes by setting the values of $\tau_{min} = 1 \times 10^{13} \text{ cm}^{-3}\text{s}$ to isolate the equilibrium plateau region of the T_e versus τ plots. With τ_{min} set we get results seen in Figures 3.8, 3.9, and 3.10. The solar comparison sees residuals reduced by a factor of three, and for the plasmas of pure heavy elements the HE model matches the *vnei* models, except for the normalization. Since *vnei* uses solar abundances and HE does not, the difference in normalization is very close to solar values it was expected to reproduce. Table 3.1 summarizes fit results for each element including the fit statistic of the fits (χ^2/DOF)(where a value of ~ 1.0 indicates an acceptable fit), the normalization value, and the solar abundances for comparison. Recall equation 3.22 that describes the expected difference in normalization.

Next we compare the HE model to the plane-parallel shock model *vpshock* for a nonequilibrium plasma with $\tau = 10^{11} \text{ cm}^{-3}\text{s}$. The single electron temperature to be used in *vpshock* for this comparison is the average T_e generated from the HE model T_e versus τ data. These values were calculated for a shock speed of 1000 km s^{-1} in several plasmas. We examine two new example plasmas with abundances of heavy elements consistent with supernova explosion ejecta from Iwamoto et al. (1999). Table 3.2 gives the heavy element abundances calculated from Iwamoto’s mass calculation of each ion for a Type II remnant and Type Ia deflagration model (model W7 in Iwamoto et al. (1999)). The discrepancies at high and low energies seen in Figures 3.11 and 3.12 show the limitations of the plane-parallel

shock models currently in use. A single electron temperature cannot accurately define the X-ray spectrum of a plasma with non-equilibrium ionization. With the HE model we can examine evolution of supernova remnant ejecta emission over a full range of ionization timescales and electron temperatures. We generated spectra for a type II explosion at 4 values of τ , from $\tau = 10^{10} - 10^{13} \text{ cm}^{-3}\text{s}$. Different levels of normalization were chosen to separate the spectra in Figure 3.13.

3.5 Discussion

The electron temperature evolution in a shocked plasma is very dependent on the chemical composition of the emitting material. We have derived a relationship between electron temperature and ionization timescale, equation 3.18, that can be solved with some basic assumptions. With only Coulomb interactions occurring behind the shock, energy is transferred from ions to electrons. These hot electrons are responsible for thermal X-ray continuum emission, and collisional excitation of heavy ions by these electrons is responsible for X-ray line emission. The electrons freed from the ions by collisional ionization by hot electrons are added to the pool of electrons in the plasma. The spectra calculated from the new heavy element shock model are consistent with simple NEI models in the large τ limit of ionization equilibrium. Before equilibrium is reached, however, significant differences can arise. When the abundances are far from solar, as is the case with our Type Ia and Type II ejecta plasmas, the HE model provides improvement over the tools currently available.

Let us review the assumptions made within our calculations. We first need to assume that the populations of the different ionization states for each element are constant over small steps in ionization timescale. Although ion density is constant and depends on the elemental abundances, the electron density varies with τ due to the ionizations occurring behind the shock. The new calculations presented here estimate these values as step functions that remain constant over a range of $\Delta\tau$. We can make our calculations as accurate as necessary by decreasing $\Delta\tau$, but at the expense of efficiency. We also must assume that elements are singly ionized in the preshock region to create an initial pool of electrons even if H and He are not present. The X-ray spectrum is insensitive to the specific ionization state in the preshock material (Borkowski et al., 2001, who choose initially neutral states based on this assumption).

Another important issue is the uniformity of the shock speed v_s . Shocks slow down over time as they run into material (either ejecta or ISM), with different rates depending on the density. The limitation of simple NEI models is the single ionization timescale, and that several NEI components need to be used to describe the plasma. If multiple shock speeds are required to describe an object's emission, then several HE shock models may need to be used. The HE shock model makes no attempt to investigate the distribution and density of elements within the emitting material. It assumes all the elements are completely mixed and evenly distributed in the SNR. This is a significant inaccuracy since new data have shown dense clumps of O and Fe in many young SNRs. These limitations are the same as for any plane parallel shock model. While full Sedov models can describe the expansion of a SNR, plane shock models are very useful with the current generation of X-ray data. The ability to perform spatially resolved spectroscopy makes the estimate of a plane shock safe over a small region. The HE shock model brings in detailed calculations of the ionization states and electron temperature for plasmas far from solar abundance ratios, and should be an important new tool in the analysis of X-ray spectra.

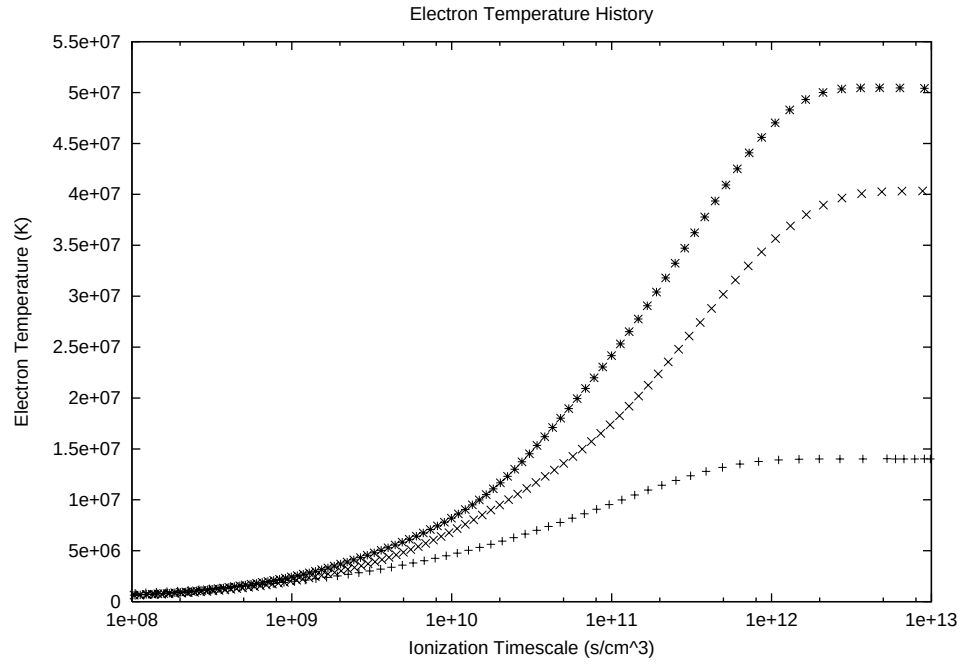


Figure 3.1: Electron temperature history of solar abundance (crosses), pure oxygen (x's) and pure iron (asterisks) plasmas for shock with velocity 1000 km s^{-1} .

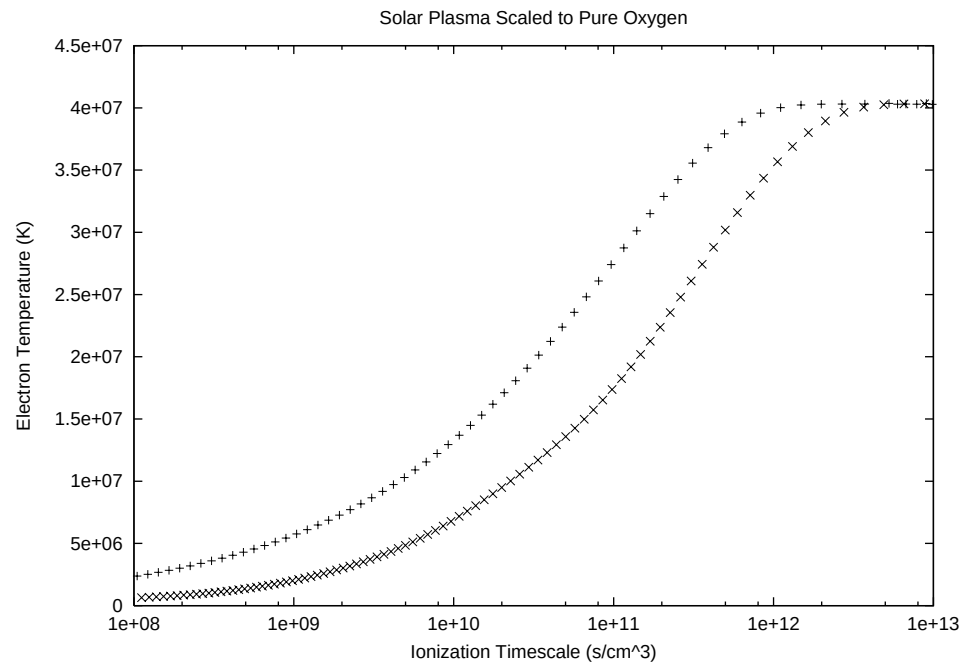


Figure 3.2: Electron temperature history of solar abundance (crosses), pure oxygen (x's) plasmas scaled to maximum solar T_e .

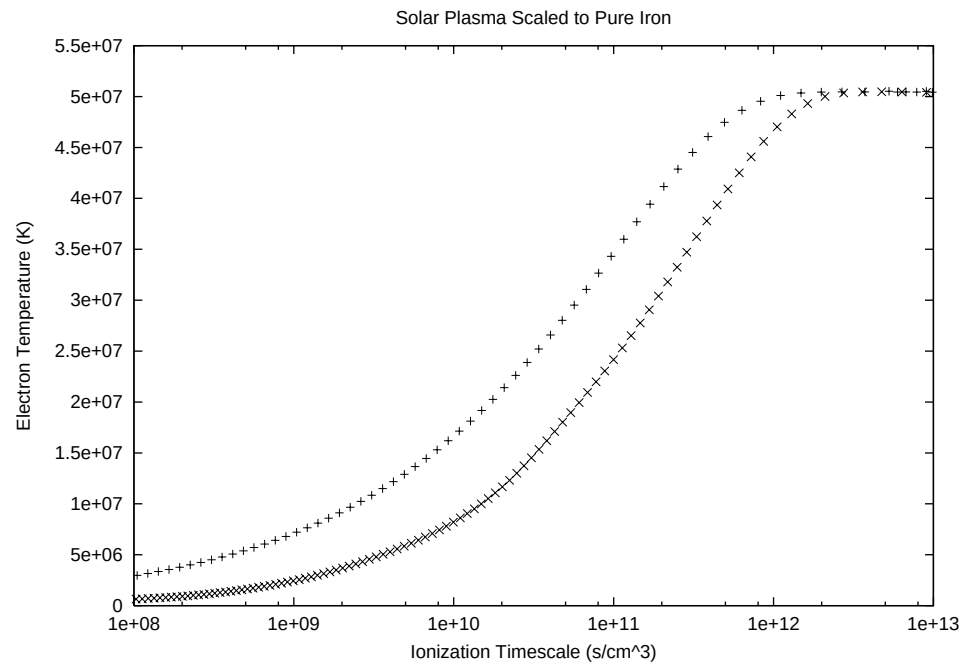


Figure 3.3: Electron temperature history of solar abundance (crosses), and pure iron (x's) plasmas scaled to maximum solar T_e .

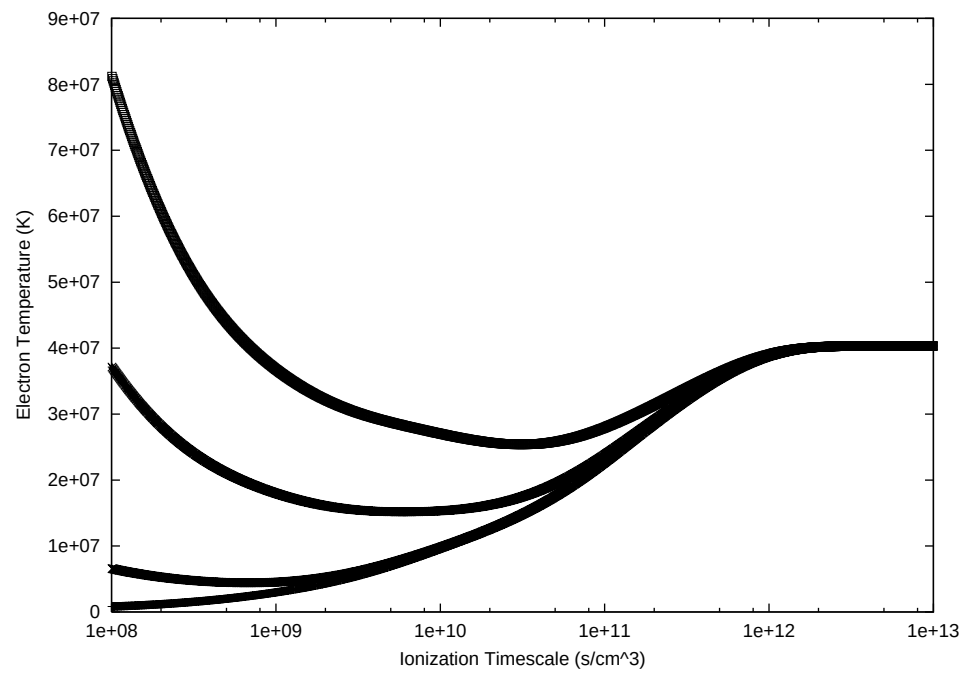


Figure 3.4: T_e versus τ plot for values of $\beta = 0.0, 0.1, 0.5$, and 1.0 (lowest to highest curves).

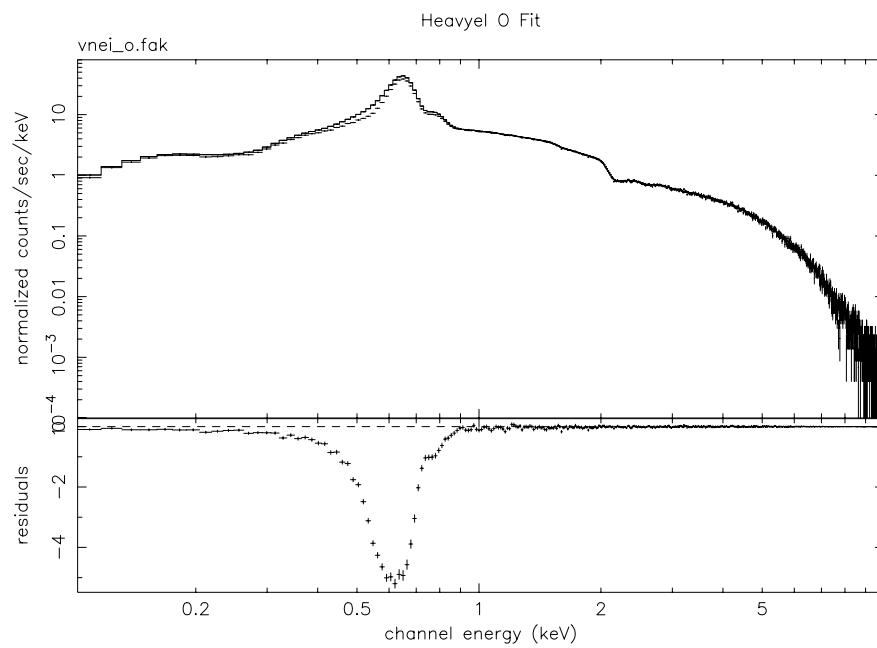


Figure 3.5: HE shock versus v_{nei} for a pure oxygen plasma.

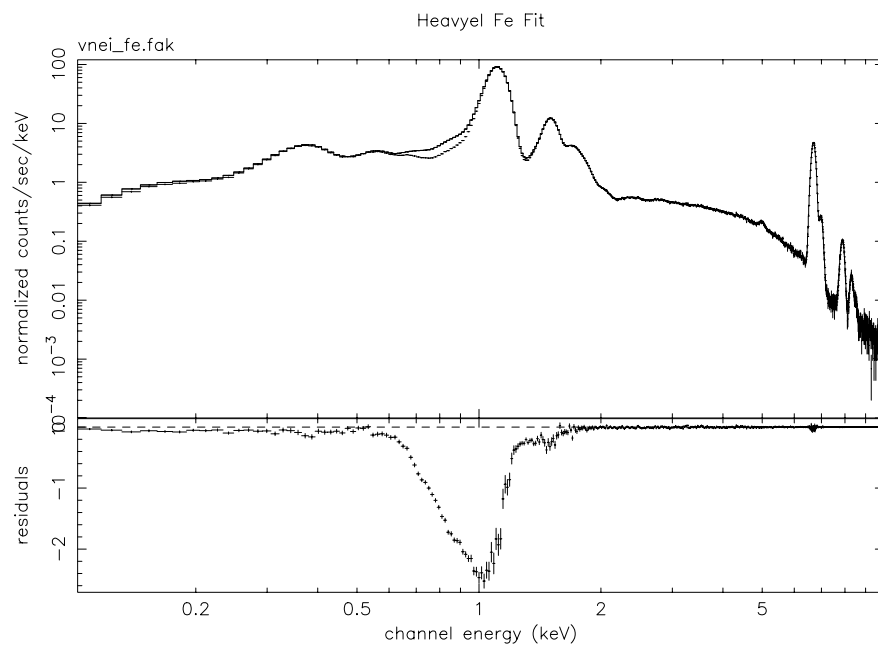


Figure 3.6: HE shock versus $vnei$ for a pure iron plasma.

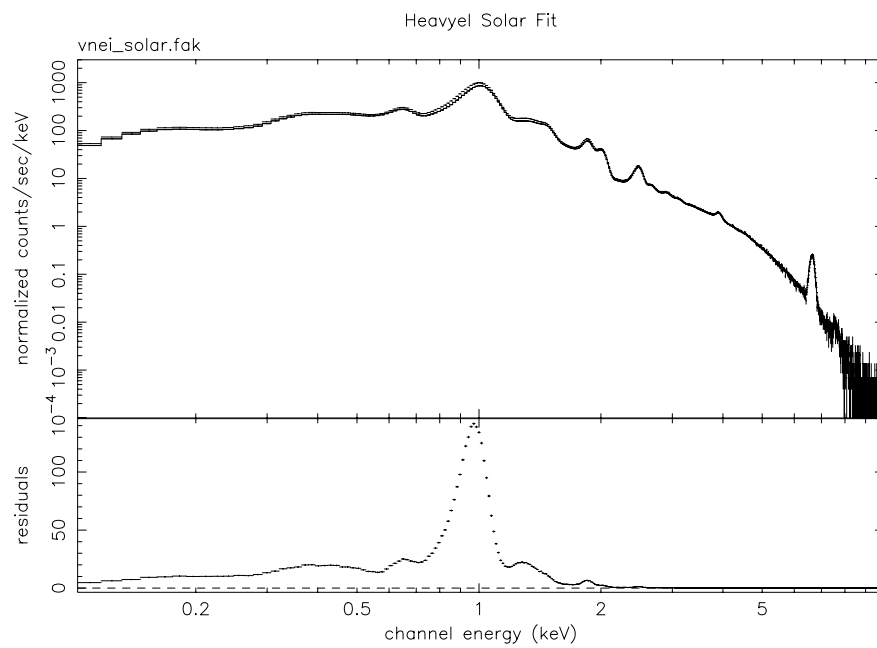


Figure 3.7: HE shock versus $vnei$ for a solar abundance plasma.

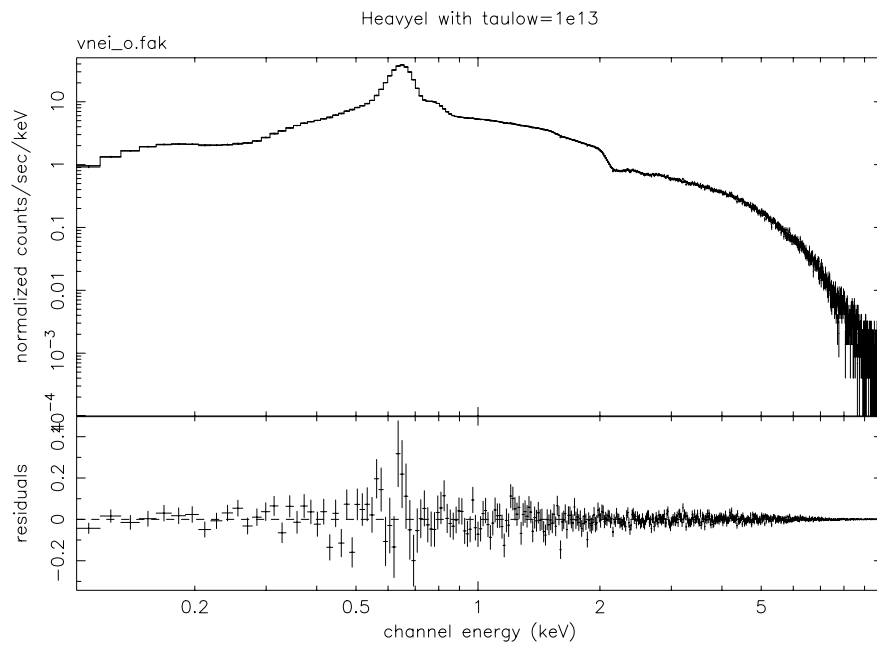


Figure 3.8: HE shock versus $vnei$ for a pure oxygen plasma, with $\tau_{min} = 1 \times 10^{13} \text{ cm}^{-3}\text{s}$.

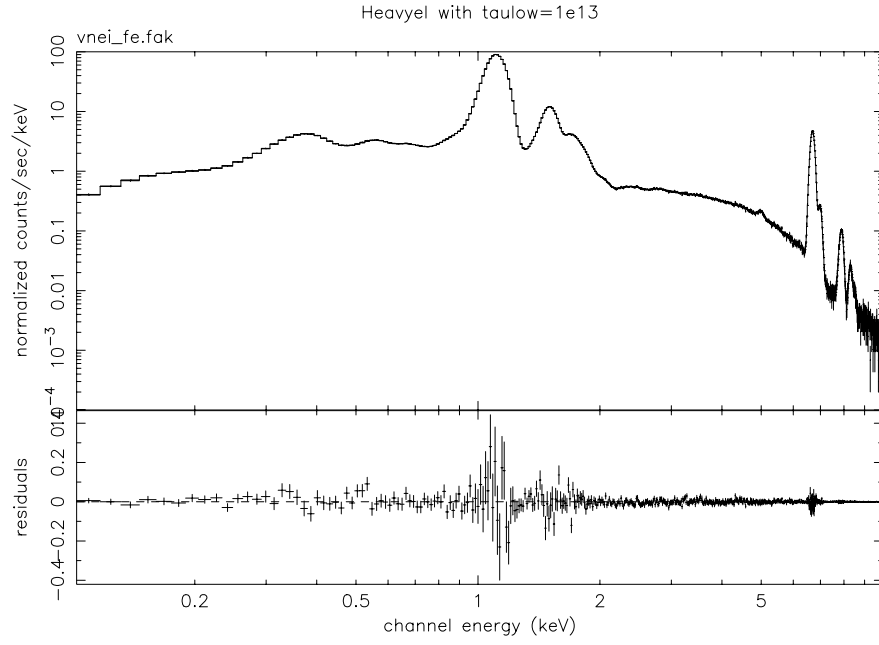


Figure 3.9: HE shock versus $vnei$ for a pure iron plasma, with $\tau_{min} = 1 \times 10^{13} \text{ cm}^{-3}\text{s}$.

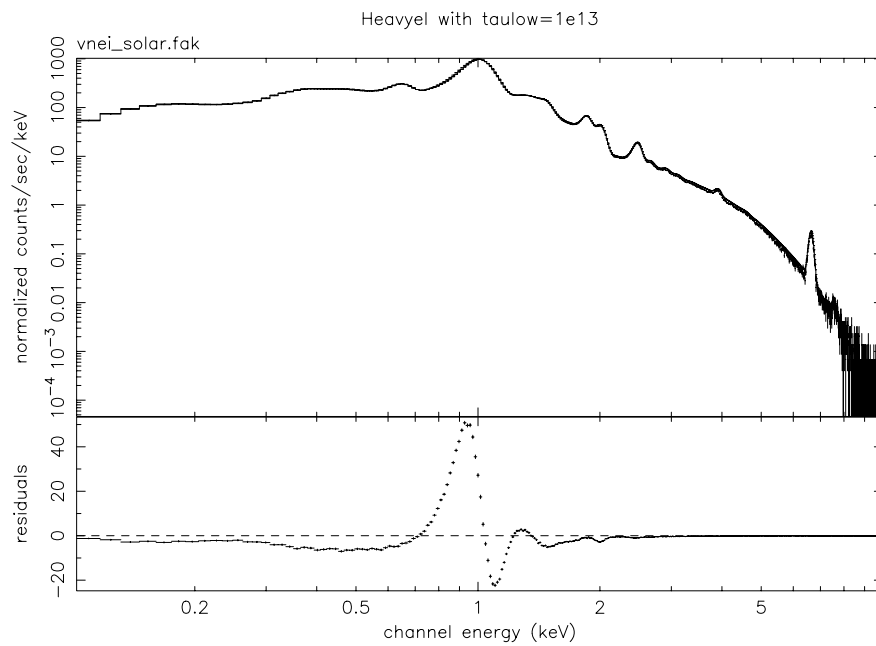


Figure 3.10: HE shock versus $vnei$ for a solar abundance plasma, with $\tau_{min} = 1 \times 10^{13} \text{ cm}^{-3}\text{s}$.

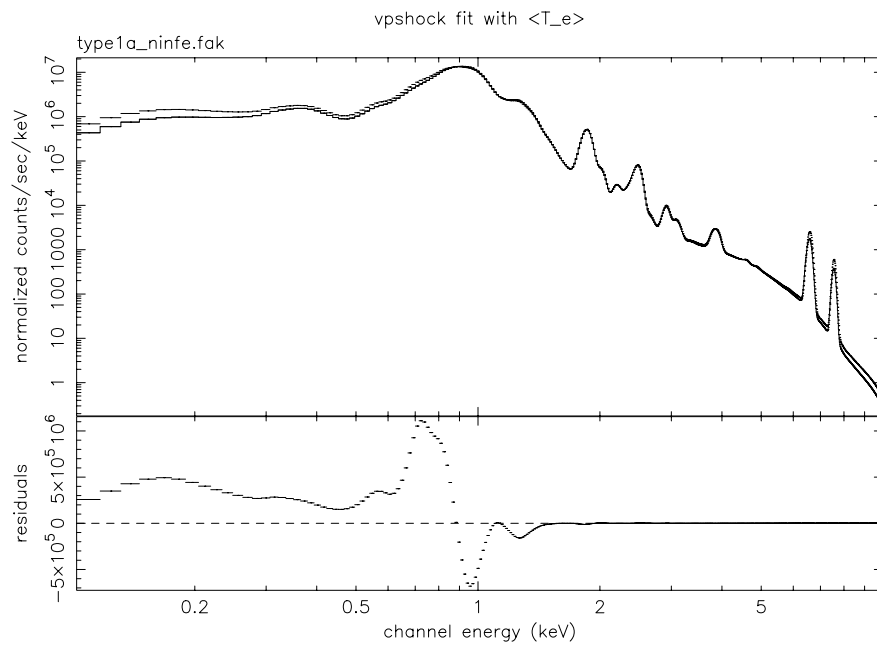


Figure 3.11: HE shock versus *vpshock* for a Type Ia abundance plasma.

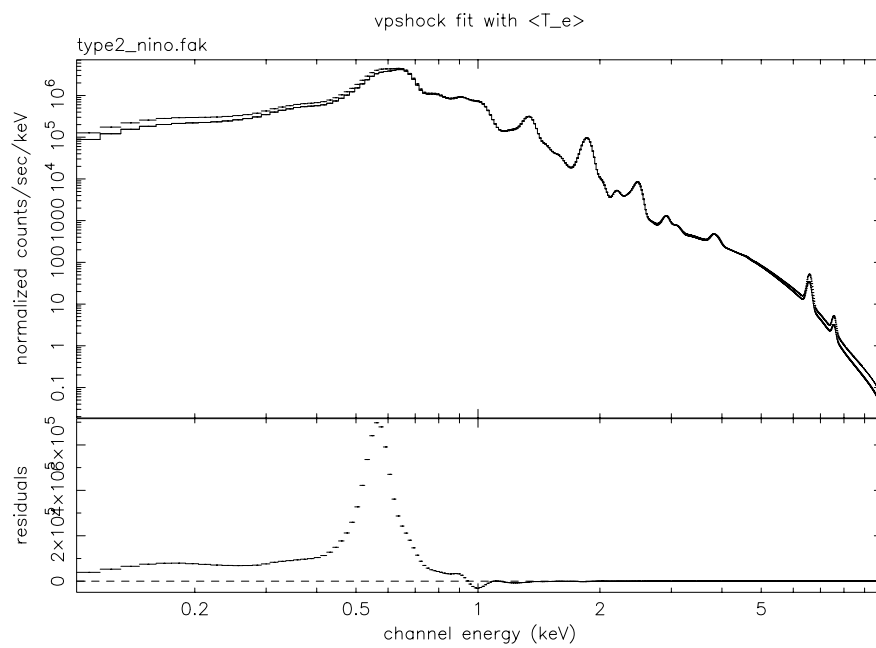


Figure 3.12: HE shock versus $vnei$ for a Type II abundance plasma.

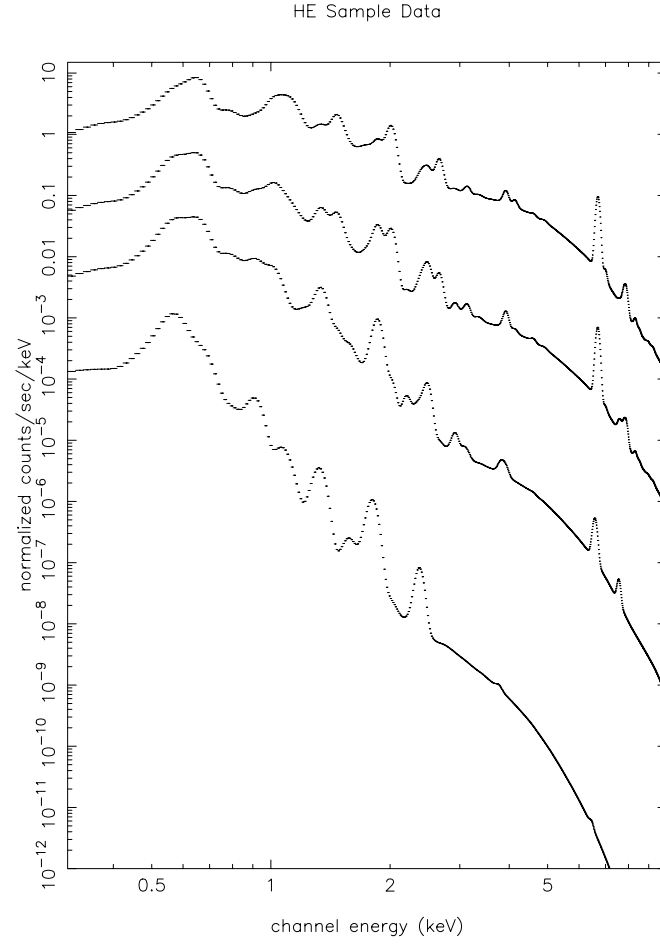


Figure 3.13: Evolution of HE shock model predicted spectra for 4 levels of ionization time-scale. Lower curve is for $\tau = 10^{10} \text{ cm}^{-3}\text{s}$, with successive orders of magnitude up to $10^{13} \text{ cm}^{-3}\text{s}$.

Table 3.1: NEI fits to HE shock model

Element	χ^2/DOF	norm	solar
C	0.84	3.615×10^{-4}	3.631×10^{-4}
N	0.82	1.111×10^{-4}	1.122×10^{-4}
O	0.99	8.437×10^{-4}	8.511×10^{-4}
Ne	0.97	1.218×10^{-4}	1.23×10^{-4}
Mg	0.98	3.737×10^{-5}	3.8×10^{-5}
Si	1.0	3.50×10^{-5}	3.55×10^{-5}
S	0.97	1.60×10^{-5}	1.62×10^{-5}
Ca	0.96	2.26×10^{-6}	2.3×10^{-6}
Fe	0.90	4.65×10^{-5}	4.68×10^{-5}
Ni	1.0	1.76×10^{-6}	1.8×10^{-6}

Table 3.2: SN ejecta abundances from Iwamoto et al. (1999) normalized to Fe

Element	Type Ia	Type II
C	0.344	4.232
N	7.09×10^{-6}	0.071
O	0.765	72.28
Ne	0.019	7.351
Mg	0.030	3.239
Si	0.477	2.785
S	0.232	0.824
Ca	0.025	0.094
Fe	1.0	1.0
Ni	0.183	0.065

Chapter 4

Maximum Energies of Shock-Accelerated Electrons in Large Magellanic Supernova Remnants

4.1 Introduction

Observations of the ion component of galactic cosmic rays show a straight power-law spectrum from a few GeV up to the “knee” around 1000 TeV (see references in Blandford & Eichler (1987) for these and other observational generalities about cosmic rays). Shock acceleration of particles by supernova remnants is believed to create cosmic rays up to this energy. However, from radio synchrotron observations, we learn only about electron energies in the range 0.1 to 10 GeV. Strong shocks should accelerate both electrons and ions, and should treat them equally above a few GeV, when all particles become relativistic. Between 5 GeV and 200 GeV, cosmic-ray ions and electrons have different power-law distributions

when viewed from Earth. However, these distributions are consistent with both ions and electrons having the same slope at the cosmic ray source, but evolving differently in the course of propagating through the Galaxy. Understanding the cosmic-ray electrons should give us some information on the cosmic-ray ions, which contain most of the energy in the cosmic-ray spectrum.

The power-law distribution of cosmic-ray electrons accelerated by a SNR shock can be cut off at high energies by several factors. Synchrotron and inverse-Compton radiative losses can cause the spectra to cut off. Finite shock age is also a factor since the time required for a particle to be accelerated to a given energy by diffusive propagation back and forth across the shock is a sharply increasing function of particle energy. Finally, there can be a change in scattering due to the lack of long wavelength MHD waves ahead of the shock. The first cutoff cause affects only the electron component, but finite age and change in scattering properties should affect ions and electrons equally.

X-ray observations have shown several SNRs with spectral evidence for nonthermal emission (SN 1006, Koyama et al. (1995); G347.3-0.5, Koyama et al. (1997); Slane et al. (1999); Cas A, The et al. (1996)). This nonthermal emission could be explained by extrapolating the radio synchrotron emission across the entire electromagnetic spectrum to X-rays. Comparing the X-ray flux to the extrapolated radio flux, we see that the electron spectrum must have steepened between the radio and X-ray regimes. We can bound from above the frequency at which this steepening begins using a synchrotron cutoff model. The cutoff frequency is related to the maximum energy to which the shock can accelerate electrons, which can therefore also be bounded from above by the observations. We must still consider whether these electron limits also apply to ion acceleration.

A previous study by Reynolds & Keohane (1999), hereafter RK, used a sample of 14 Galactic SNRs to constrain the maximum energies of electron acceleration. In this paper, we have undertaken a parallel investigation, studying 11 SNRs in the LMC, using data from the ASCA Public Archive. Although there is little evidence concerning the cosmic-ray spectrum in the LMC, the SNRs in the LMC appear to be similar in most respects to those in the Galaxy. Since we are attempting to define the electron acceleration properties of SNRs in general, the LMC provides an excellent sample. It has many SNRs at many different stages of development, all at about the same distance. This sample has young remnants where E_{max} may still be rising due to increasing age, and mature remnants where E_{max} may be dropping due to radiation losses competing with slowing rates of particle

acceleration as the shock decelerates.

In Section 2, we will review the sample of LMC remnants, and discuss the data and the spectral models used in this study. Section 3 will summarize our results. In Section 4, we will discuss the implications of the cutoffs we infer in the electron spectrum, and what can be learned about the ion spectrum. Finally, Section 5 will list our conclusions.

4.2 Analysis Technique

4.2.1 The Sample

Radio and X-ray properties of the SNRs appear in Table 1. The radio parameters quoted here are from several sources. The 1 GHz flux densities in column 2 were extrapolated from the flux at the observed frequency from 408 MHz (Mathewson et al., 1983) and 4.75 GHz (Filipovic et al., 1998), using the spectral indices, α , in column 3 from these same references, except as noted in the table (we define $S_\nu \propto \nu^{-\alpha}$). SNR DEM L71 (0505-67.9) is a very dim radio source, and flux densities are available only at 843 MHz (Mills et al., 1984), so a spectral index of 0.5 was assumed. X-ray size and ROSAT PSPC count rate in columns 4 and 5 come from Williams et al. (1999). The final two columns give the extrapolation of the synchrotron power law into the X-ray regime, assuming no change in slope. This 1 keV flux density is listed both in mJy and in photons $\text{s}^{-1} \text{cm}^{-2} \text{keV}^{-1}$ for the convenience of radio and X-ray observers.

The supernova remnants in this study fall into two main categories: mature remnants, most likely in the Sedov phase, and younger remnants from type Ia SNe, most of which exhibit Balmer lines in their optical spectrum. Remnants that appear to be in the Sedov phase include N23 (0506-68.0), N63A (0535-66.0), N132D (0525-69.6), and N49B (0525-66.0) (Hughes et al. (1998), hereafter HHK). DEM L71 seems to be in a transition state from ejecta-dominated to dominated by swept-up ISM. 0534-69.9 has not been well studied, but its size and morphology (diffuse emission across the face of the shell; Williams et al. (1999)) suggests Sedov phase. Remnants suspected to have type Ia progenitors include 0453-68.5, N103B (0509-68.7), DEM L71, 0509-67.5, 0519-69.0, and 0548-70.4. The last four objects on this list also have an optical spectrum dominated by Balmer line emission (Smith et al., 1994). N103B is identified by characteristic lines in the X-ray spectra as type

Ia (Hughes et al., 1995), and 0453–68.5 is classified by Chu & Kennicutt (1988) as type Ia. The youngest remnants in this survey are 0509–67.5, N103B, and 0519–69.0 (Hughes et al., 1995). Optical spectroscopy gives shock velocities implying an age of 750–1500 years for 0519–69.0, and an upper limit of 1000 years for 0509–67.5 (Smith et al., 1994). N132D has high-velocity oxygen-rich filaments within the shell (HHK), and both N103B (Hughes et al., 1995) and N63A (HHK) contain optically bright knots.

4.2.2 The X-ray Data

The data for the 11 LMC SNRs were retrieved from the ASCA data archive on the HEASARC web page (heasarc.gsfc.nasa.gov). The unscreened events files for each observation were downloaded and processed using the FTOOLS software package. ASCA contains four X-ray telescopes, each with a detector. Two detectors are solid-state imaging spectrometers (SIS0 and SIS1), with each containing an array of four CCD camera chips. The other two detectors are gas imaging spectrometers (GIS2 and GIS3), which are gas scintillation proportional counters. More details on ASCA and its components can be found in Tanaka et al. (1994).

Table 2 summarizes the observations, and the exposure times for both the SIS and GIS detectors. SIS0 and SIS1 data in 1-CCD mode were used. Data in faint datamode were converted to bright mode and corrected for dark frame error. These files were added to data in bright datamode to increase photon counts. The combined data were screened using the standard Revision 2 procedure. High bit rate data were used for each SIS observation. GIS2 and GIS3 data were collected in PH datamode, also with high bit rate. Standard Revision 2 screening criteria again were used. Details on Revision 2 procedure can be found at the HEASARC web site.

Extracting the spectra involved drawing circular regions about the image of each remnant in SAOimage, making sure that each region included all the radiation from the source. To get the background spectra, all obvious sources were eliminated from the image before extraction. For the SIS data in 1-CCD mode, the chip has a FOV of $10' \times 10'$. The largest SNR in our survey is only slightly larger than $2' \times 2'$, so there was sufficient space on the chip to extract an uncorrupted background spectrum. For GIS data, the field of view is a circle $25'$ in radius, so there was less of a problem getting a source-free background spectrum.

Next, the appropriate response matrices were created using the FTOOLS software package. For SIS data, both the response matrix file and the ancillary response file had to be created for each observation. The response matrix files are the same for all GIS2 and GIS3 observations, and those files come with the data sets. Ancillary response files still needed to be created, using the FTOOLS package. All of the data were grouped with a minimum of 25 counts per bin.

The four spectra for each object were analyzed simultaneously in XSPEC, with relative normalizations of the fits to the spectra allowed to vary. This allowed us to freeze one data set and shift the others along the vertical axis to account for small differences in effective area. Bad channels were ignored, as were channels above 10.0 keV. For the SIS spectra, the range below 0.8 keV was ignored, and below 0.4 keV for the GIS spectra. A full study of each remnant is in preparation; this paper will concentrate on constraining the maximum electron energies. Each SNR was fit with a Sedov model and the synchrotron cutoff model (called *SRCUT* in the XSPEC spectral-analysis software), as described below.

4.2.3 The Models

The details of the Sedov model used in this paper can be found in Borkowski et al. (2001). Calculations of the X-ray spectra include reliable atomic data for Fe L-shell lines, and are based on the updated Hamilton-Sarazin spectral code (Hamilton et al., 1983). The code allows the electron temperature to differ from the (mean) shock temperature, to allow for the possibility of incomplete temperature equilibration between electrons and ions. In order to speed up the fitting procedure, we used a pre-calculated table of Sedov model solutions, with the fitting program interpolating between grid points. The Sedov grid of models was created by the FTOOL *kbmodels* (K. Arnaud, private communication). The grid covers a range of shock temperatures from 0.1 to 3.0 keV, the same range in electron temperature, and a range of ionization timescales of $\tau \equiv n_e t = 10^9$ to 10^{13} s cm $^{-3}$. For each parameter kT_s , kT_e , and τ , ten grid points were created. The models allow arbitrarily variable abundances, but since the Sedov model presumes emission dominated by the blast wave presumably encountering normal-abundance material, we froze abundances to 0.4 solar as appropriate for the LMC.

Any nonthermal emission is modeled with the *SRCUT* model. This model assumes a spatially uniform electron spectrum, $N(E) = KE^{-s} \exp(-E/E_{\max})$, radiating in a uni-

form magnetic field – that is, the synchrotron volume emissivity is just the convolution of $N(E)$ with the single-electron emissivity. The spectral shape is a power-law with a cutoff dropping roughly as $\exp(-(\nu/\nu_{\text{rolloff}})^{+1/2})$ where $\nu_{\text{rolloff}} = 1.82 \times 10^{18} E_{\text{max}}^2 B$, all quantities in cgs units. For more details on this model, refer to Reynolds (1998). Since the model assumes continuity between the radio and X-ray emission, this rolloff frequency, the radio spectral index, and the flux at 1 GHz are the only parameters, the flux normalization being set by the observed radio flux. Values for the spectral index and flux at 1 GHz for our target objects have been reported in the literature (see references in Table 1), and were frozen during fitting. Thus only a single adjustable parameter describes synchrotron emission for each remnant.

4.2.4 The Spectral Fitting Methods

The spectra of the 11 LMC SNRs were fit using standard Wisconsin absorption (Morrison & McCammon, 1983) applied to the Sedov model and the synchrotron cutoff model. Abundances were frozen at 0.4 solar. The radio parameters from Table 1 were frozen as well. The parameters that remained to be fitted were absorbing column density N_H , electron temperature kT_e , shock temperature kT , ionization timescale τ , and cutoff frequency ν_{rolloff} . In some cases the column density was assumed to be $1.2 \times 10^{21} \text{ cm}^{-2}$ to the LMC.

Our purpose is to bound the maximum energy to which a SNR can accelerate electrons, from X-ray observations. We adopt the strategy of allowing the fitting program to employ a synchrotron component in addition to our best idea of the appropriate thermal model. The fitting process will then produce values for the one adjustable synchrotron parameter, ν_{rolloff} . If we had total confidence in the appropriateness of these thermal models, and if the fits were of high statistical quality, we could accept the results as indicating the actual presence of a synchrotron component. However, since neither of these is the case, we shall conservatively interpret the values of ν_{rolloff} as indicating the *maximum* possible synchrotron component allowable by the data. To be even more conservative, we shall use as our upper limit not the best-fit value of ν_{rolloff} , but the 3σ upper bound to that fitted value, as determined by XSPEC. Remarkably, we obtain powerful constraints even with this highly conservative approach.

RK used a different method of data analysis. That study of Galactic remnants used

only GIS data from the ASCA satellite, and the study limited itself to Galactic remnants that fit within the GIS field of view. Their spectral modeling used standard Wisconsin absorption and allowed the *SRCUT* model to account for the entire continuum. For some remnants it was necessary to add a bremsstrahlung component, and Gaussian lines were also added when needed. However, they made the same interpretation of the fitted values of ν_{rolloff} as upper limits.

To test the dependence of our results on the analysis method, we refit all the remnants using methods similar to RK. For all of the remnants in this LMC survey a bremsstrahlung component was needed. Only GIS data were used in this case, for purposes of comparison to RK. The results of this comparison can be found in Column 4 on Table 3. We tested the RK method on the LMC remnants, but we also tested our methods on a Galactic remnant (3C 391) from the RK survey to compare with their results.

4.3 Results

The results of this study can be found in Table 3. Column 2 lists the reduced χ^2_ν from our fits with the Sedov model and the synchrotron cutoff model. Column 3 gives the best fit value for ν_{rolloff} , and Column 4 gives the 3σ upper limit on ν_{rolloff} determined from the model fits, which we shall use to obtain limits on E_{max} as described above. Parallel results using the RK method fits on the LMC remnants (that is, 3σ upper bounds on the fitted values) are found in column 5. The final two columns are the calculated E_{max} upper limits in units of ergs (column 6) and TeV (column 7).

While the compilations of radio data did not quote errors for flux densities, the range of errors quoted for spectral index varied from ± 0.03 for N132D to ± 0.19 for N49B. Typical errors in radio parameters are 10% in flux density and ± 0.1 in spectral index α . This error in α dwarfs all other sources of uncertainty in radio data and in relative calibration of radio and X-ray flux scales, since a change in α of 0.1 results in a factor of 7 variation in extrapolating a spectrum from 1 GHz to 1 keV ($= 2.41 \times 10^{17}$ Hz). All other calibration errors are considerably smaller than a factor of 2. In order to test the dependence of our results on the radio parameters, we refit the data varying α by 0.1 and $S_{1\text{GHz}}$ by 10%. Tables 4 and 5 show the results of this analysis for DEM L71 and N63A, respectively. The boldface values in these tables are the fitted results. For all the remnants, even in the worst-case

conspiracy of extreme radio uncertainties, ν_{rolloff} rises by less than a factor of 6.3 implying an increase of only $6.3^{1/2} = 2.5$ in E_{max} . The extreme insensitivity of our derived limits on E_{max} to all these sources of error and uncertainty is a major reason for our confidence in these results.

The error analysis for N63A found in Table 5 behaves as one would expect. When the spectral index increases by 0.1, the rolloff frequency increases (the inverse is also true). The effect of the 10% change on the flux shows an inverse relationship with the frequency. This is the expected result, if only the radio parameters were changed and the thermal parameters were frozen to their values from the original fit. As mentioned above, we do not have total confidence in the thermal models. For this reason, the thermal parameters were allowed to vary as we refit the remnants to create Tables 4 and 5. For N63A the thermal parameters remained relatively the same for each fit in Table 5, but that was not the case for DEM L71 in Table 4. The thermal parameters varied such that the rolloff frequency for each fit did not change for the lower flux values. For the higher flux values, we see the expected variation in the rolloff frequency.

Plots of the spectral fits for SNRs N63A and N132D are displayed as folded spectra in counts $\text{s}^{-1} \text{keV}^{-1}$ in Figures 1 and 2, respectively. Both the Sedov blast wave model and the synchrotron cutoff model component are shown on the plots, with the synchrotron component the lower curve at low energies.

From Table 3, we can see that no supernova remnant in this survey is producing an unbroken power-law distribution of electrons beyond 100 TeV. This is the major result of our paper. Table 3 also includes the RK method results which show that the Sedov model with the *SRCUT* component gives much lower cutoff frequencies than the values from the RK method. There are some changes in rank between the methods seen in the middle five remnants but the top four objects and the bottom two remained the same. The method of RK does seem to give us a higher (more conservative) upper limit on the cutoff frequency. The same result was seen in the analysis of 3C391, where the break frequency was slightly lower using our methods (0.049 keV) than the RK result (0.058 keV).

4.4 Discussion

Observations of the spectrum of cosmic-ray electrons indicate that the spectrum has already begun to fall below the cosmic-ray ion spectrum at 1 TeV (Blandford & Eichler, 1987, see references in). It is cosmic-ray ions that extend out to the “knee” at 1000 TeV; we have no direct evidence of cosmic ray electrons near that energy. In this regard, our results that SNRs cannot accelerate cosmic ray electrons beyond 100 TeV are consistent with observations.

However, there is cause for concern. Cosmic-ray observations indicate that ions and electrons have similar power-law spectra at the cosmic-ray sources, while theory reassures us that strong shocks should treat electrons and ions equally at energies well above 1 GeV. The one obvious explanation for differences between electron and ion spectra is radiative losses on the electrons. If all our remnants have electron spectra limited by such losses, the ion spectra could extend unbroken to much higher energies. However, the finite age of the shock, and a change in scattering, will affect both ions and electrons. A cutoff of the electron spectrum below 100 TeV would then cast doubt on the ability of supernova remnants to accelerate cosmic ray protons up to the “knee” at 1000 TeV.

The Galactic remnant study (RK) was limited to relatively small-diameter, probably young SNRs, since the observations had to fit well inside the ASCA GIS field of view. But the oldest, largest remnants in the LMC are small enough in angular diameter to fit easily on a single SIS chip. Young remnants have had only a short time to accelerate particles, but since their shocks are strong and accelerate particles rapidly, they have a high limit for radiative losses. Older remnants, while having more time to accelerate particles, have much lower limits for radiative losses (Reynolds, 1998). Such SNRs could become loss limited after a few thousand years. Therefore, older remnants could have an ion spectrum that extends beyond the maximum energies we have reported.

In the previous study of Galactic remnants, the synchrotron cutoff model was allowed to represent the entire continuum (with gaussian lines and a bremsstrahlung component, when necessary). We can see that the E_{max} values decrease when the more realistic Sedov model is used. This result is expected due to the thermal nature of most SNR X-ray spectra. By allowing the Sedov model to account for most of the continuum in each object, we have attempted to isolate any nonthermal component to the spectrum. The synchrotron cutoff model then fits the part of the spectrum not covered by our single-component

Sedov model. In the younger remnants, emission from ejecta is probably also significant; this would require an additional thermal component. More detailed thermal models should continue to reduce the upper limits for electron acceleration found in this paper.

4.5 Conclusion

This survey of LMC supernova remnants has found that their blast waves’ ability to accelerate electrons can be bounded from the X-ray spectra. By determining the point at which the electron spectrum begins to steepen, we have found that 80 TeV is the upper limit for the maximum energy to which these LMC remnants can accelerate electrons in an unbroken power law from radio-emitting energies. With more detailed models and better X-ray data, the conservative upper limits on E_{max} found in this study will become lower. The question of supernova-remnant shock waves being able to accelerate particles up to the “knee” at 1000 TeV remains uncertain. The ion component could continue beyond the electron component, but finite shock age and change in scattering properties should affect both equally. It is unknown if there is a similar “knee” in the LMC cosmic ray spectrum, but the properties of LMC supernova remnants should be consistent with remnants in the Galaxy. The fact that both the youngest and the oldest remnants in the LMC fail to accelerate electrons to within even an order of magnitude of the “knee” suggests that some revision may be needed in the picture of galactic cosmic ray origin.

We are pleased to acknowledge helpful discussions with K. Borkowski, K. Arnaud, and K. Dyer. This research has made use of data obtained through the High Energy Astrophysics Science Archive Research Center Online Service, provided by the NASA/Goddard Space Flight Center.

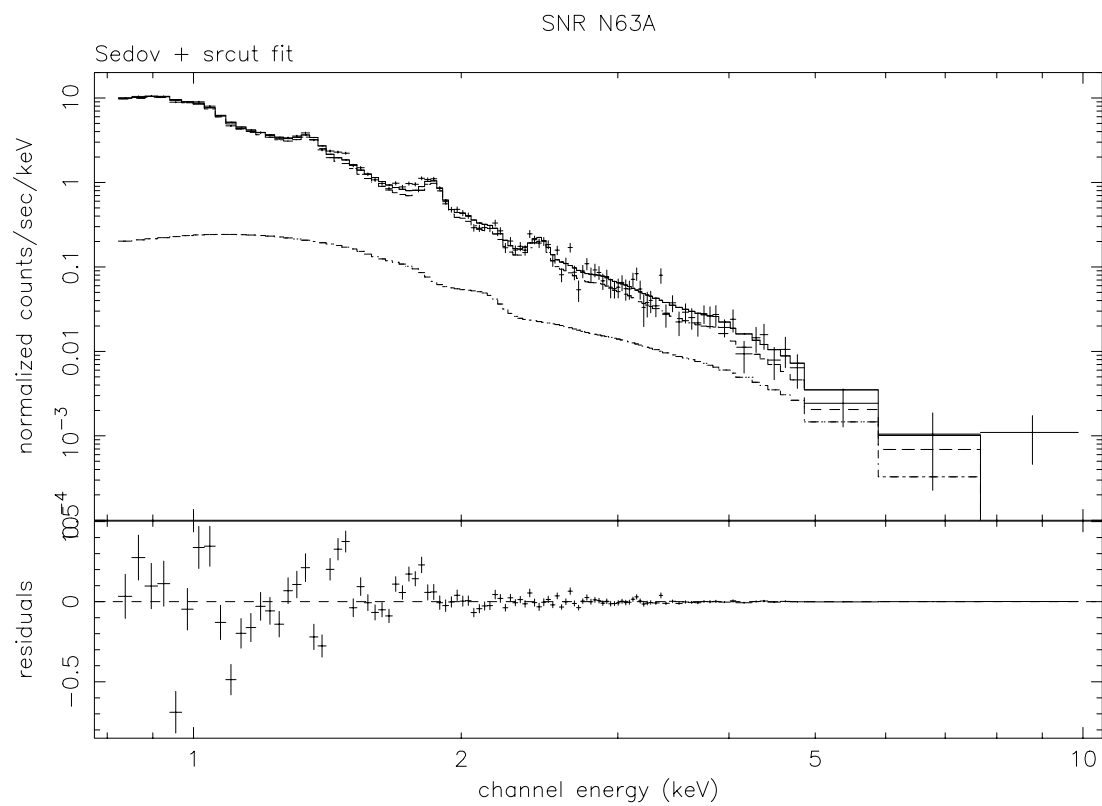


Figure 4.1: Fitted spectrum of N63A with combined SIS0 and SIS1 data. The synchrotron component is the curve with lower values at lower X-ray energies.

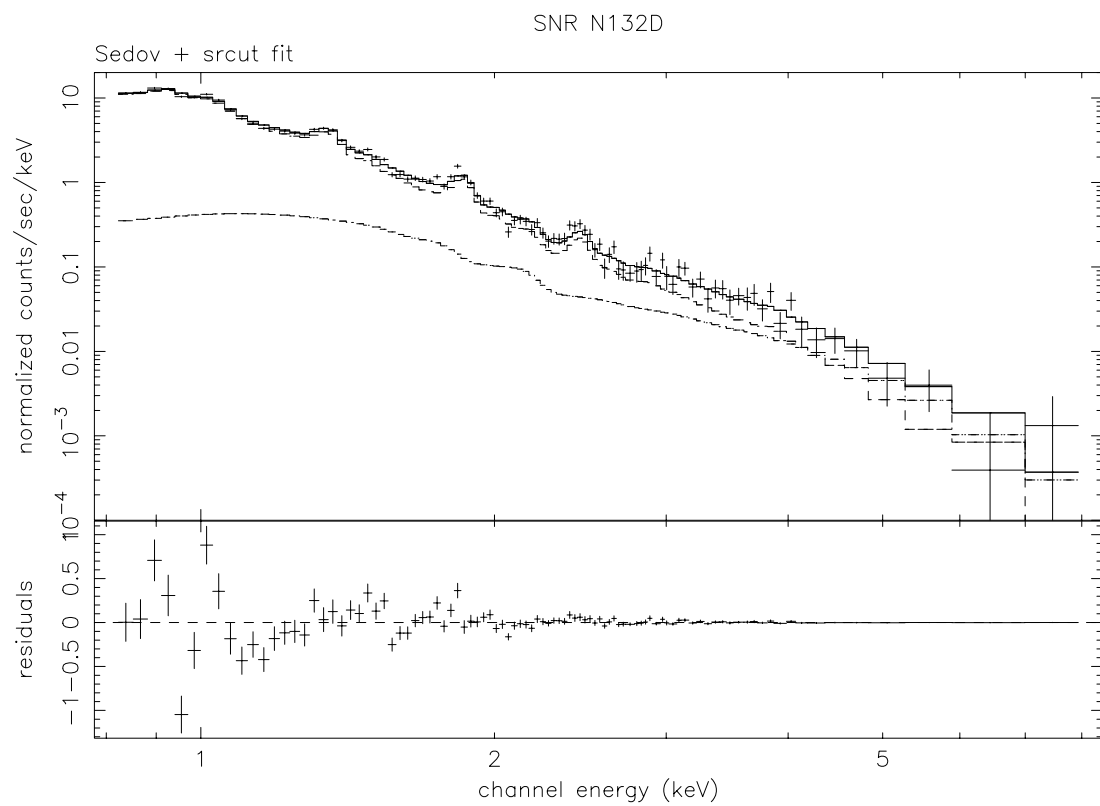


Figure 4.2: Fitted spectrum of N132D with combined SIS0 and SIS1 data. The synchrotron component is the curve with lower values at lower X-ray energies.

Table 4.1: A summary of the radio parameters for SNRs in the LMC in order of decreasing flux density. References for the flux at 1 GHz and the spectral indices: (1) Mathewson et al. (1983), (2) Mills et al. (1984), (3) Dickel & Milne (1995), (4) Dickel & Milne (1998), and (5) Filipovic et al. (1998). X-ray information is from Williams et al. (1999). The final column is the extrapolation of the synchrotron power law to 1 keV, listed in two units. For DEM L71, no spectral index has been published; we have assumed a typical value of 0.5.

Object	1-GHz Flux Jy	Spectral Index α	Radio Refs.	X-ray Size arcmin	PSPC Rate cts/s	Extrapolated 1 keV Flux Den.	
						mJy	$\frac{\text{photons}}{\text{cm}^2 \text{ keV}}$
N132D	5.6	0.58	3,5	2.2×1.7	4.5	0.077	0.12
N63A	1.9	0.53	1	1.35×1.35	2.8	0.068	0.10
N103B	0.73	0.57	3,5	0.65×0.6	0.72	0.012	0.018
N49B	0.56	0.57	4	2.8×2.4	0.57	0.009	0.014
N23	0.40	0.60	4	1.55×1.4	0.41	0.004	0.006
0453-68.5	0.24	0.40	1	2.4×2.3	0.20	0.108	0.16
0534-69.9	0.11	0.48	5	2.4×2.3	0.086	0.001	0.002
0548-70.4	0.11	0.55	5	1.9×1.8	0.059	0.003	0.004
0519-69.0	0.11	0.47	3,5	0.75×0.75	0.60	0.013	0.019
0509-67.5	0.066	0.46	1	0.6×0.6	0.20	0.009	0.014
DEM L71	0.009	(0.50)	2	1.2×1.2	0.65	0.001	0.001

Table 4.2: Summary of X-ray observations

Object	Sequence Number	Observation Dates	Exposure Time	
			SIS (ksec)	GIS (ksec)
N132D	50022000	9/23/93–9/23/93	5	6
0519–69.0	51013000	11/8/93–11/9/93	8	10
N103B	51012000	11/9/93–11/9/93	7	10
N23	51002000	11/9/93–11/10/93	21	16
0509–67.5	51003000	11/10/93–11/10/93	10	8
N63A	51004000	11/21/93–11/21/93	12	13
N49B	52010000	10/1/94–10/2/94	10	13
0548–70.4	53027000	8/24/95–8/25/95	16	16
DEM L71	53026000	10/8/95–10/9/95	13	11
0453–68.5	53018000	11/25/95–11/26/95	10	7
0534–69.9	55050000	5/12/97–5/14/97	20	20

Table 4.3: Rolloff frequencies and energy upper limits

Object	χ^2_ν	$h\nu_{\text{rolloff}}$ (keV)	3σ $h\nu_{\text{rolloff}}$ (keV)		$E_{\text{max}}\sqrt{(10\mu G/B)}$	
			This work	RK Method	(erg)	(TeV)
DEML71	1.9	1.03	1.32	9.4	130	80
N49B	3.4	0.28	0.39	0.71	72	40
N103B	3.7	0.18	0.25	0.60	57	40
N23	1.5	0.17	0.20	0.41	52	30
0509-67.5	6.9	0.12	0.16	0.18	45	30
N132D	1.8	0.13	0.15	0.27	45	30
0548-70.4	1.5	0.08	0.12	0.18	39	20
N63A	1.9	0.08	0.10	0.21	36	20
0519-69.0	4.2	0.006	0.07	0.25	30	20
0534-69.9	2.2	0.01	0.03	0.07	19	10
0453-68.5	1.2	0.004	0.02	0.03	17	10

Table 4.4: Rolloff frequency dependence on radio parameters for DEM L71

$\alpha \downarrow S_{\text{1GHz}} \rightarrow$	0.00792	0.0088	0.00968
0.40	1.03	–	0.17
0.50	–	1.03	–
0.60	1.03	–	1.50

Table 4.5: Rolloff frequency dependence on radio parameters for N63A

$\alpha \downarrow S_{1\text{GHz}} \rightarrow$	1.701	1.89	2.079
0.43	0.038	–	0.037
0.53	–	0.083	–
0.63	0.26	–	0.23

Chapter 5

On the X-ray Spectrum of LMC

SNR 0548-70.4

5.1 Introduction

Observations of supernova remnants provide important information on the evolution of the interstellar medium. SN explosions are responsible for depositing energy ($\sim 10^{51}$ ergs) and newly synthesized material into the space between the stars. The Large Magellanic Cloud (LMC) provides an excellent sample of remnants at a known distance of 50 kpc in a wide range of evolutionary stages. Young remnants are dominated by the star's ejecta, while older remnants reach the Sedov phase (Sedov, 1959) where swept-up ISM material dominates the dynamics and emission. The current generation of X-ray telescopes have excellent angular resolution that allows spatially resolved spectroscopy of SNRs. By examining the emission in different regions of a remnant we can attempt to determine if ejecta or ISM material is responsible for the observed emission. If ejecta are present, we shall seek to investigate the elemental abundances to determine the chemical composition of the stellar material. Iwamoto et al. (1999) provide calculations of ejected masses from Type II and Type Ia explosions (see Table 3.2 of this work), that can be compared to our observed results. Our goal is to use X-ray observations to type the explosion mechanism of

the supernova pregenitor.

We present a study of *Chandra* observations of the LMC supernova remnant 0548-70.4. Previous X-ray observations with ROSAT indicate a “diffuse face” morphology, where central emission is as bright as the limbs (Williams et al., 1999). Also detected in that work was a ‘gap’ region of weaker emission in the shell. Observations of this SNR in the optical by Smith et al. (1991) indicate that 0548-70.4 belongs to the class of Balmer-dominated remnants. This class of remnants exhibit strong hydrogen Balmer series emission without the forbidden emission lines of O, N, and Si seen from the radiative cooling of shocked ISM (Fesen et al., 1985) in most SNRs. A different mechanism must therefore exist in Balmer-dominated remnants. A possible explanation of this emission is nonradiative shocks (Chevalier & Raymond, 1978; McKee & Hollenbach, 1980), where a collisionless shock travels into partially neutral material and gives the optical line profile observed (Chevalier et al., 1980). Other remnants in the class include Galactic objects SN 1572 (Tycho) and SN 1006, and LMC remnants DEM L71 (0505-67.9), 0509-67.5, and 0519-69.0. Several of these remnants are thought to be from Type Ia explosions, presenting an interesting connection that, unfortunately, is not well understood. Our *Chandra* observations are the first to take a look at the structure hinted at by the ROSAT data and study the spectrum of the bright features in detail.

5.2 The Data

5.2.1 Observation

SNR 0548-70.4 was observed on 2001 September 16 for 60.1 ksec, as *Chandra* observation ID 1992. *Chandra* analysis software CIAO version 2.2.1 was used to process the data. Filtering was performed for good time interval, bad grades, observation specific bad pixels, and high background times. Notes on data reduction can be found at the *Chandra* X-ray Center (CXC) web page <http://cxc.harvard.edu> and are summarized in Appendix A. With an angular extent of ~ 100 arcsecs, SNR 0548-70.4 fit entirely within the back-illuminated S3 chip, near the telescope’s aimpoint. Based on this angular extent, we estimate the radius to be 12.1 pc, assuming a distance of 50 kpc to the LMC. Below we shall investigate different regions of the remnant by extracting spectra from each. A region

is defined from the image, and a spectrum is extracted from the filtered dataset. For proper statistical purposes, each spectrum was binned to contain a minimum of 15 counts per bin. A background spectrum must be extracted as well, to subtract diffuse X-ray background and internal telescope effects in order to clearly discern the source of emission. The background effects in *Chandra* usually vary across a chip and a blank sky background spectrum would need to be created from calibration data. This is not the case with the S3 chip, however. We were therefore able to use a source-free region of the chip as the background region for each spectrum extracted. The corresponding response files for each were created, and the ACISABS model was applied to account for quantum efficiency degradation in the detector caused by hydrocarbon contamination.

Figure 5.1 shows an image of 0548-70.4 that has been smoothed with the CIAO tool *csmooth* over a small, 2-3 pixel, scale to preserve the details of the features. The features discussed in Williams et al. (1999) can be clearly seen. The central emission lies closest to the west limb, with very dim emission between it and the bright eastern limb. The shapes on the image indicate the regions of extraction in Hendrick et al. (2003). These three bright regions were the obvious first choice for spectral analysis, but a complete study of the remnant requires that we examine several other regions. Figure 5.2 shows the remaining extraction regions: the somewhat brightened regions to the NW and SE of the central region and the eastern and western halves of the central region. Attempts were made to study the dimmest portions of the SNR in the northern and southern limbs, as well as the regions between the east limb and the central emission, but the S/N in those sections were quite poor.

5.2.2 Spectral Analysis

Limb emission in supernova remnants is presumably the result of the outer blast wave interacting with the ISM. Such emission may be well described by a Sedov model. Although this emission exists across the entire remnant, it is most pronounced in the limbs due to the path length along our line of sight through the material. Assuming Sedov dynamics would allow for some simple calculations of the remnant age, the mass of swept-up ISM, and explosion energy. Sedov modeling of SNR emission has been done in the past notably by Hamilton et al. (1983). We employ a Sedov model (Borkowski et al., 2001) created for use with the XSPEC (Arnaud, 1996) data analysis package. In order

to speed up the calculations of this model, we use a tabular form of the model which extrapolates the spectrum from a range of ten gridpoint values in each of three parameters: ionization timescale τ from $5 \times 10^{10} \text{ cm}^{-3}\text{s}$ to $1 \times 10^{12} \text{ cm}^{-3}\text{s}$, and shock temperature kT_s and electron temperature kT_e from 0.1 to 2.0 keV. One important question remains: what are the abundances in the LMC ISM? We begin our investigation by assuming that abundances are 0.4 solar values (as used in Chapter 4). Published works (Russell & Dopita, 1992; Korn et al., 2002; Peimbert, 2003) present varying results in determining LMC ISM composition, but those results indicate that abundance ratios between the heavy elements are somewhat different from their solar values. Until this question is successfully resolved, an assumption of ISM abundance is required. Fitting with a value of 0.4 for each element is an important first step, since the results would show which elements are deficient. We could then investigate enhanced abundances of an individual element by freeing it from the solar ratios implicit in our assumption. Without full knowledge of the LMC ISM composition, we must seek the best approximation for the limb emission possible.

Emission from the central regions can be more complicated in the presence of ejecta. We must remember that emission from the outer blast wave and ISM exists throughout the remnant. The first question then is to determine if the bright central regions are consistent with projected ISM emission, or a combination of that plus emission from reverse shock heated ejecta. At the ejecta-ISM interface, a reverse shock is generated which travels backwards through the ejecta. We apply a Sedov model to the central emission to see if another component is necessary. We can also use a single plane shock model with abundances set at 0.4, and investigate any overabundances. Since *Chandra* is most sensitive between 0.5-1.0 keV, Fe and O lines in that range can be indications of ejecta. If it seems that ejecta are present, we can turn to a two component model. This first component would be a Sedov model to describe the ISM emission. The second component would account for any reverse-shocked ejecta with a plane-parallel shock model. We will determine the elemental abundances of the ejecta using a shock model. First we employ the *vpshock* model in XSPEC that is characterized by an electron temperature T_e , assumed constant, and a maximum ionization timescale $\tau \equiv n_e t$ of the shocked material. All nonequilibrium ionization (NEI) models in XSPEC include modern atomic data for the Fe L-shell complex (Liedahl et al., 1995). Results of that will be compared to two-component models with the new HE shock model described in Chapter 3 of this work accounting for the ejecta. We assume all H and He have been cleared from the central regions by the blast wave and that

the elements in the ejecta are completely mixed. We confine our fits for the ejecta to only the most abundant elements that appear prominently in the X-ray spectra. Each model is used to determine O, Ne, Mg, Si, and S abundances relative to Fe. This approach should give a clear indication of the chemical composition of the ejecta.

5.3 Results

5.3.1 Limb Region Results

Our investigation begins with the bright limb regions of 0548-70.4. The E limb spectrum contained 1764 counts and the W limb had 1456 counts between 0.3 and 5.0 keV. These regions were fit with a Sedov table model with uniform abundances of 0.4. The problem with solar ratios in the LMC was highlighted in the initial fits that indicate an overabundance of Ne in the limb regions. The fits were poor in the vicinity of the Ne $K\alpha$ line at 0.9 keV. In order to improve the fit quality, we refitted the limb spectra with Ne free. The comparison of these fits are quoted in Table 5.1. Each model included two absorption components, for the LMC and the Galaxy. For the foreground Galactic absorption toward the LMC, we used the hydrogen column density of $7 \times 10^{20} \text{ cm}^{-2}$ based on observations of diffuse neutral hydrogen in our Galaxy (Dickey & Lockman, 1990). The derived LMC hydrogen column density N_H is higher than average for the LMC, particularly for 0548-70.4. Observations of the 21 cm HI line emission in this direction reveal a substantial ($\sim 3 \times 10^{21} \text{ cm}^{-2}$) amount of HI in the LMC (Luks & Rohlfs, 1992, scaled up by 50% according to Staveley-Smith et al., 2003), and additional absorption may be caused by ionized and/or molecular gas.

For 0548-70.4, Smith et al. (1991) determined a shock speed v_s of 670 km s^{-1} for the E limb material from the observed width of the broad component of the $H\alpha$ line (assuming no collisionless electron heating at the shock). This shock speed was used to constrain the fit in the E limb, where the optical emission was detected. The previous fit of the E limb had better χ^2/DOF statistics, but was poorly constrained, and error analysis failed to get well bounded results. Table 5.1 summarizes all the fits to the limbs discussed so far. The fits with error ranges included on the table are our choices for the best description of the limb emission. The fitted values for the W limb are consistent, within errors, with

the E limb results. The results of these fits can be seen in Figures 5.3 and 5.4. In each Sedov model, the initial electron temperature was essentially zero, with upper limits quoted in Table 5.1. Those upper limits are lower than the shock temperature values indicating a lack of equilibration between ions and electrons at the shock front. We derive the shock age from Sedov dynamics by $t = 2R/5v_s = 7,100$ yr for 0548-70.4. We calculated densities in two ways, from the emission measure parameter assuming an emitting volume of 2.2×10^{58} cm³ in the E limb, and from the relationship between age and ionization timescale. We approximated the limb volume as the cap of a spherical shell of the radius of the SNR, with thickness and angular extent of the cap determined from the regions of extraction. For the E limb, we found a thickness of 1.2×10^{18} cm and an angular extent of 90 degrees, which for a distance of 50 kpc give the volume we quote. With an ambient density of 0.18 cm^{-3} , we determine the total mass of the swept-up ISM to be $50M_{\odot}$. Still assuming Sedov dynamics, the explosion energy is $E = (R/1.15)^5 \rho / t^2 = 3.0 \times 10^{50}$ ergs for 0548-70.4. This apparently low explosion energy could be the result of energy going into cosmic ray acceleration.

5.3.2 Central Emission Results

We begin our investigation with the large central region, seen in Figure 5.1. This region has 8723 counts and a spectrum that seems to have enhanced level of several elements. Figure 5.5 compares the central spectra to the E limb. Notice the large difference in emission in the 0.55-1.5 keV range, as well as the strength of the Si line at 1.9 keV. Single-component *vpshock* fits to the center of 0548-70.4 with all the abundances linked were poorly constrained. Freeing Fe from the solar ratios, but keeping the other abundances linked to one another, we get results, and 90% confidence ranges, of 0.96 (0.80-1.2) for Fe and 0.52 (0.42-0.72) for the rest of the elements, with $\chi^2/\text{DOF} = 165/86$. That is, iron abundance relative to the other heavy elements is roughly twice its solar ratio. This result indicates that there is Fe emission beyond the LMC ISM contribution in the central region, so we will add another model component to account for this excess. The two-component fit results of the central region can be seen in Figures 5.6 & 5.7. The lower curve represents the Sedov model component which is based on the fit to the E limb (column 5 Table 5.1) with only the normalization allowed to vary in the two component fit. The results of the two-component fitting, one plane shock model plus the Sedov model component for blast-wave emission projected against the center, are found Table 5.2. In the central region we find

low ionization timescales and shock temperatures similar to the limb values. Fe, O, Mg, Si, and S emission is prominent. We also allowed neon to be fit in the central regions, but only a small amount was needed in 0548-70.4. Results using the new heavy-element model agree closely with the *vpshock* results, falling within the 90% confidence ranges quoted. Computational constraints do not allow more detailed error analysis of the HE shock at this time. With the heavy-element model we determine the speed of the reverse shock to be 420 km s^{-1} for the central region of 0548-70.4.

This rather large region was broken up into eastern (EC) and western halves (WC) to search for any spectral variations. Figure 5.8 compares the two spectra. For single component fits we linked elemental abundances together, except for Fe. Iron is freed from solar ratios to investigate the presence of ejecta. The results of that were values, and 90% confidence ranges, of $\text{Fe} = 0.84$ (0.62-1.3) and abundances of 0.48 (0.30-0.92) in EC, and in the WC $\text{Fe} = 0.90$ (0.63-1.4) with the other elements at 0.52 (0.32-0.84). These fits were poorly constrained, but these numbers again confirm that elevated levels of Fe exist. We next repeat the two component fits described above on these two regions with both *vpshock* and the HE shock model. Those results are also found in Table 5.2. The initial fits find less oxygen in the WC, closer to the center of the remnant, with higher levels of sulfur when compared to the large central region. The EC exhibits the opposite, with higher oxygen levels and lower sulfur closer to the limb. Within the error ranges quoted in Table 5.2, however, these regions are consistent with each other and the large central region.

We expect an O/Fe ratio of 70 (by number) for a $20 M_{\odot}$ core-collapse explosion, and 0.75 for a Type Ia explosion (Iwamoto et al., 1999). For the central region of SNR 0548-70.4 we find a ratio of 16 (WC also has O/Fe = 16, with EC slightly less at 13.4), much higher than expected for a Type Ia remnant as the optical data suggest. One explanation could be that O is located in denser ejecta than Fe (Dwarkadas & Chevalier, 1998). Particularly strong O lines are expected in deflagration models of Type Ia explosions, with O/Fe emission measure ratios of ~ 100 (Badenes et al., 2003). However, plane-shock models assume a uniform density for each element, so differences in emission measure are interpreted as differences in abundances. The residual pattern in the O complex at 0.55-0.65 keV seen in each spectrum is a sign that the model underpredicts O $K\alpha$ and overpredicts O $L\alpha$ emission, compared with the data. This indeed hints at the presence of cool and dense O-rich ejecta unaccounted for in the simple plane-shock model. If we interpret the O/Fe values in Table 5.2 as the ratios of emission measures instead of abundances, there is no

problem anymore with a high O abundance, allowing us to preserve the interpretation of 0548-70.4 as a Type Ia remnant. Another scenario producing a relatively high O abundance in a Type Ia remnant is that if a substantial amount of shock energy is being transferred to cosmic rays, our estimate of the shock velocity based on the fitted electron temperature is too low, and our age estimate too high. Then the apparent O/Fe abundance ratio would increase because not all Fe might have been shocked by the present time. In either case, the O/Fe ratio determined would not be as high as the 70 expected for Type II explosions.

To determine the mass of the ejecta, we examine the upper-limit case where the filling fraction, f , is equal to 1, and use the emission measure for each element and the volume. The central region was approximated as a sphere of radius 1.8×10^{19} cm, giving a volume of 2.5×10^{58} cm³. Volumes of the western and eastern halves of the center are assumed to be 1.25×10^{58} cm³. In Table 3, the EM_{EJ} term is defined as $\int n_e n_i dV$, as is appropriate for a heavy element dominated shock. In order to calculate the contribution of each element to the emission measure, we can replace n_i with n_{elem}/n_i , where n_{elem} is the ion density of a single element. The ratio of electron to ion density is calculated by the HE shock model, so we can determine the value of n_e from the emission measure and the volume. With $n_e/n_i = 8.3$, we find that n_e is equal to 0.026 cm⁻³. The electron density relates the emission measure of a single element to its total number N_i , so we can sum the contributions of each to the total mass. Our upper-limit calculation of the central region of 0548-70.4 gives $1.28M_\odot$ of ejecta. In the EC, the HE shock model gives $n_e/n_i = 8.6$. That corresponds to an electron density of 0.023 cm⁻³, and an ejecta mass of $0.44M_\odot$. Similarly, in the WC we find $n_e = 0.025$ cm⁻³, when $n_e/n_i = 8.4$, which gives an ejecta mass of $0.50M_\odot$.

We have assumed that the ejecta are pure heavy elements with no mixing of hydrogen and helium, though both remnants are old enough to have possibly undergone significant mixing. Mixed H and He in the ejecta will increase the inferred electron density and lower the density of the heavy ions, for the same observed spectrum. We can estimate how much H and He might be present by comparing the ram pressure ρv^2 of the blast wave to that of the reverse shock, using our fitted shock speeds and calculated densities. The difference in ram pressures is 1.7×10^{-9} dyne cm⁻² between the central region and the outer blast wave (the blast-wave pressure being larger). Since we expect approximately uniform pressure in the remnant interior, we attribute the deficit in reverse-shock pressure to our not including H and He in the ejecta, and determine the densities of those elements needed

for pressure balance. Our pressure calculation for the blast wave is based on the outermost portion of the shock. Pressure should decrease somewhat moving inward, so these values represent an upper limit to the amount of H and He mixing. The amount of mixing we infer has little effect on the plane-shock fits to the centers, so we can repeat our mass calculations. With maximum mixing, our electron density is now 0.98 cm^{-3} and our lower limit on ejecta mass (times $f^{1/2}$) becomes $0.41 M_{\odot}$ in 0548-70.4. Our results in EC and WC were not as consistent as before; the reason can be seen in Table 5.3. The EM_{EJ} values for the HE fits are the expected value of nearly half the center value. This is not the case with the *vpshock* emission measures, so we cannot compare the mixing calculations in these regions.

The next region to examine is the SE central region, as seen in Figure 5.2. This region contains 1872 counts. As with the other central regions, a Sedov model failed to describe the emission. A single component *vpshock* model with abundances linked except for Fe, was poor. In this case we have to freeze the linked elements to 0.4, and fit only for Fe. That single component fit had $\chi^2/\text{DOF} = 83/53$, and values, with 90% confidence ranges, of 1.29 (1.07-1.56) for Fe. Once again we find indication of the presence of ejecta. Although there is an indication of elevated abundances in this region, the low counts made any two-component fitting impractical.

One outer region in 0548-70.4 that does not meet the limb profile is the bright NW region from Figure 5.2. This appears to be connected to the central emission in the image, despite its position on the limb. Figure 5.9 compares spectra from this region (with 1944 counts) to the nearby W limb region. The most notable difference is the emission between 0.7 and 0.9 keV, the range for Fe L-shell lines. The Sedov model fit to this region was quite poor, so we turn to a single plane shock model, *vpshock*. Table 5.3 shows the plane shock results. We freed Ne to mimic the Sedov model fitting, and investigated the effects of freeing Fe from the solar ratios. The results of the Fe free fit did not improve much over the fit with only Ne free, and barely elevated Fe above 0.4. Since the Ne free fits are consistent with our limb assumptions, the emitting material could simply be ISM. To investigate any ejecta components, a two-component fit with the W limb Sedov model fit and the central region *vpshock* component result. The S/N in NW limb did not allow for fitting each parameter, but the normalizations of each component were fit with $\chi^2/\text{DOF} = 66.5/54$. Figure 5.10 shows that this region could be described as a superposition of blast wave and ejecta emission where in this case the ISM material is the dominant contributor. The analysis of this region could be complicated by the presence of nonradiative shocks.

Smith et al. (1991) identifies it as the brightest source of the Balmer line $H\alpha$ emission in the SNR. It exhibits stronger optical forbidden-line emission than the rest of the remnant, which leads those authors to suggest that 0548-70.4 is in transition from the nonradiative shocks, to the radiative shocks normally seen in SNRs.

5.4 Discussion

We have used plane shock models to describe emission from ejecta in SNR 0548-70.4. It is important to recognize the limitations of these models. Both *vpshock* and the HE shock oversimplify the ejecta structure by assuming uniform density and full mixing of the elements. The presence of oxygen clumps was addressed, but we did not attempt to account for possible depletion of heavy elements, either by dust formation or the fact that all the ejecta have yet to be heated by the reverse shock. Dwarkadas & Chevalier (1998) find for Type Ia remnants much denser ejecta near the contact discontinuity than at the reverse shock. Based on our swept up mass estimates and explosion energy, assuming $1.4M_{\odot}$ of ejecta with the density profile from Dwarkadas & Chevalier (1998), their calculation indicates that the reverse shock would have propagated to the center of 0548-70.4 in 6200 yr. We calculated an age of 7100 yr for this remnant and claim reverse shock heated ejecta are still present. This discrepancy could be accounted for by a substantial amount of explosion energy transferred to cosmic rays at the forward shock, or by improper interpretation of the plane shock model results.

SNR 0548-70.4 in the LMC has well defined limbs that are consistent with a Sedov blast wave. The central regions of this object show enhanced emission that two component fits show can be explained as ejecta from a Type Ia supernova explosion. This assessment agrees with the apparent link between Balmer dominated SNRs and Type Ia events. In the brightest region of Balmer emission, the NW limb, we find a spectrum that varies from the other limb regions. There is an indication the superposition of ejecta and blast wave emission in the NW limb. The presence of nonradiative shocks in that region could have an affect as well. Type Ia progenitors modify their environments prior to explosion. Polar out-flow from the white dwarf's accretion disk, or the planetary nebula formed from its outer layers could have significant effects on the local ISM. The fact that a plane shock model improves over Sedov models in the NW limb is a sign that nonradiative shocks are

present. Overall our results show that ejecta detection is possible for mature SNRs, and that our tools can help determine progenitor type.

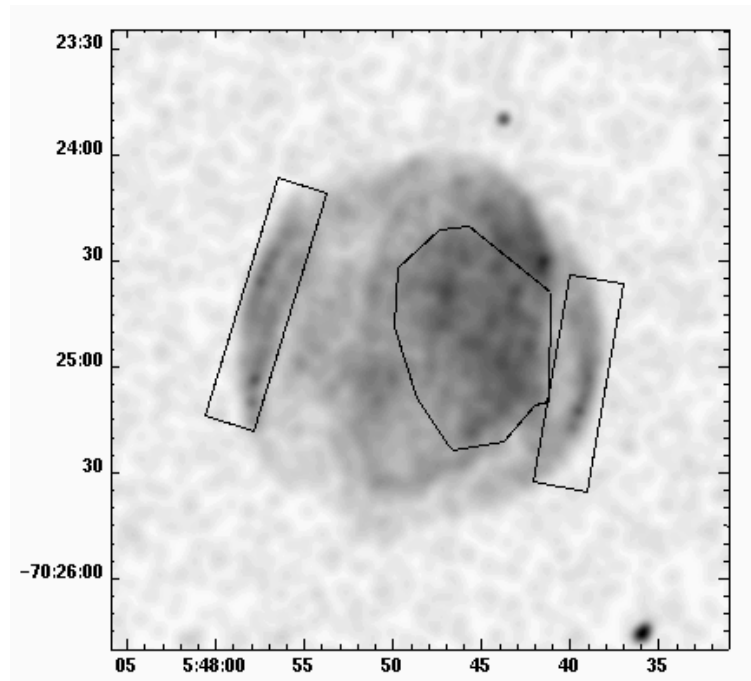


Figure 5.1: Image of 0548-70.4, with regions from Hendrick et al. (2003).

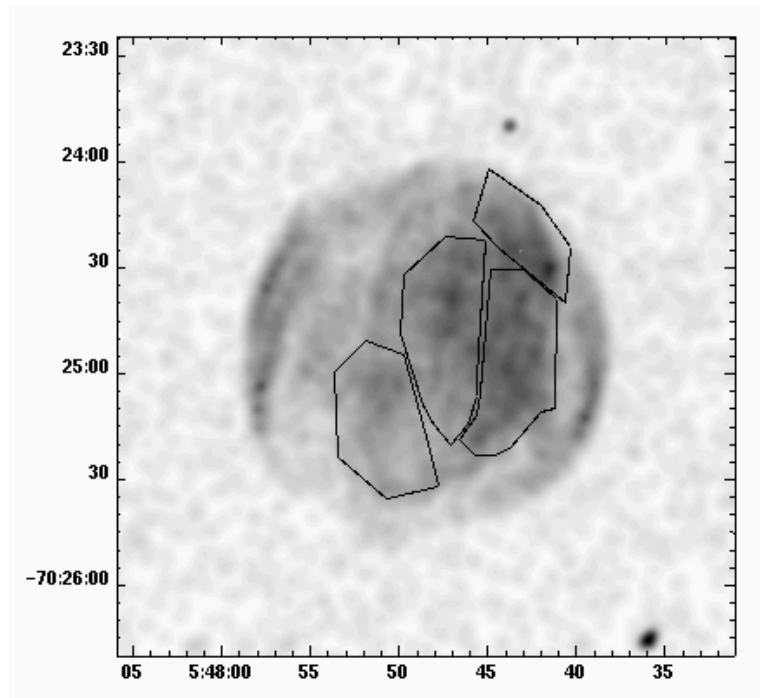


Figure 5.2: Same as Figure 5.1, with the NW and SE regions and the central region halves.

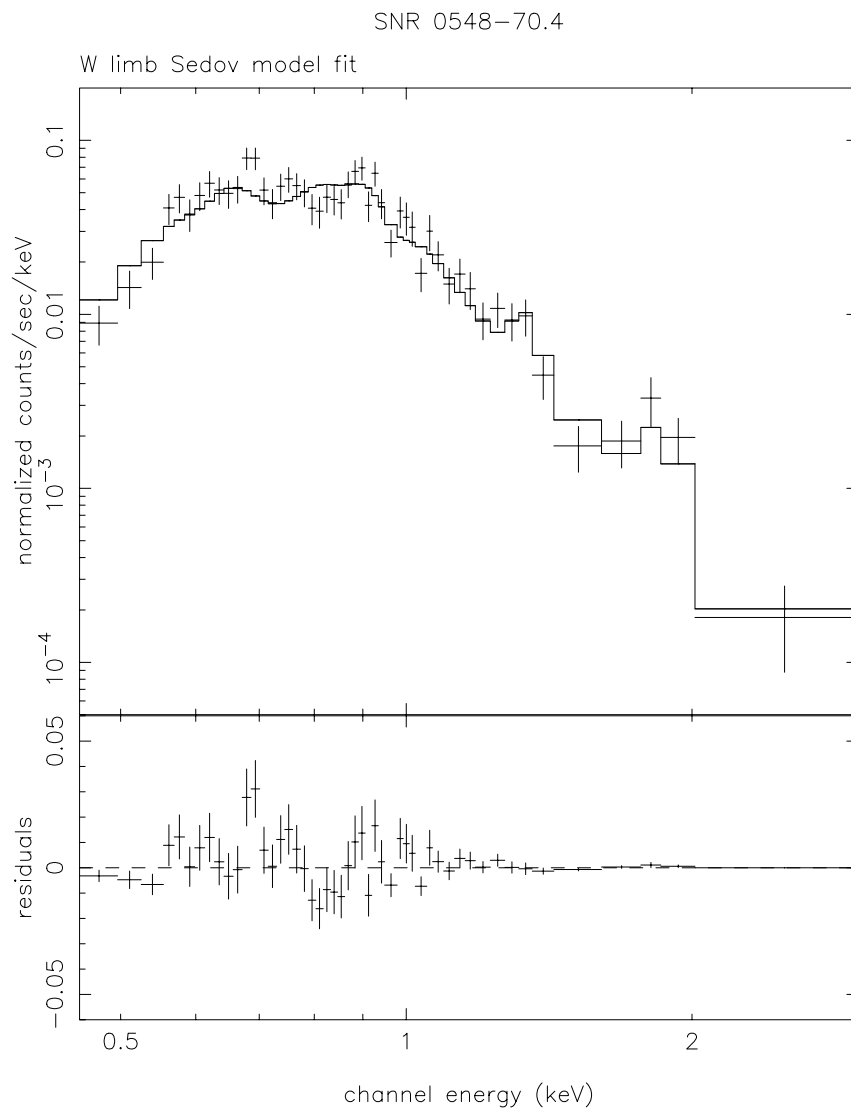


Figure 5.3: Sedov model fit to West limb spectrum.

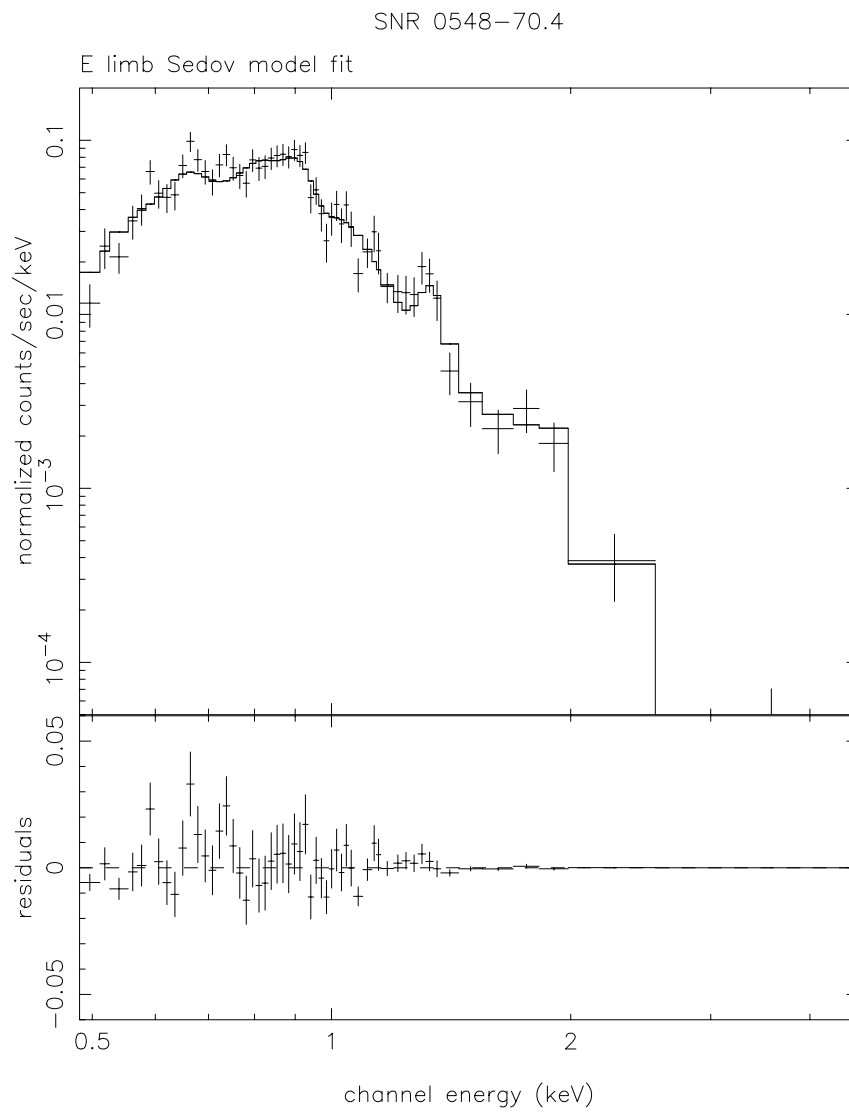


Figure 5.4: Sedov model fit to East limb spectrum.

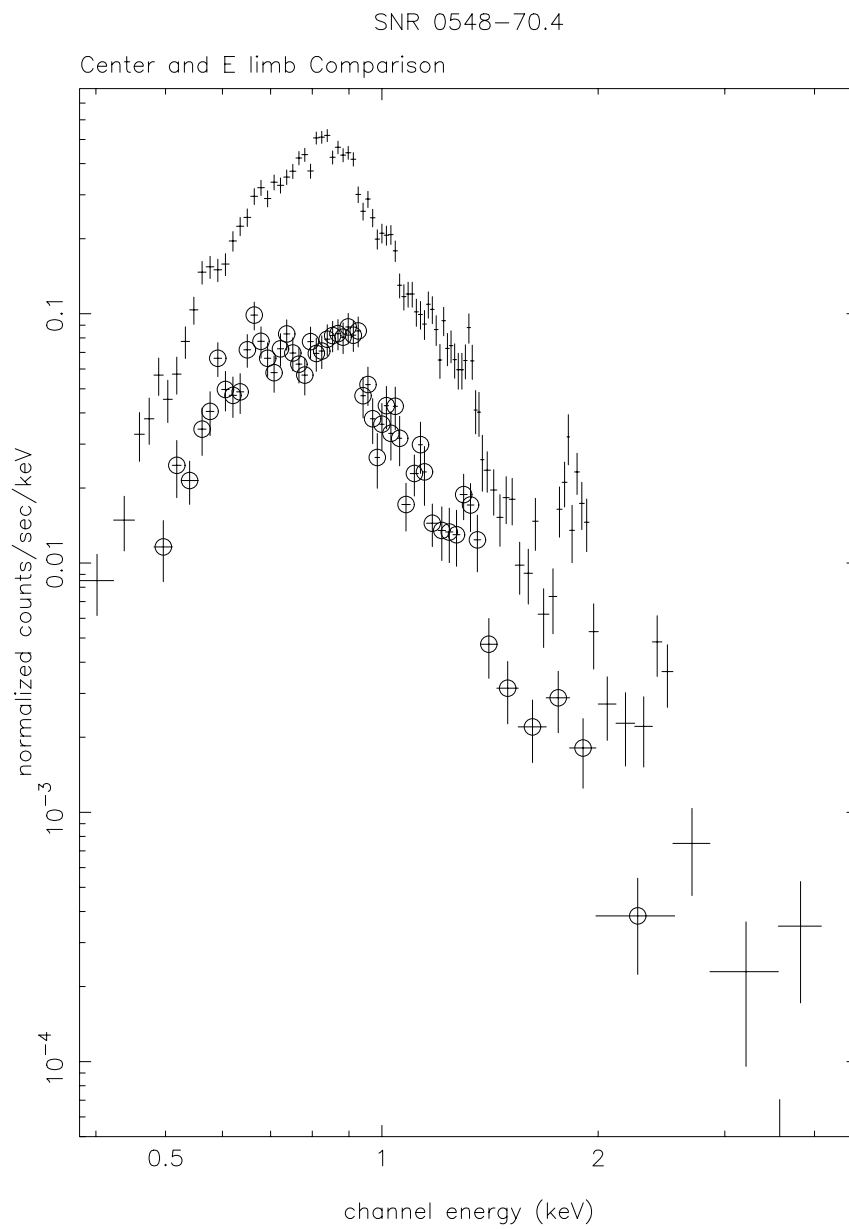


Figure 5.5: Comparison of Central region to the East limb (circles).

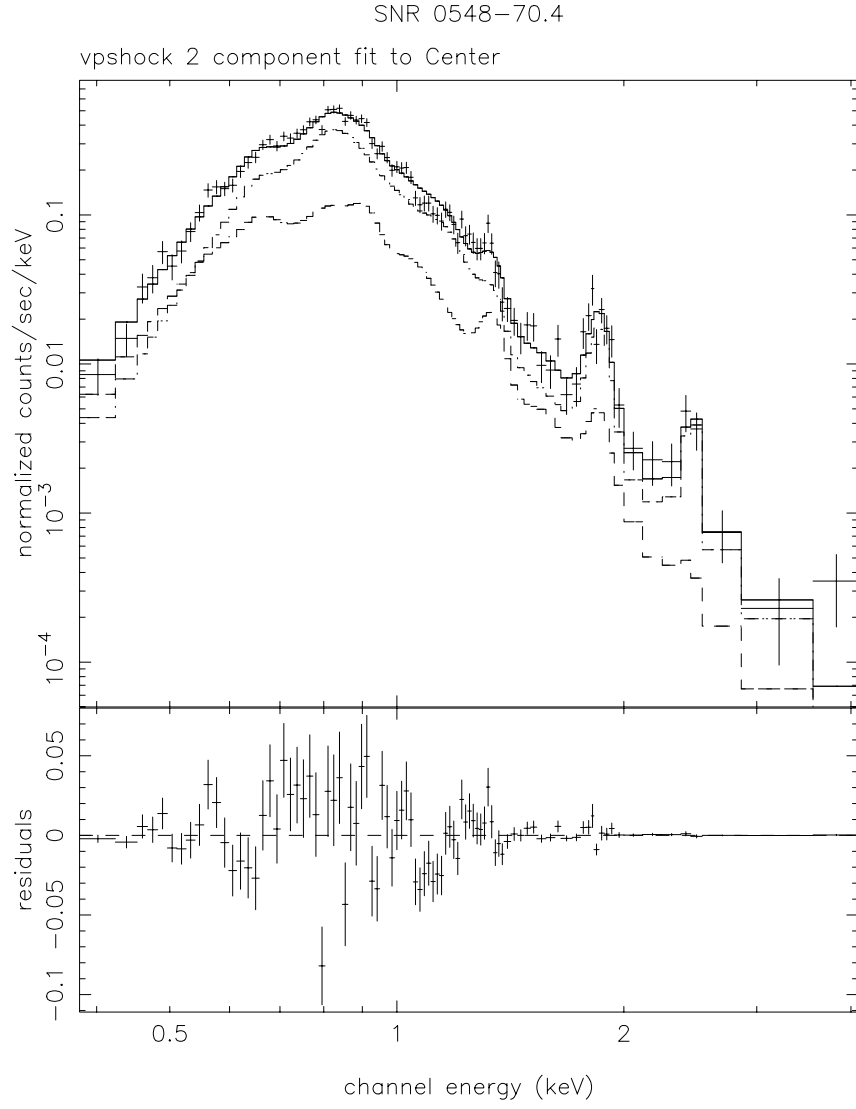


Figure 5.6: Two-component fit to the Center region. The curve through the data is the fit, with the two components shown separately as the two lower curves. The lower of the two curves represents the Sedov model component, which matches the limb spectrum in shape, and the upper curve is the *vpshock* model.

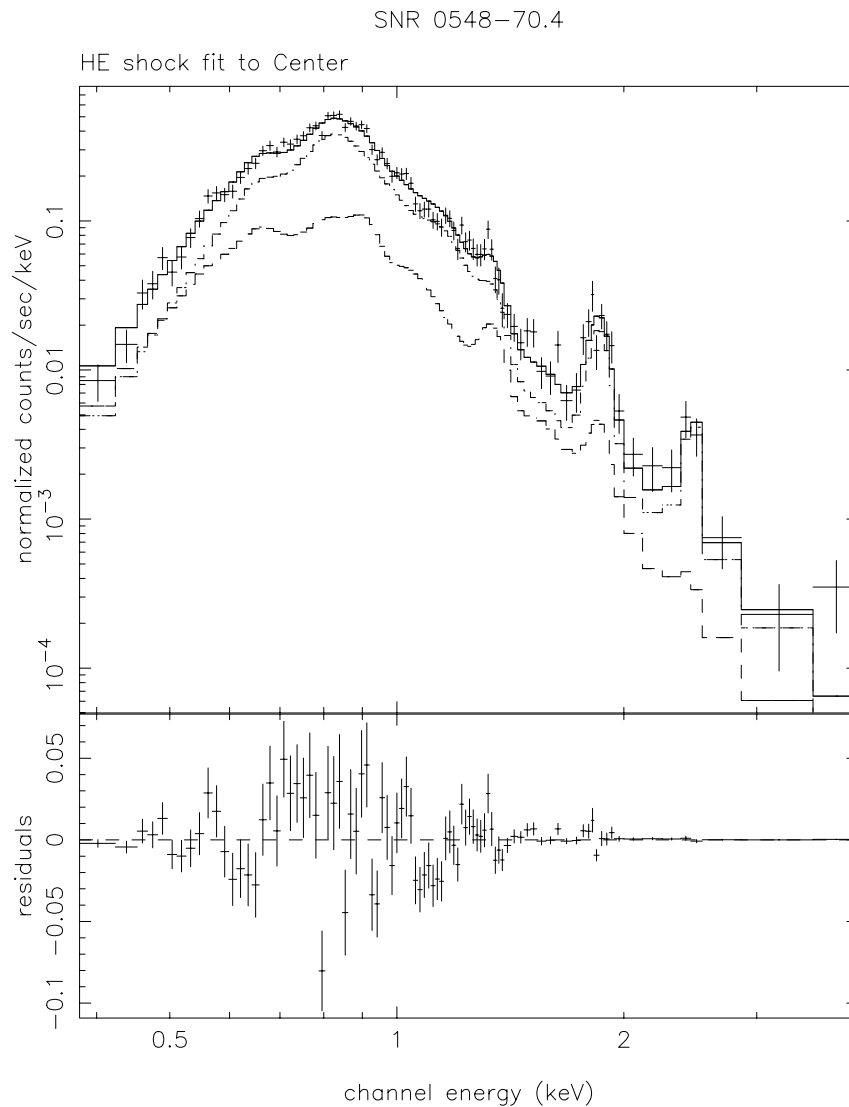


Figure 5.7: Two-component HE shock fit to Center. The curve through the data is the fit, with the two components shown separately as the two lower curves. The lower of the two curves represents the Sedov model component, which matches the limb spectrum in shape, and the upper curve is the HE model.

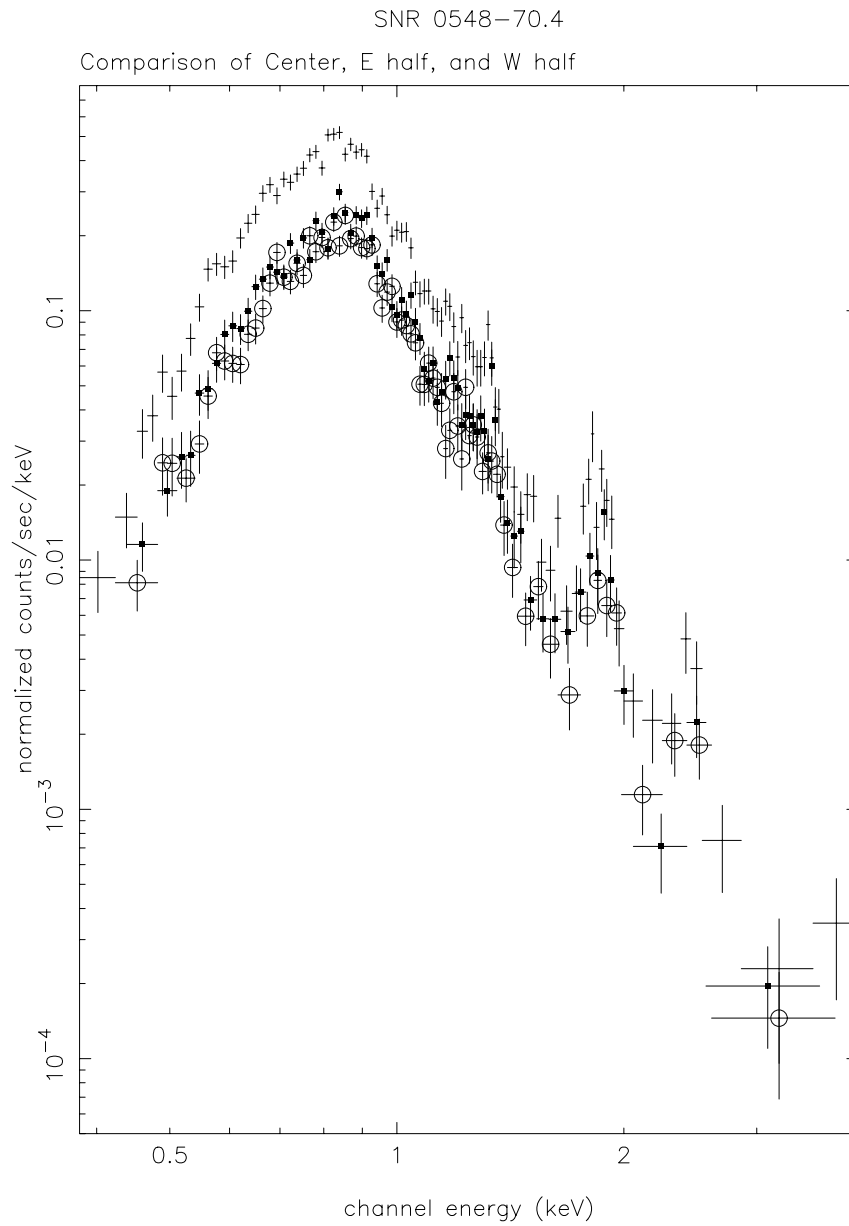


Figure 5.8: Comparison of the Central region to EC (circles) and WC (squares) halves.

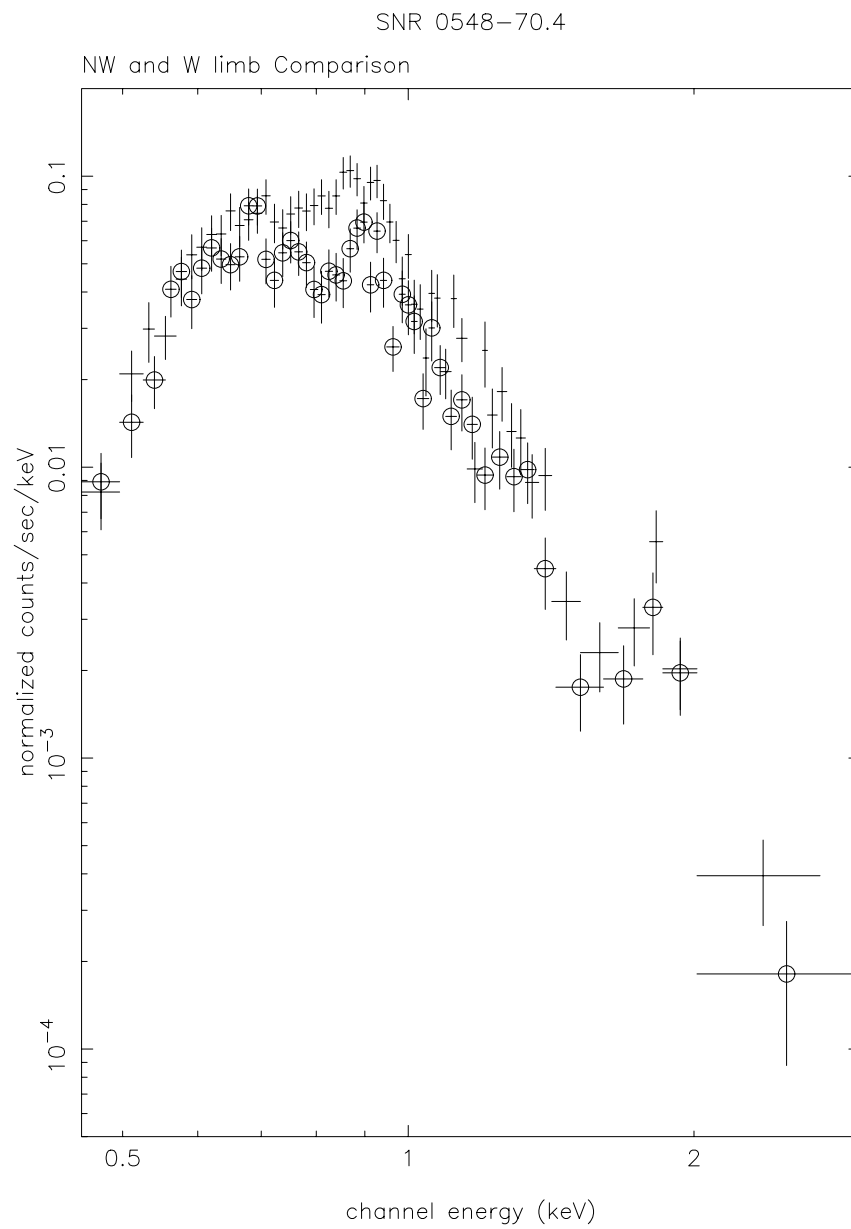


Figure 5.9: Comparison of the NW region and the West limb (circles).

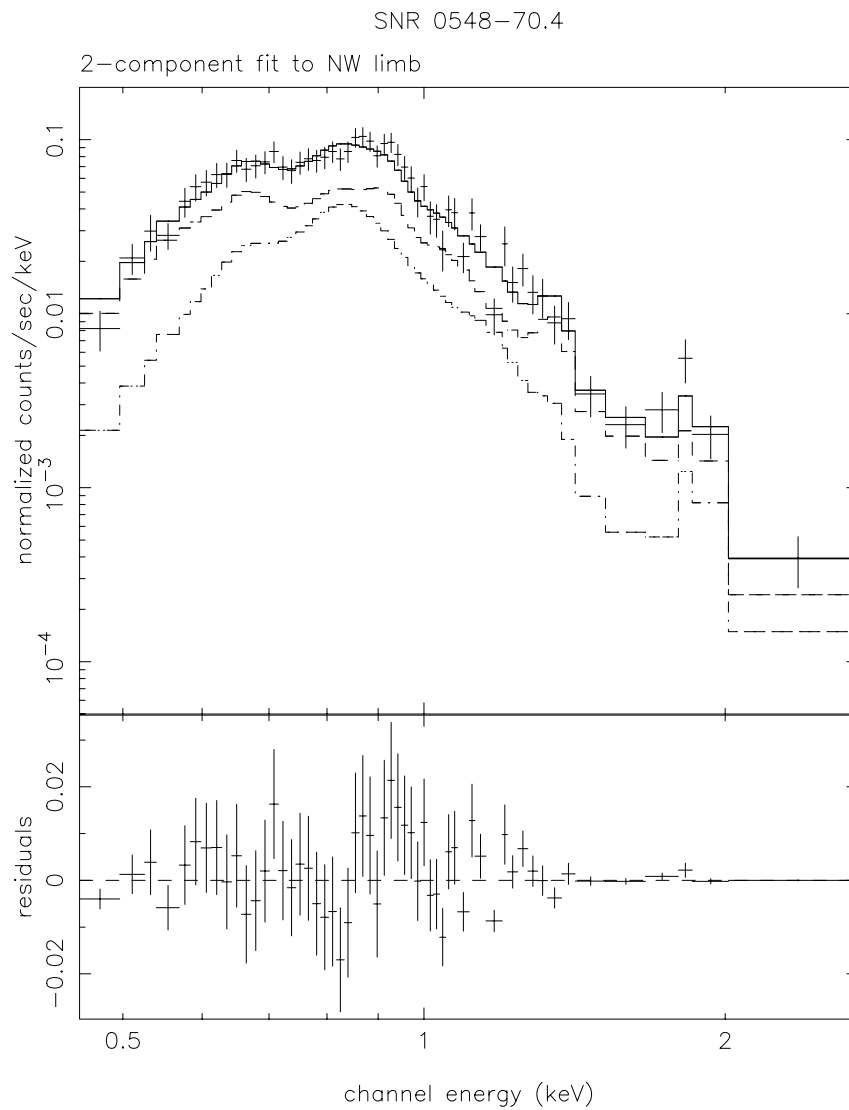


Figure 5.10: Two-component *vpshock* fit to the NW limb. The curve through the data is the fit, with the two components shown separately as the two lower curves. The lower of the two curves represents the shock model, and the upper curve is the Sedov model component.

Table 5.1: Sedov model fits to limb regions. Columns 3 and 5 are fit with Ne free, while column 6 freezes T_s to Smith et al. (1991) value.

Parameter	WL	WL+Ne	EL	EL+Ne	EL+ T_s
χ^2/DOF	71.4/45	69.7/44	63.6/51	61.6/50	63.7/51
$N_H/10^{21} \text{ cm}^{-2}$ (LMC)	3.35	$3.67^{+1.9}_{-1.1}$	6.80	6.73	$5.22^{+1.1}_{-1.2}$
Abund	0.4	0.4	0.4	0.4	0.4
Ne	-	$0.52^{+0.17}_{-0.15}$	-	0.50	$0.55^{+0.15}_{-0.11}$
kT_s (keV)	0.66	$0.79^{+0.96}_{-0.54}$	0.34	0.25	0.54
kT_e (keV)	0.0	$0.0^{+0.44}$	0.10	0.0	$0.0^{+0.2}$
$\tau/10^{11} \text{ cm}^{-3}\text{s}$	2.64	$1.89^{+2.41}_{-1.15}$	1.87	2.64	$1.89^{+0.95}_{-0.71}$
$(EM/4\pi d^2)/10^{10} \text{ cm}^{-5}$	1.63	1.64	11.3	14.4	4.37

Table 5.2: Two-component fits to central regions (EM units are cm^{-5})

Parameter	C vps	C HE	EC vps	EC HE	WC vps	WC HE
χ^2/DOF	146/83	148/83	98/64	99/64	104/68	107/68
$(EM_L/4\pi d^2)/10^{10}$	$7.1^{+2.6}_{-2.7}$	6.5	$3.5^{+1.5}_{-0.9}$	2.8	$3.3^{+1.7}_{-1.7}$	3.2
kT_e (keV)	$0.62^{+0.04}_{-0.03}$	-	$0.58^{+0.04}_{-0.06}$	-	$0.62^{+0.06}_{-0.04}$	-
v_s (km s^{-1})	-	423	-	406	-	424
O/Fe	$20.22^{+6.15}_{-6.04}$	15.6	$14.4^{+7.6}_{-7.5}$	13.4	$22.6^{+9.5}_{-9.3}$	16.0
Ne/Fe	$0.32^{+0.58}_{-0.08}$	0.15	$0.53^{+1.01}_{-0.11}$	0.33	$0.58^{+0.84}_{-0.14}$	0.35
Mg/Fe	$0.18^{+0.09}_{-0.08}$	0.20	$0.17^{+0.14}_{-0.11}$	0.19	$0.21^{+0.12}_{-0.14}$	0.22
Si/Fe	$0.46^{+0.12}_{-0.12}$	0.44	$0.45^{+0.25}_{-0.18}$	0.43	$0.52^{+0.20}_{-0.17}$	0.52
S/Fe	$0.66^{+0.33}_{-0.26}$	0.62	$1.2^{+0.79}_{-0.58}$	0.97	$0.62^{+0.46}_{-0.38}$	0.62
$\tau/10^{11} \text{ cm}^{-3}\text{s}$	$4.15^{+1.99}_{-1.40}$	2.84	$5.34^{+4.62}_{-1.94}$	3.37	$4.72^{+4.78}_{-2.07}$	3.07
$(EM_{EJ}/4\pi d^2)/10^6$	$8.62^{+0.9}_{-0.8}$	6.95	$8.43^{+1.2}_{-1.1}$	2.55	$10.5^{+1.5}_{-1.6}$	3.16

Table 5.3: Fit summary of NW limb region

Parameter	<i>vpshock</i> +Ne	<i>vpshock</i> +Ne+Fe
χ^2/DOF	65.1/51	65.0/50
$N_H/10^{21} \text{ cm}^{-2}$ (LMC)	4.69	4.74
Abund	0.4	0.4
Ne	0.77	0.77
Fe	-	0.42
kT_e (keV)	0.46	0.46
$\tau/10^{11} \text{ cm}^{-3}\text{s}$	1.16	1.06
$(EM/4\pi d^2)/10^{10} \text{ cm}^{-5}$	3.30	3.24

Chapter 6

On the X-ray Spectrum of LMC

SNR 0534-69.9

6.1 Introduction

Our next target, SNR 0534-69.9, is a middle-aged, Sedov-phase remnant in the LMC. The ROSAT observation reviewed in Williams et al. (1999) describes this remnant as having a “diffuse face” morphology within a well defined shell. This is not a particularly well-studied remnant. Our *Chandra* observation is the first to describe the emission in detail. We shall perform spatially-resolved spectroscopy on different regions of this remnant to determine if a Type Ia or core-collapse explosion occurred.

6.2 Data and Analysis Technique

Our *Chandra* observation of SNR 0534-69.9 occurred on 2001 September 1 for 60.2 ksec (OBSID 1991). For its angular extent of ~ 120 arcsec, the radius is estimated to be 13.7 pc, assuming a distance of 50 kpc to the LMC. The target fit entirely within the back-illuminated S3 chip, allowing the same processing and filtering methods described in Chapter 5 and Appendix A with CIAO 2.2.1 software. Different regions for investigation

have been identified in the complex morphology of 0534-69.9. There are some clearly defined features that can be seen in Figures 6.1-6.3. Each image has been smoothed over a small, 2-3 pixel, scale with the CIAO tool *csmooth*. Spectral extraction regions are drawn in each figure. Figure 6.1 shows the bright NW limb region, the large, dim southern region, and the small, brightened eastern limb. Figure 6.2 zooms into the NW quarter of 0534-69.9 that has several interesting structures. A bright line of emission exists directly behind the NW limb. There is also a region of emission in the western limb that loops inwards towards the center of the remnant. The circular region is where these two regions come together with the NW limb, which we have dubbed the western knot. Central emission regions are found in Figure 6.3. The brightest central emission is in the nearly circular regions in the NE and S. Examination of those regions was the focus of ejecta detection efforts in Hendrick et al. (2003). Here we include emission from the dimmer central regions in the NW and SE to complete the study.

Spectral analysis will proceed as it did in Chapter 5. Limb emission is presumed to be consistent with the blast wave interacting with the LMC ISM. Sedov table models (Borkowski et al., 2001) will be applied to each limb region in Figure 6.1. We will also use this Sedov model to investigate the regions in Figure 6.2, where the features could be projection effects of the outer blast wave. If the the results are poor for these Sedov fits, we will turn to a parallel plane shock model, *vpshock* in XSPEC (Arnaud, 1996). With this model we vary the abundance of individual elements, particularly Ne and Fe, to detect any enhancement beyond solar ratios. This is also the first step in the study of the central regions. We can examine the overabundance of Fe by freeing it from the solar ratios of the other elements in that model. Enhanced Fe emission is a sign that there may be ejecta present, so we will then use two-component models. The first component is the Sedov model fit to a nearby limb region to account for the ISM emission from the parts of the shell along the line of sight. The second component will then account for any excess of emission beyond the shell contribution that we attribute to ejecta. First we run the Sedov model plus *vpshock*, and we compare those results to another two-component model of the same Sedov fit now with the HE shock model described in Chapter 3 accounting for the ejecta. For all of the models described above, we used a two-component absorption model. The foreground Galactic absorption toward the LMC is fixed at a hydrogen column density of $7 \times 10^{20} \text{ cm}^{-2}$ based on observations of diffuse neutral hydrogen in our Galaxy (Dickey & Lockman, 1990). We then used a second component, with abundances set to 0.4 solar,

to determine the column density through the LMC (the values of N_H quoted in our results tables are this second component).

6.3 Results

6.3.1 Limb Emission

Table 6.1 lists the Sedov model fit results for the limbs of 0534-69.9. The bright NW limb contains 4395 counts, the large dim S limb region contains 3912 counts, and the small E limb has 1466 counts. We began fitting with abundances set at 0.4 solar, but the initial fits were poor near the Ne $K\alpha$ line. As with SNR 0548-70.4, Sedov model fits improved when Ne was allowed to float beyond solar ratios. This is an indication of the uncertainty in the actual abundance ratios in the LMC. The fits in Table 6.1, with Ne free, are our best approximation of the limb emission until the question of the LMC ISM is resolved. Figure 6.4 depicts the Sedov model fit to the bright NW limb. Mean shock temperatures for each region are approximately 0.34 keV. The electron temperatures were essentially zero with upper limits well below the lower limits on the shock temperatures. We derive the shock age from Sedov dynamics by $t = 2R/5v_s = 10, 100$ yr for 0534-69.9, with a shock speed of 530 km s^{-1} in the NW limb. Shock speed is calculated from the shock temperature, $v_s = 9.1 \times 10^7 \sqrt{kT_s}$ (for kT_s in keV) where $\mu \sim 1.3$ is consistent with solar abundance ratios assumed in the LMC ISM. Similar numbers for shock speed and age were found using the S limb and E limb results. We calculated the ambient density in the NW limb to be 0.17 cm^{-3} , and with an emitting volume of $1.0 \times 10^{58} \text{ cm}^3$, we determine the total mass of the swept-up ISM to be $60M_\odot$. (We modeled the limbs as a cap of a spherical shell assuming a distance of 50 kpc, with thickness and angular extent for the NW limb of $8.3 \times 10^{18} \text{ cm}$ and 60 degrees.) The explosion energy of 0534-69.9 is 2.6×10^{50} ergs where $E = (R/1.15)^5 \rho / t^2$. Again, we find an explosion energy well below the expected 10^{51} ergs. Each limb region Sedov parameter is within the 90% confidence ranges of the others, indicating a uniform LMC ISM, despite the differences in brightness and size between the regions. Figure 6.5 presents a comparison of the spectra from each limb region.

6.3.2 NW Regions

The regions seen in Figure 6.2, which we refer to as the NW inner, W loop, and W knot, present interesting questions about the morphology of this remnant. How are these structures formed and how did they evolve into their current state? It would take complex 3D hydrodynamic modeling to answer these questions fully, but the X-ray spectrum can give us an indication. We shall begin with Sedov model fits to these regions to determine if they are projections of dense shell regions. Table 6.2 summarizes the Sedov fits to each. The W loop region is well defined by a Sedov model, as seen in Figure 6.6. Although the Ne levels are slightly higher than found in Table 6.1, the shock temperature and ionization timescale yield similar calculations of age and explosion energy as found in the section above. The NW inner region is poorly fit by the Sedov table model. Turning to a *vpshock* fit of this region we attempt to investigate any overabundance of Fe. We find in the NW inner region Fe values of 0.45 (0.39–0.53) when the abundances were frozen to 0.4. The W knot does not seem to require any additional Fe. The mystery there is the lower levels of Ne found in the Sedov fits. This is the only region in this target, or our previous work with 0548-70.4, that does not need elevated Ne to improve the fit. This NW region is the brightest section in 0534-69.9, where the blast wave is presumably encountering dense material. Dust destruction in the dense ISM could be affecting our results. The emission and structures in this part of the remnant will require further investigation in the future.

6.3.3 Central Emission

The bright central regions of 0534-69.9 each have indications of elevated levels of Fe emission in the plane shock fitting. We allowed all abundances to be linked and varied, with Fe free from solar ratios. We quote the resulting abundance and Fe values, with their 90% confidence ranges, for the single component fits. The NE center, with 4071 counts, had elemental abundances (excepting Fe) of 0.79 (0.71–0.87) solar, with Fe equal to 2.3 (2.1–8.0), with $\chi^2/\text{DOF} = 125/61$. The S center, with 6063 counts, had elemental abundances of 0.72 (0.45–1.6) with Fe equal to 2.8 (1.9–5.8) with $\chi^2/\text{DOF} = 123/67$. Here we infer even higher ratios of Fe to other metals of 3–4. The SE central region, with 4248 counts, is rather well described by a *vpshock* model. With $\chi^2/\text{DOF} = 67/61$, we find elemental abundances of 0.45 (0.41–0.50) and an Fe value of 1.61 (1.3–4.1). Once again we find a

ratio of Fe to other metals of almost 4. Two component fits will show if any other elements have heightened abundances in this region, or if only excess Fe is present. The final region to check is the NW, containing 3446 counts. Poorly constrained results were found when elemental abundances (excepting Fe) were allowed to float, so we had to freeze them at 0.4 solar. Our resulting Fe abundance was 0.9 (0.81-1.0), this time only slightly more than twice the level of the other elements. For each central region, we find an indication of elevated Fe abundances which we attribute to the presence of ejecta.

Our two component fit results to the bright NE and S central regions are summarized in Table 6.3. We determined the values of Fe, O, Mg, Si, and S, and found no Ne was necessary for either of these central regions (the upper limit of Ne is quoted in Table 6.3). Figure 6.7 depicts the Sedov plus *vpshock* two component fit to the NE center, which closely matches (within 90% confidence ranges) the two component fit with the HE shock model. The S central spectrum can be seen in Figure 6.8, with the HE shock fit result. Our derived reverse-shock speeds are 360 km s^{-1} for the S center of 0534-69.9, and 500 km s^{-1} for its NE center. The other central region two component fit results are found in Table 6.4. As before, no Ne was needed in these fits, but much more O was found. Figure 6.9 is the *vpshock* two component model of the NW center, and the same for the SE center in Figure 6.10. Notice the large difference in normalization between the Sedov limb components between these two figures. In the dim central regions, our method of accounting for the blast wave emission seems to be affected by the variations seen across the limb region. HE model fits were not performed on the regions, so the second emission measure term in Table 6.4 is still in solar units.

Shock speeds and other inferences obtained from plane shock fits are uncertain, because plane shocks most likely oversimplify the ejecta structure. This is the case for Type Ia remnants, where Dwarkadas & Chevalier (1998) find much denser ejecta near the contact discontinuity than at the reverse shock. For estimated explosion energies and swept-up masses listed above, and assuming $1.4M_{\odot}$ of ejecta with an exponential density profile (Dwarkadas & Chevalier, 1998), the reverse shock should have propagated to the center of 0534-69.9 at the age of 7200 yr. This might not have happened yet if a substantial amount of energy has been transferred to cosmic rays at the forward shock, causing us to overestimate the ages.

In 0534-69.9, enhanced oxygen was detected in the NE center (an O/Fe ratio by number of 6.0) and in the S center (2.8). These values are much smaller than expected for a

Type II explosion, so our results indicate that a massive-star progenitor for SNR 0534-69.9 is unlikely. Our results from Table 6.4 for the fainter central regions have much higher values, however. The SE center has an O/Fe ratio of 10.6, while the value is even higher in the NW, at 27.1. Our results now seem to be approaching the O/Fe ratio of ~ 70 for Type II. As we discussed in Chapter 5, such high values could be the result of the plane shock assumption of uniform densities for each element. We see the same residual footprint in Figures 6.7 and 6.8 near the O complex at 0.55-0.65 keV which could be an indication of the cool, dense clumps of O that cause the discrepancy in our O/Fe ratios compared to Type Ia predictions.

Our upper-limit mass calculations yield $0.46M_{\odot}$ in the S center region, with electron density of 0.045 cm^{-3} and a n_e/n_i ratio of 10.2, and $0.20M_{\odot}$ in the NE center, with n_e of 0.06 cm^{-3} and a ratio of 8.9. Volume and ram pressure were calculated in the same fashion as before. (We estimate the radii of the S and NE central regions, again assumed spherical, as $1.2 \times 10^{19} \text{ cm}$ and $9.5 \times 10^{18} \text{ cm}$, respectively.) Since we are looking for an upper limit to H and He mixing, we assume no interactions between the pressures of the central regions. The excesses of blast-wave over reverse-shock ram pressure are $4.6 \times 10^{-10} \text{ dyne cm}^{-2}$ in the NE center of 0534-69.9, and $7.4 \times 10^{-10} \text{ dyne cm}^{-2}$ in its S center. With maximum mixing, our electron density becomes 0.35 cm^{-3} giving a lower limit on ejecta mass (times $f^{1/2}$) of $0.09M_{\odot}$ in the S center. The NE center has a mixed electron density of 0.21 cm^{-3} implying at least $0.12M_{\odot}$ of ejecta.

6.4 Discussion

The X-ray spectrum of LMC SNR 0534-69.9 shows a clearly defined shell that is well described by a Sedov model indicating outer blast wave interaction with the ISM. Within this shell, there exist many structures that we have investigated. We find slightly elevated Fe levels in the NW inner and western ‘knot’ regions, that could be the result of ejecta mixing to the outer regions of the remnant. Moving inwards, the central region of 0534-69.9 has two clearly brightened regions that show heightened elemental abundances. The values of the O/Fe number ratio in those regions can be consistent with a Type Ia origin for this supernova if O is in higher density regions. Ejecta mass estimates favor a Type Ia progenitor. Turning to the dimmer central regions, we also find evidence for SN

ejecta. This mature remnant, that was thought to be in the Sedov phase of evolution, still has significant emission from ejecta after $\sim 10,000$ yr.

Our results are dependent on our spectral models, and the assumptions therein. The assumption of uniform density and full mixing of the heavy elements in the shocked plasma is clearly not the case in supernovae, as discussed in the literature (Dwarkadas & Chevalier, 1998; Badenes et al., 2003). The plane shock models used to account for reverse shock heated ejecta are approximations that represent an important first step in understanding the X-ray emission. As hydrodynamical modeling continues to evolve, so shall our knowledge of the detailed structures that have been discovered in this and other *Chandra* observations of supernova remnants.

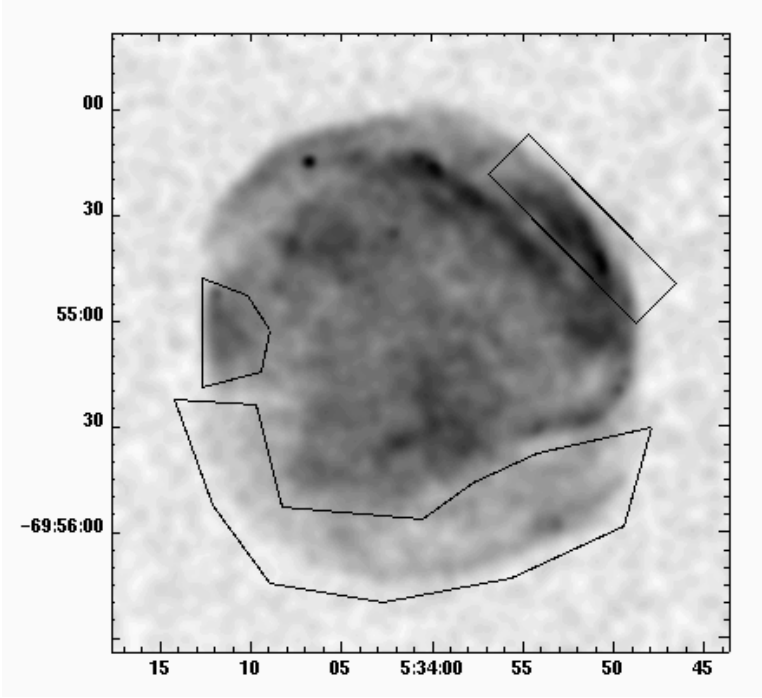


Figure 6.1: Image of 0534-69.9, with NW, S, and E limb regions.

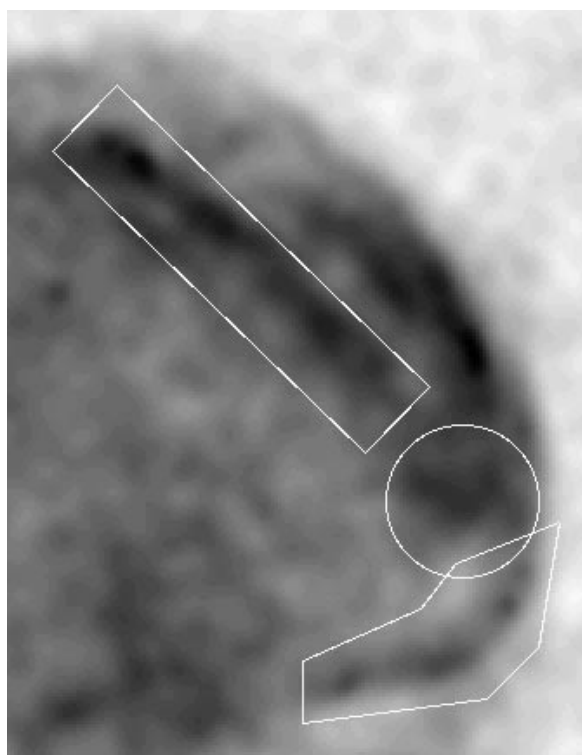


Figure 6.2: NW region of 0534-69.9 behind the limb. Note the detailed morphology.

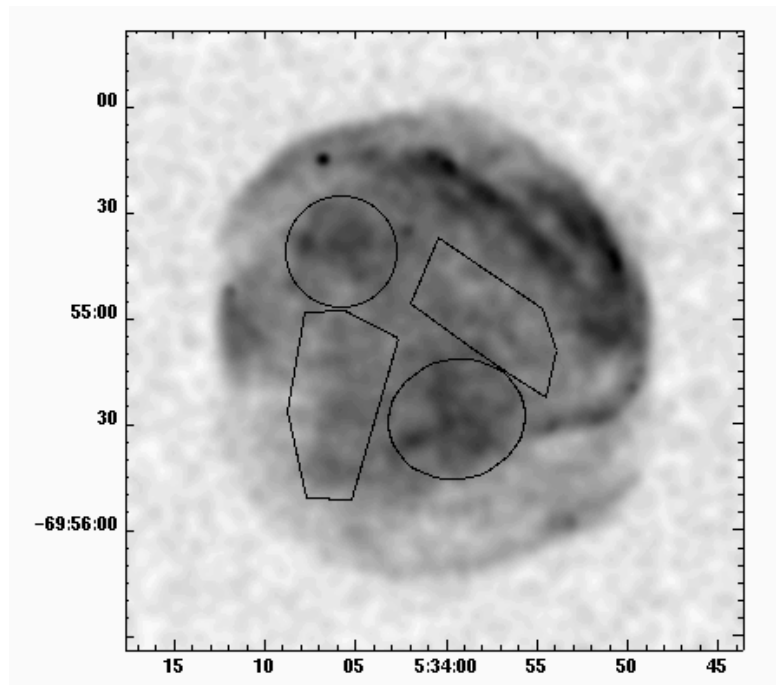


Figure 6.3: Same as Figure 6.1, with NE, S, NW, and SE central regions.

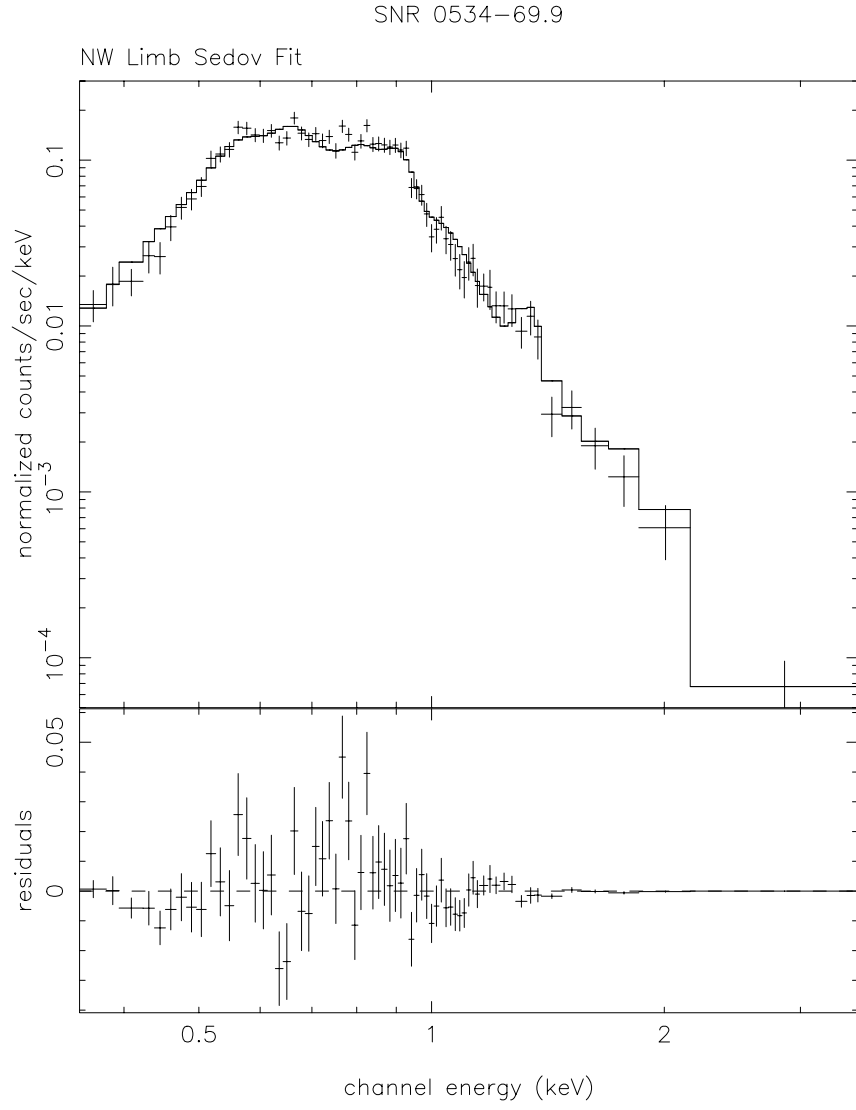


Figure 6.4: Sedov model fit to the NW limb spectrum.

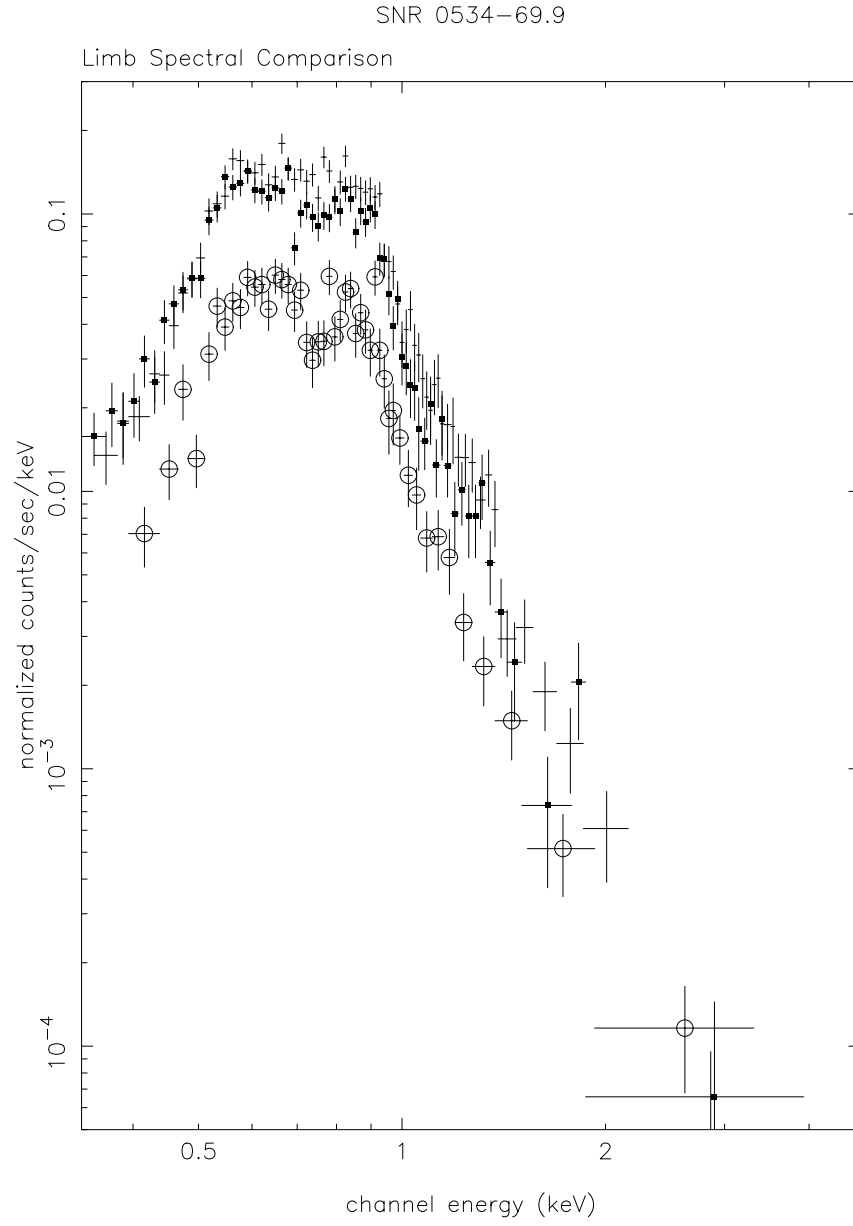


Figure 6.5: Spectral comparison of the NW, S (squares), and E limb (circles) regions.

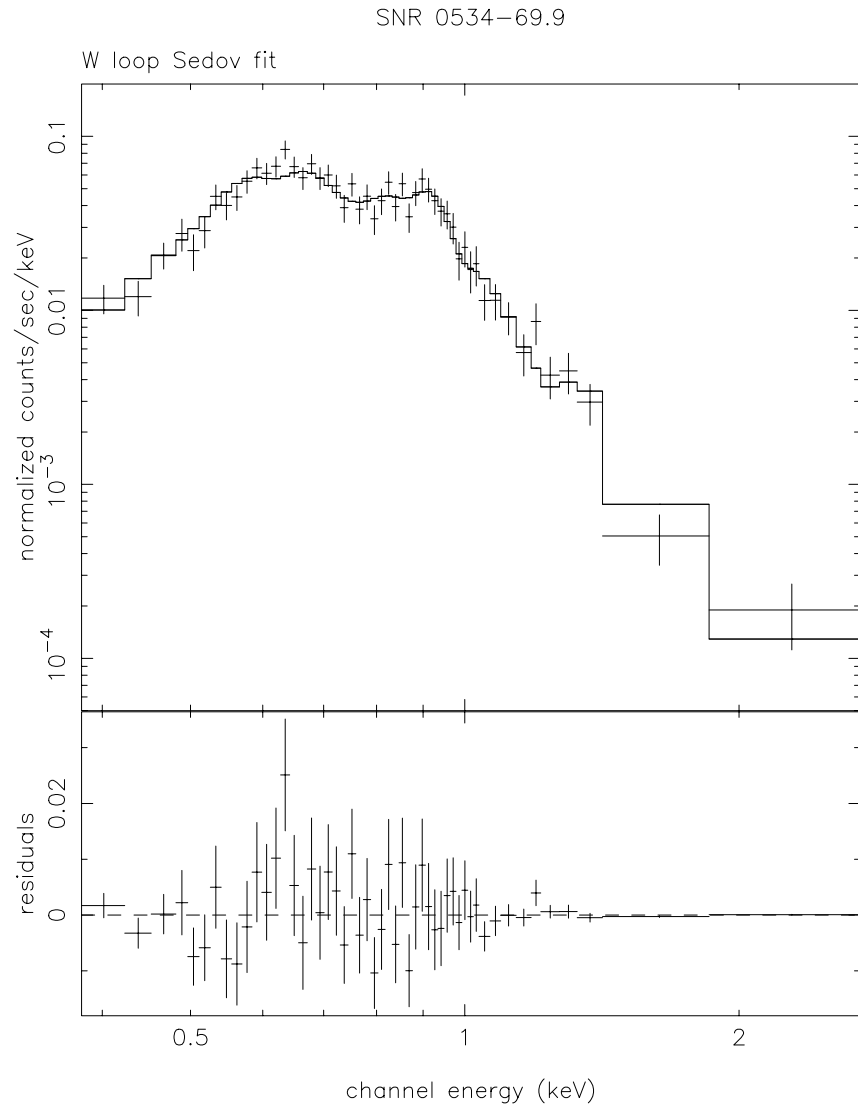


Figure 6.6: Sedov model fit to the W loop region.

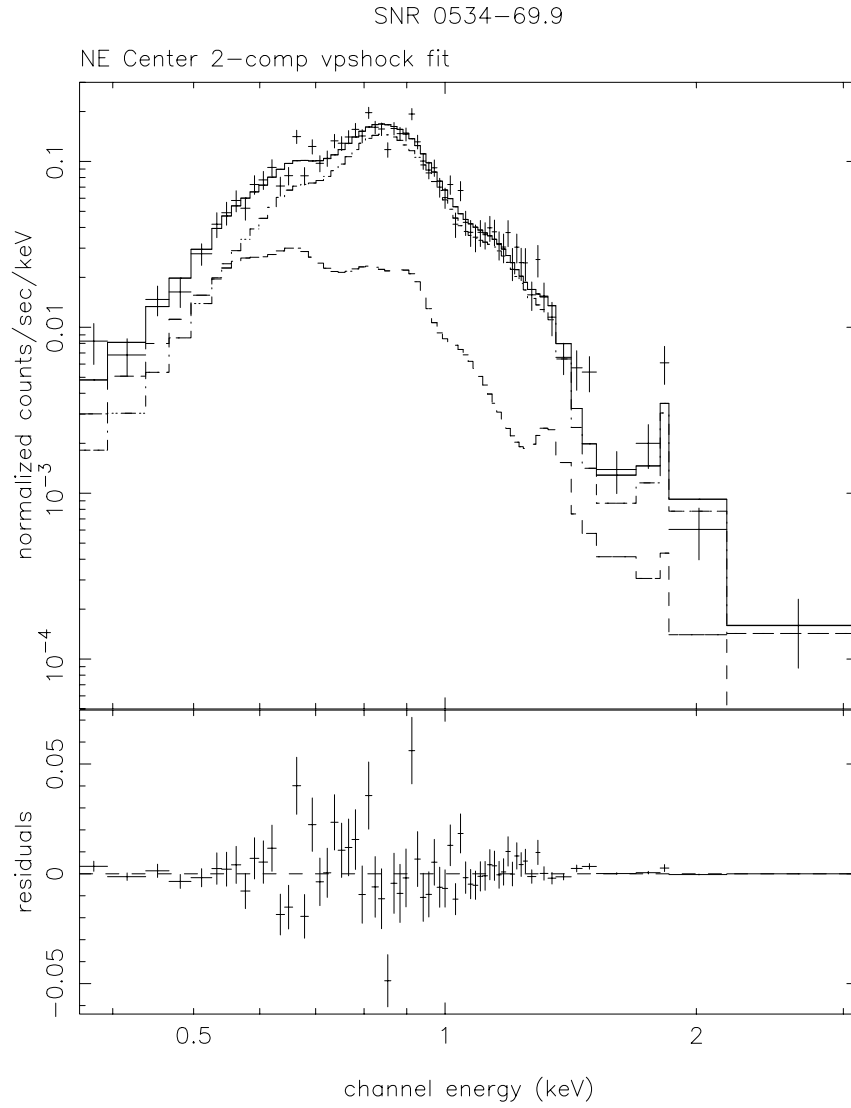


Figure 6.7: Two-component *vpshock* fit to NE center. The curve through the data is the fit, with the two components shown separately as the two lower curves. The lower of the two curves represents the Sedov model component, which matches the limb spectrum in shape, and the upper curve is the *vpshock* model.

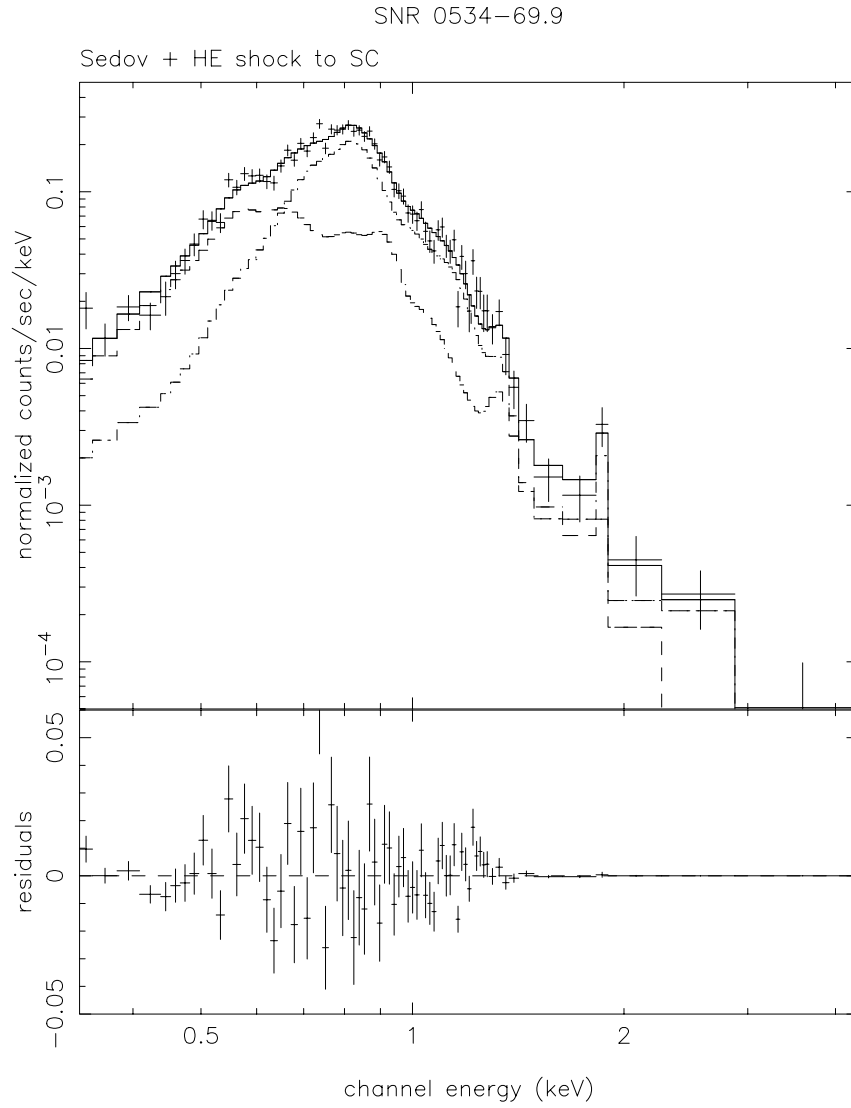


Figure 6.8: Two-component HE shock fit to S center. The curve through the data is the fit, with the two components shown separately as the two lower curves. The lower of the two curves represents the Sedov model component, which matches the limb spectrum in shape, and the upper curve is the HE model.

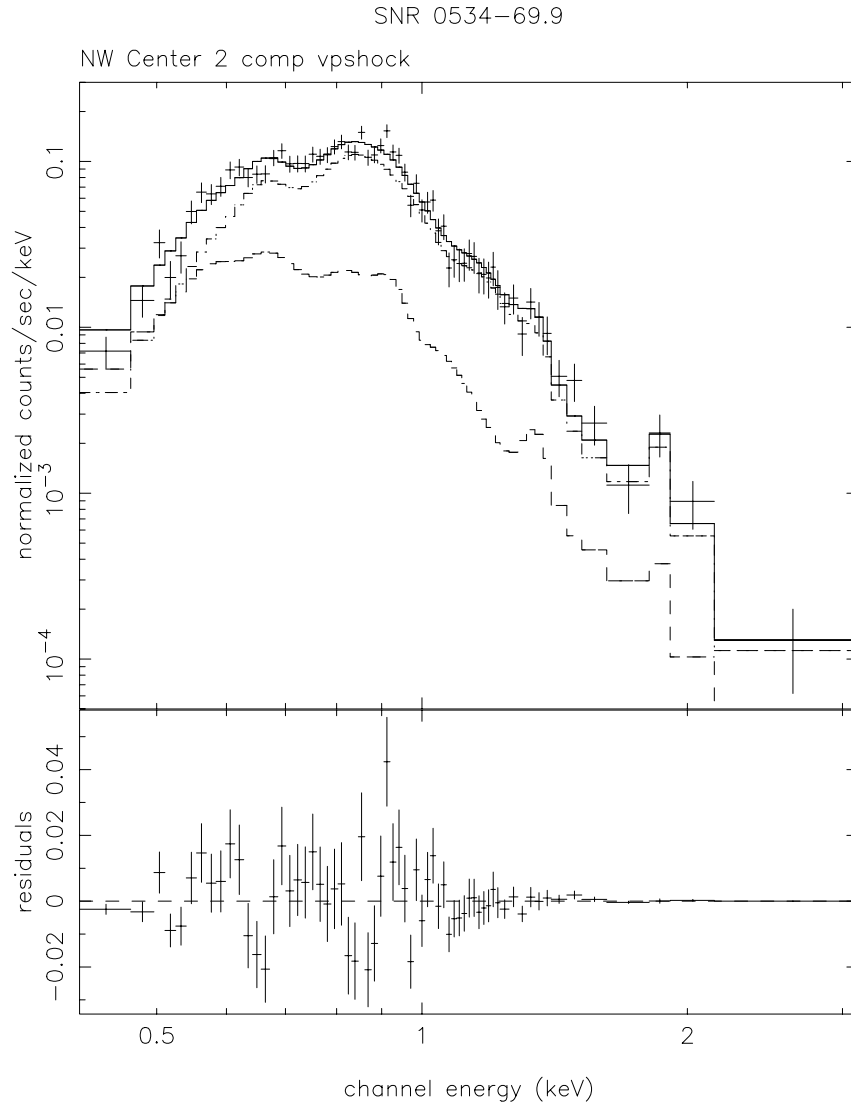


Figure 6.9: Two-component *vpshock* fit to NW central region. The curve through the data is the fit, with the two components shown separately as the two lower curves. The lower of the two curves represents the Sedov model component, which matches the limb spectrum in shape, and the upper curve is the *vpshock* model.

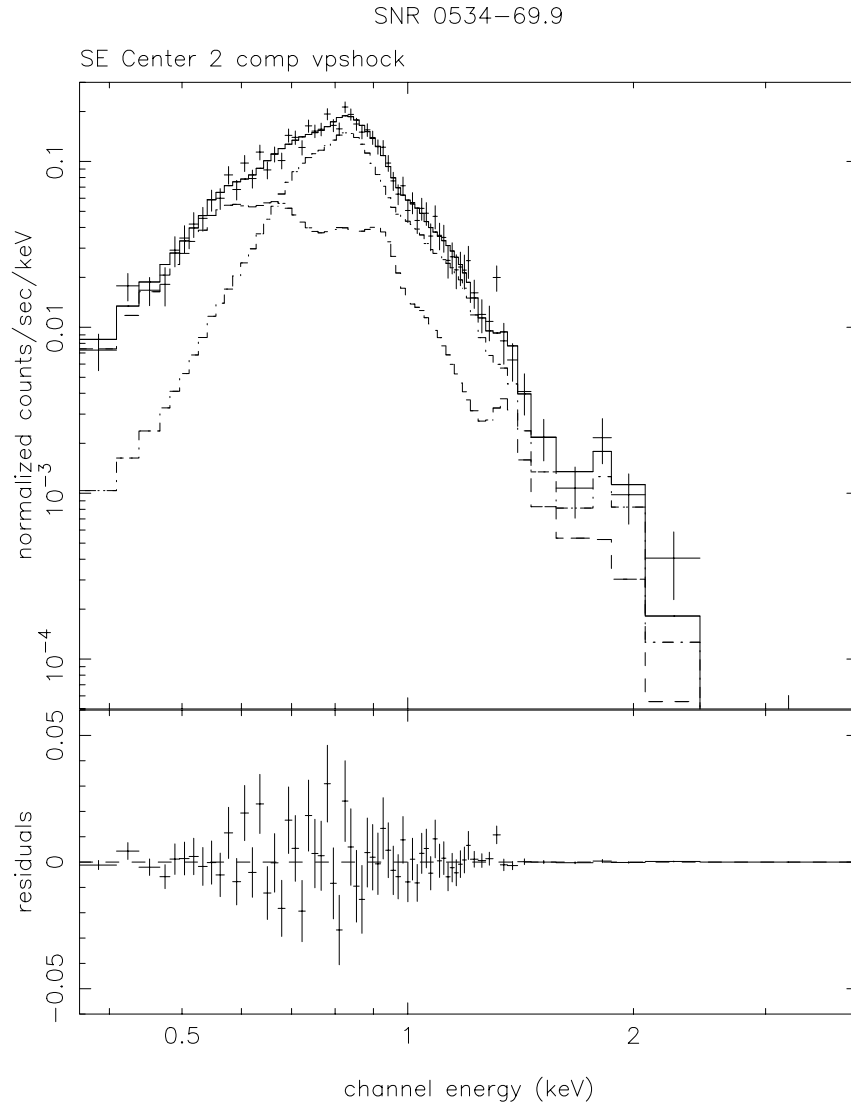


Figure 6.10: Two-component *vpshock* fit to the SE central region. The curve through the data is the fit, with the two components shown separately as the two lower curves. The lower of the two curves represents the Sedov model component, which matches the limb spectrum in shape, and the upper curve is the *vpshock* model.

Table 6.1: Sedov model fits to limb regions

Parameters	NW Limb	S Limb	E Limb
χ^2/DOF	91.5/61	83.3/65	63/41
$N_H/10^{21} \text{ cm}^{-2} \text{ (LMC)}$	$2.6^{+3.1}_{-0.2}$	$1.9^{+1.6}_{-0.7}$	$2.9^{+1.2}_{-0.7}$
Abundances	0.4	0.4	0.4
Ne	$0.59^{+0.08}_{-0.14}$	$0.68^{+0.12}_{-0.12}$	$0.52^{+0.12}_{-0.11}$
$kT_s \text{ (keV)}$	$0.34^{+0.02}_{-0.19}$	$0.33^{+0.04}_{-0.13}$	$0.34^{+0.06}_{-0.10}$
$kT_e \text{ (keV)}$	$0.0^{+0.10}$	$0.0^{+0.13}$	$0.0^{+0.14}$
$\tau/10^{11} \text{ cm}^{-3}\text{s}$	$2.64^{+2.64}_{-1.10}$	$2.43^{+1.71}_{-0.75}$	$1.89^{+2.87}_{-0.39}$
$(EM/4\pi d^2)/10^{10} \text{ cm}^{-5}$	5.30	3.52	2.28

Table 6.2: Sedov model fits to NW regions

Parameters	NW Inner	W Knot	W Loop
χ^2/DOF	179/71	86/49	44/45
$N_H/10^{21} \text{ cm}^{-2}$ (LMC)	3.49	3.70	1.50
Abundances	0.4	0.4	0.4
Ne	0.43	0.38	0.79
kT_s (keV)	0.34	0.34	0.34
kT_e (keV)	0.0	0.0	0.0
$\tau/10^{11} \text{ cm}^{-3}\text{s}$	3.68	1.89	1.89
$(EM/4\pi d^2)/10^{10} \text{ cm}^{-5}$	5.30	3.52	3.11

Table 6.3: Two-component fits to NE and S central regions

Parameter	NEC+NWL		SC+SL		Solar Ratios
	<i>vpshock</i>	HE shock	<i>vpshock</i>	HE shock	
χ^2/DOF	116/58	114/58	111/64	109/64	
$(EM_L/4\pi d^2)/10^{10} \text{ cm}^{-5}$	$1.01^{+0.4}_{-0.4}$	1.04	$2.00^{+0.5}_{-0.4}$	2.12	
$kT_e \text{ (keV)} / v_s \text{ (km s}^{-1}\text{)}$	$0.92^{+0.48}_{-0.30}$	501	$0.61^{+0.09}_{-0.10}$	356	
O/Fe	$5.98^{+2.17}_{-2.07}$	5.96	$2.75^{+3.07}_{-1.75}$	2.79	18.2
Ne/Fe	$0^{+0.58}$	0	$0^{+0.32}$	0	2.63
Mg/Fe	$0.16^{+0.10}_{-0.08}$	0.15	$0.18^{+0.09}_{-0.08}$	0.18	0.81
Si/Fe	$0.23^{+0.08}_{-0.10}$	0.22	$0.22^{+0.14}_{-0.14}$	0.22	0.76
S/Fe	$0.18^{+0.26}_{-0.18}$	0.17	$0.51^{+0.46}_{-0.46}$	0.50	0.35
$\tau/10^{11} \text{ cm}^{-3}\text{s}$	$0.97^{+4.03}_{-0.27}$	1.49	$1.46^{+1.64}_{-0.47}$	2.71	
$(EM_{EJ}/4\pi d^2)/10^6 \text{ cm}^{-5}$	$2.84^{+0.24}_{-0.68}$	5.20	$7.00^{+0.67}_{-0.71}$	4.84	

Table 6.4: Two-component fits to SE and NW central regions

Parameter	SEC+SL	NWC+NEL	Solar
χ^2/DOF	<i>vpshock</i> 61/58	<i>vpshock</i> 83/56	Ratios
$(EM_L/4\pi d^2)/10^{10} \text{ cm}^{-5}$	1.54	1.16	-
$kT_e \text{ (keV)} / v_s \text{ (km s}^{-1}\text{)}$	0.53	0.68	-
O/Fe	10.6	27.1	18.2
Ne/Fe	0.13	0.0	2.63
Mg/Fe	0.15	0.22	0.81
Si/Fe	0.23	0.18	0.76
S/Fe	0.08	0.0	0.35
$\tau/10^{11} \text{ cm}^{-3}\text{s}$	21.5	4.62	-
$(EM_{EJ}/4\pi d^2)/10^{10} \text{ cm}^{-5}$	1.55	1.4	

Chapter 7

On the X-ray Spectrum of LMC

SNR 0453-68.5

7.1 Introduction

As we study supernova remnants, the question of the explosion type is important in understanding the evolution of the ISM. In the previous two chapters we have examined the X-ray spectra for signs of ejecta emission that would indicate their chemical composition. We have determined that both SNRs 0548-70.4 and 0534-69.9 were the result of Type Ia explosions. The final SNR in our sample, 0453-68.5, has a much clearer indication of its progenitor. Examining our *Chandra* observation in the hard band (2.0-10.0 keV), we find the presence of a small, elliptical source in the center of the remnant. We identify this as a pulsar wind nebula (PWN) due to its hard, nonthermal spectrum which is much harder (flatter) than thermal emission from the shell. A pulsar is only created in the gravitational collapse of a massive star, a core-collapse explosion. PWN are sources of bright emission from the bubble of relativistic electrons and magnetic fields around the pulsar, the most famous of which is the PWN in the Crab Nebula (created in SN 1054). Synchrotron emission is important as the electrons are caught in the magnetic field surrounding the pulsar. Some recent examinations of PWN have been done in Gaensler et al. (2001); Petre et al. (2002);

Gonzalez & Safi-Harb (2003). We will attempt to isolate the PWN spectrum from the shell emission in 0453-68.5. Following our X-ray detection of the nebula, we took part in a radio observation with the ATCA (Australian Telescope Compact Array). We present a summary of the X-ray spectrum of 0453-68.5, and a comparison the radio and X-ray images.

7.2 The Observations

SNR 0453-68.5 was observed with *Chandra* on 6 November 2001, for 40 ksec. *Chandra* analysis software CIAO version 2.2.1 was used to process the data. Filtering was performed for good time interval, bad grades, observation specific bad pixels, and high background times, as described in Appendix A. The largest remnant in our sample, 0453-68.5 has an angular extent of ~ 142 arcsec which corresponds to a radius of 17.2 pc, assuming a distance to the LMC of 50 kpc. It fell entirely within the back-illuminated S3 chip in the *Chandra* ASIC-S array. Figure 7.1 presents a smoothed image of the remnant with the regions of limb spectra extraction: E limb, NW limb, and S limb. Also included in the figure are the ‘spokes’ in the NW quadrant, which we will refer to as the NW horizontal and NW vertical. Spectra were extracted from each region, and the corresponding response file created and corrected for hydrocarbon contamination. The same source-free region of the chip was used for background subtraction in each extracted spectrum. Figure 7.2 presents a comparison of the soft band (0.3–2.0 keV) and hard band (2.0–8.0 keV) images of 0453-68.5. The PWN can clearly be seen in the hard band image.

Radio observations took place over three dates, 24 July, 27 July, and 26 August 2002, for a total of 39.0 hours with ATCA. ATCA is an array of six 22-m radio antennae; 5 lie along a 3 km east-west track, with the sixth 3 km further west. With wavelength bands of 20 cm, 13 cm, 6 cm, and 3 cm, ATCA can cover a wide range of frequencies. Our colleague Bryan Gaensler handled these observations, and kindly provided the radio images seen here. Figure 7.3 is a set of spectral tomography maps of 0453-68.5 between the 20 cm and 13 cm observations. A spectral tomography map takes a possible value for the spectral index α_t , and compares emission from two wavelengths. The values on the maps correspond to $S_{20} - S_{13}(20/13)^{-\alpha_t}$, where spectral indices α are defined by $S_\nu \propto \nu^\alpha$. The panel where a given feature flips from positive to negative indicates the spectral index of that feature ($\alpha = \alpha_t$). The core flips at $\alpha = -0.1$, with the shell flipping at $\alpha = -0.45$. The most

interesting features in the radio image are the core emission from the PWN (much larger than the X-ray emission) and the SW limb region (this feature is much weaker in X-rays compared to the rest of the shell).

7.3 Image Analysis

The CIAO software package includes the spectral fitting program *Sherpa* that is analogous to the XSPEC routine we have been using. One feature that makes *Sherpa* different is its ability to perform 2D fitting of an image. We make use of the method described in Murray et al. (2002) to attempt to isolate the pulsar within the PWN. We have far fewer counts to work with, but we can determine a lower limit on the number of counts in the pulsar. We used a three component model. One component was an elliptical Gaussian (fitted values: eccentricity of 0.32, amplitude 2.8, and FWHM 28 pixels), to define the shape of the PWN, and another component was a constant factor to account for background emission. The final component used the point-spread function (PSF) of the *Chandra* telescope appropriate for the PWN region to determine the level of emission from a possible point source. Figure 7.4 shows the amount of noise in the dataset, but a clear peak is noticeable in the center. However, the PSF component determined 49 counts from the PWN. Unfortunately because of the quality of our data, 90 % confidence ranges were ± 26 , with 3σ (99.7 % confidence) at ± 49 . The data hint at the presence of a point source, but are also consistent with none.

7.4 Results

7.4.1 Limb Emission

The bright limb regions of 0453-68.5 are each well described by Sedov table models, as summarized in Table 7.1. We froze elemental abundances to 0.4 solar, but allowed Ne to float as described in previous chapters. The E limb, dimmest of our limbs, contained 5126 counts between 0.3 and 5.0 keV. The S limb had 7976 counts, while the NW limb had 11,798 counts. Figure 7.5 shows the Sedov model fit to the NW limb, and a comparison of the three limbs can be seen in Figure 7.6. The NW limb had a slightly lower shock

temperature, 0.24 keV, than the E and S limbs which both had $T_s = 0.34$ keV. The electron temperature was essentially zero for each fit, with upper limits quoted in the table. Each model fit used a two component absorption model. One component was frozen at 7×10^{20} cm $^{-2}$ for Galactic absorption (Dickey & Lockman, 1990), and the other component was fit assuming LMC abundances. The second component only is quoted in Table 7.1.

The assumption of Sedov dynamics allows the calculation of several important quantities in a SNR. Our shock temperatures give values of shock speed, $v_s = 9.1 \times 10^7 \sqrt{kT_s}$, as ~ 530 km s $^{-1}$ for the S and E limb ($v_s \sim 454$ km s $^{-1}$ in the NW limb). Shock age is then determined by $t = 2R/5v_s = 12,700$ yr. By estimating the limb regions as caps of spherical shells, we are able to determine the mass swept up in each region. We find $24.4M_\odot$ of ISM in the E limb, with $\sim 50M_\odot$ in both the NW and S limbs. Applying a Sedov model to the entire SNR (minus the PWN), we can calculate the swept up mass for the entire remnant to be $\sim 282M_\odot$ for an upstream ion density $n_o = 0.39$ cm $^{-3}$ determined from the equation after equation 10 in Hamilton et al. (1983). This much larger value includes contributions from the entire SNR shell, not just the limb regions we have defined above. For each of these fits, we determined an explosion energy of 5×10^{50} ergs. 0453-68.5 seems to be yet another low energy SNR.

Figure 7.7 compares the radio observation to the X-rays. The central regions compare the radio emission to the X-ray PWN. Notice how the X-ray PWN is off center from the much larger radio region. The most interesting feature is the radio bright SW corner of the SNR. We have used the region in the SW corner of the X-ray image in Figure 7.7 to extract the X-ray spectrum. It was drawn to avoid overlap with the S limb region, and include the diffuse emission beyond that limb. Figure 7.8 compares that spectrum to the S limb to look for any interesting features. There are very counts in the SW corner so fitting was not performed. The source of the bright radio emission could be from radiative shocks, but then that region should be bright in optical which is not the case.

7.4.2 Central Features

The NW horizontal and vertical features have no evidence for ejecta, and are most likely projections of the outer limbs. Table 7.2 summarizes the Sedov fit results for these regions. The horizontal part has a higher shock temperature than the limb regions, but the NW vertical agrees with the limb results. The vertical region does not seem to require the

slightly elevated Ne abundances observed in all the limbs. We investigated these regions with *vpshock* to detect any elevated Fe level, but none was needed. The emission in these regions could be the result of the blast wave encountering a dense cloud in the NW.

7.4.3 PWN Spectrum

We have attempted several methods to isolate the spectrum of the PWN from the shell emission. First we tried to define an annular region around the PWN to be our background, but that background subtraction overwhelmed the emission. The answer was found examining Figure 7.2. By cutting off the spectrum below 2 keV, we eliminate most of the shell spectrum, and are left with the hard spectrum of the PWN (once the source-free background spectrum is subtracted). The flat spectrum of the PWN should be well described by a power law model with a spectral index α . We can compare this to the result of the radio observation. Figure 7.9 is the best fit result of the PWN spectrum from 1.5 – 6.0 keV. We find that $\alpha = -0.9$ with 90% confidence range of -0.50 to -1.3. Another method was to use a two component model, Sedov plus power law, to determine the spectral index over the entire energy range. The result of that fit can be seen in Figure 7.10, where the power law model determines $\alpha = -0.74$ with 90% confidence range of -0.58 to -0.89. These values are typical for X-ray PWNs (Petre et al., 2002).

7.5 Discussion

The most striking result of this study is the discovery of a very old PWN within the interior of SNR 0453-68.5, definitively classifying it as core collapse supernova progenitor. As in the previous observations, we find limb emission in 0453-68.5 that is well described by a Sedov model, but this time without any detection of ejecta in the interior. 0453-68.5 is the largest, and presumably oldest, remnant in our survey so it is not surprising that we discovered no evidence for ejecta. It is most likely that the reverse shock has traveled through the interior and compressed the PWN. This could be the reason why the PWN is off center from the radio image. The radio signal comes from particles injected prior to the compression of the reverse shock. The X-rays emission is from recently injected material. The effects of the reverse shock on PWNs are not well understood, but such a collision

would be very interesting to observe. This PWN may be one of the oldest known.

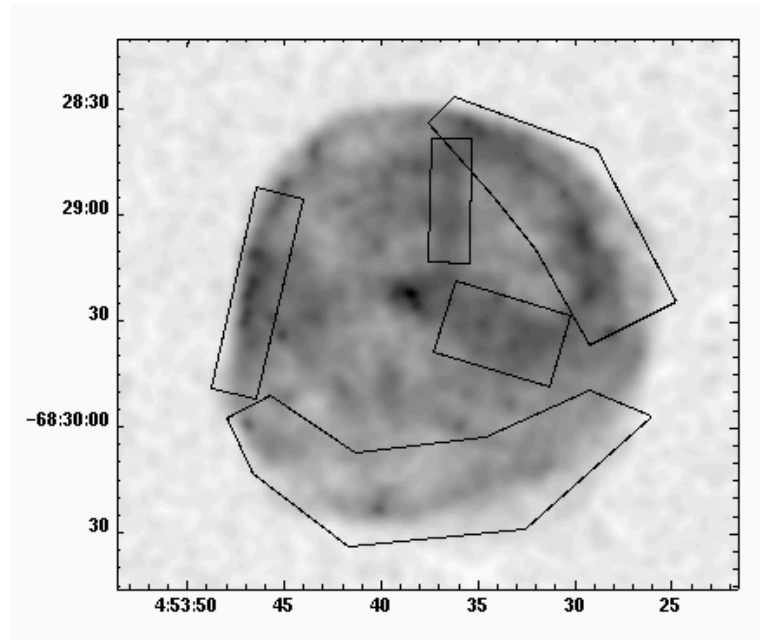


Figure 7.1: Smoothed image of 0453-68.5, with regions drawn for the E, NW, and S limbs. The interior regions represent the NW horizontal and NW vertical components.

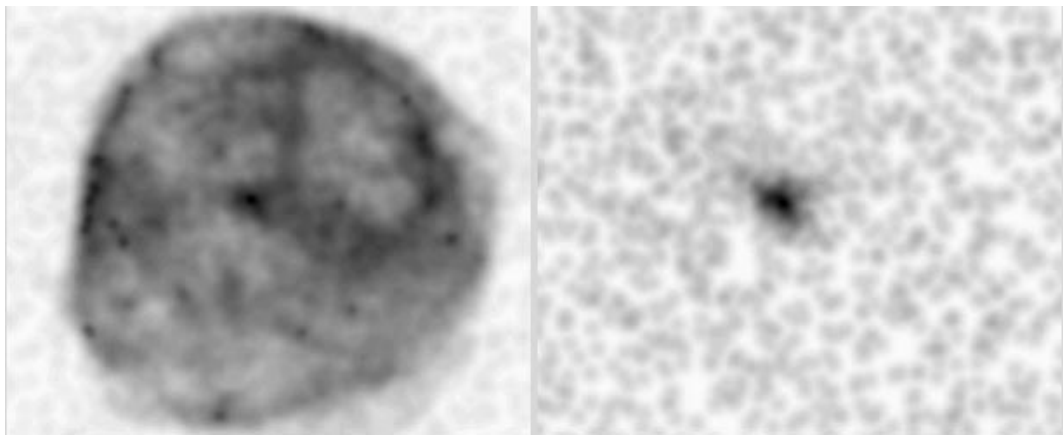


Figure 7.2: On the left, the soft band (0.3-2.0 keV) smoothed image of 0453-68.5, compared to the hard band (2.0-8.0 keV) on the right.

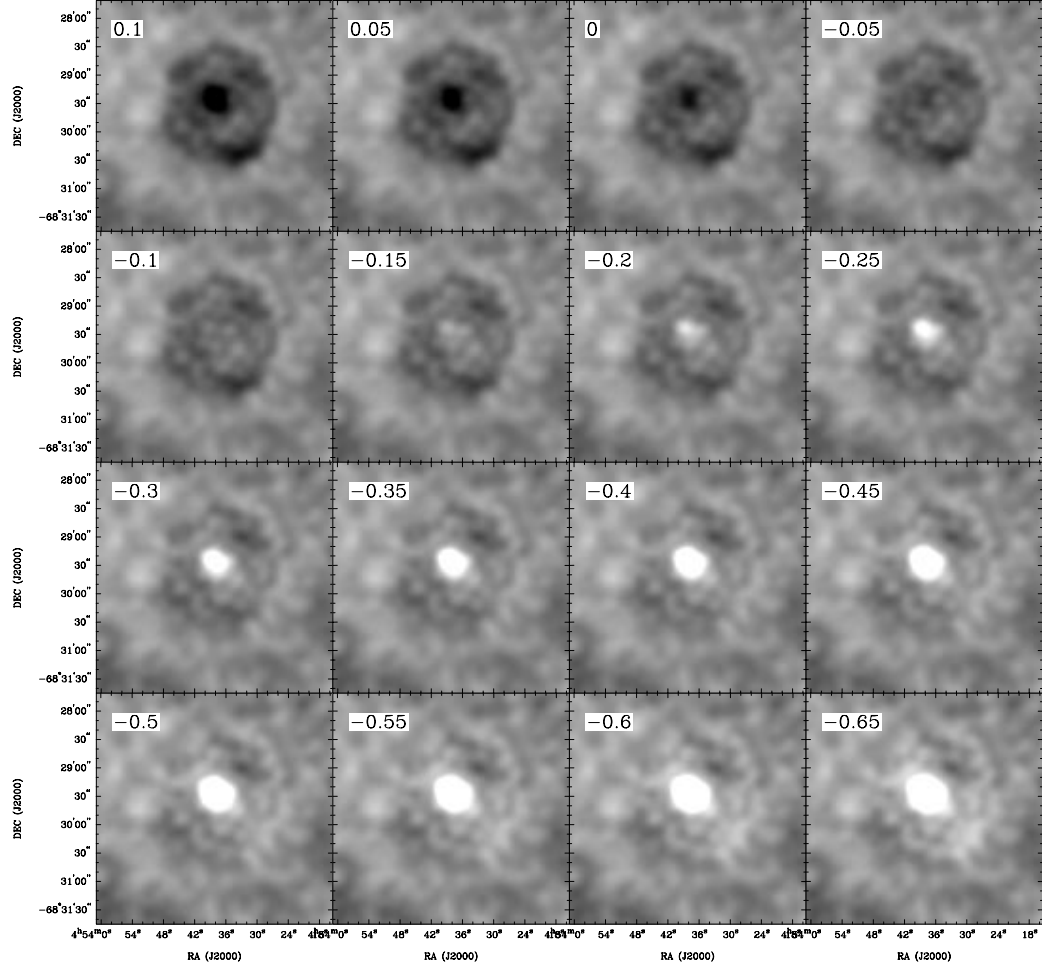


Figure 7.3: Spectral tomography maps between 20 cm and 13 cm radio emission.

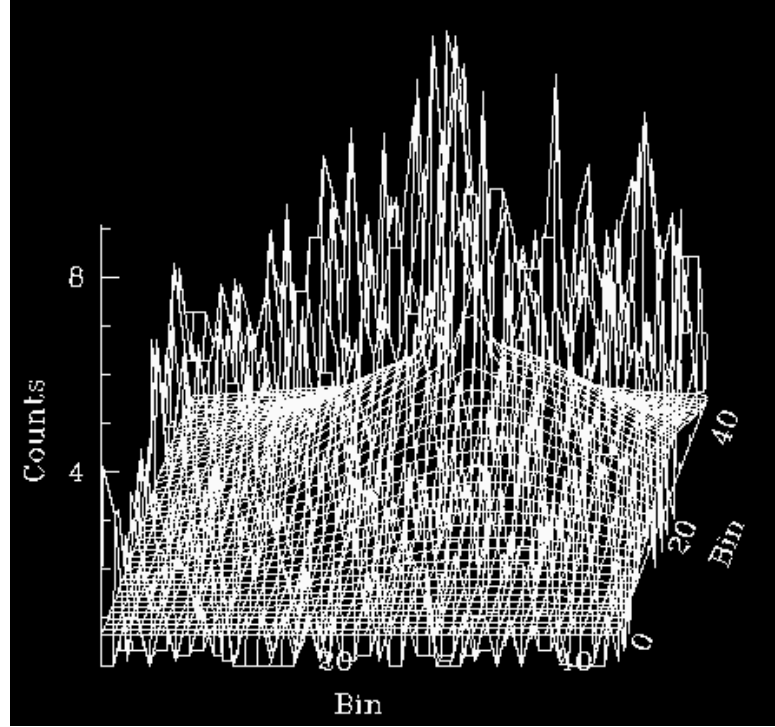


Figure 7.4: 2D fit results to the PWN image.

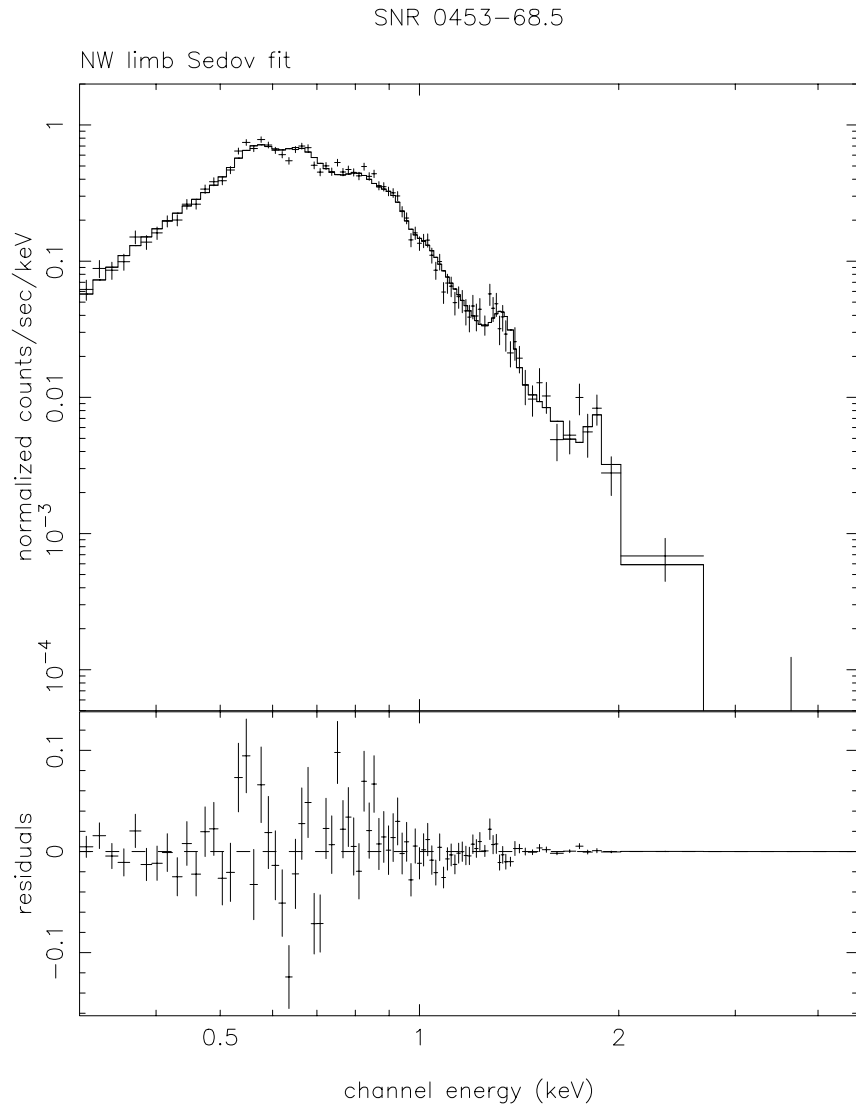


Figure 7.5: Sedov model fit to the NW limb

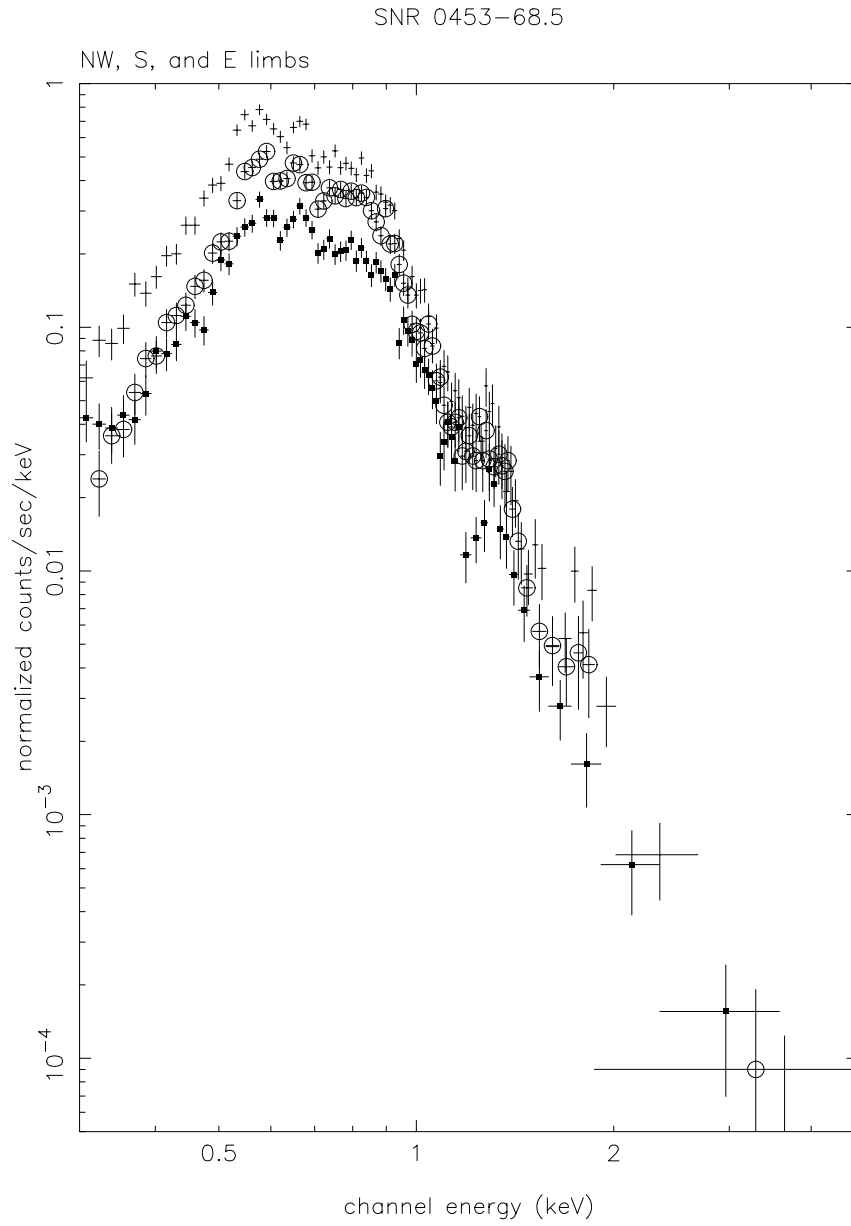


Figure 7.6: Spectra of the three limb regions: NW, E (boxes), and S (circles).

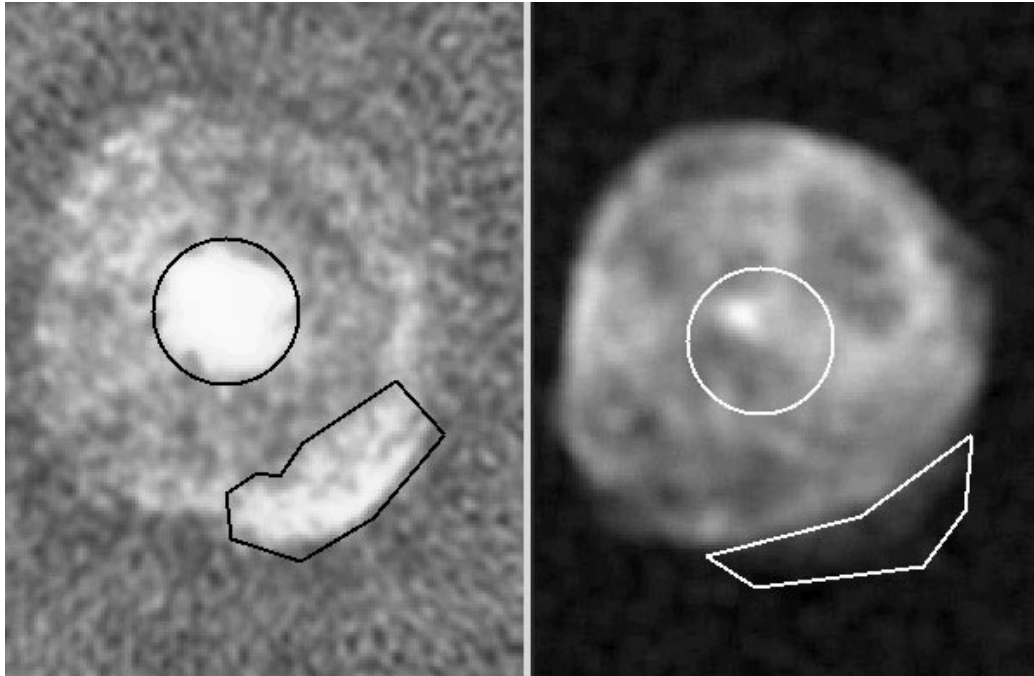


Figure 7.7: Comparison of radio (left) and X-ray (right) emission.

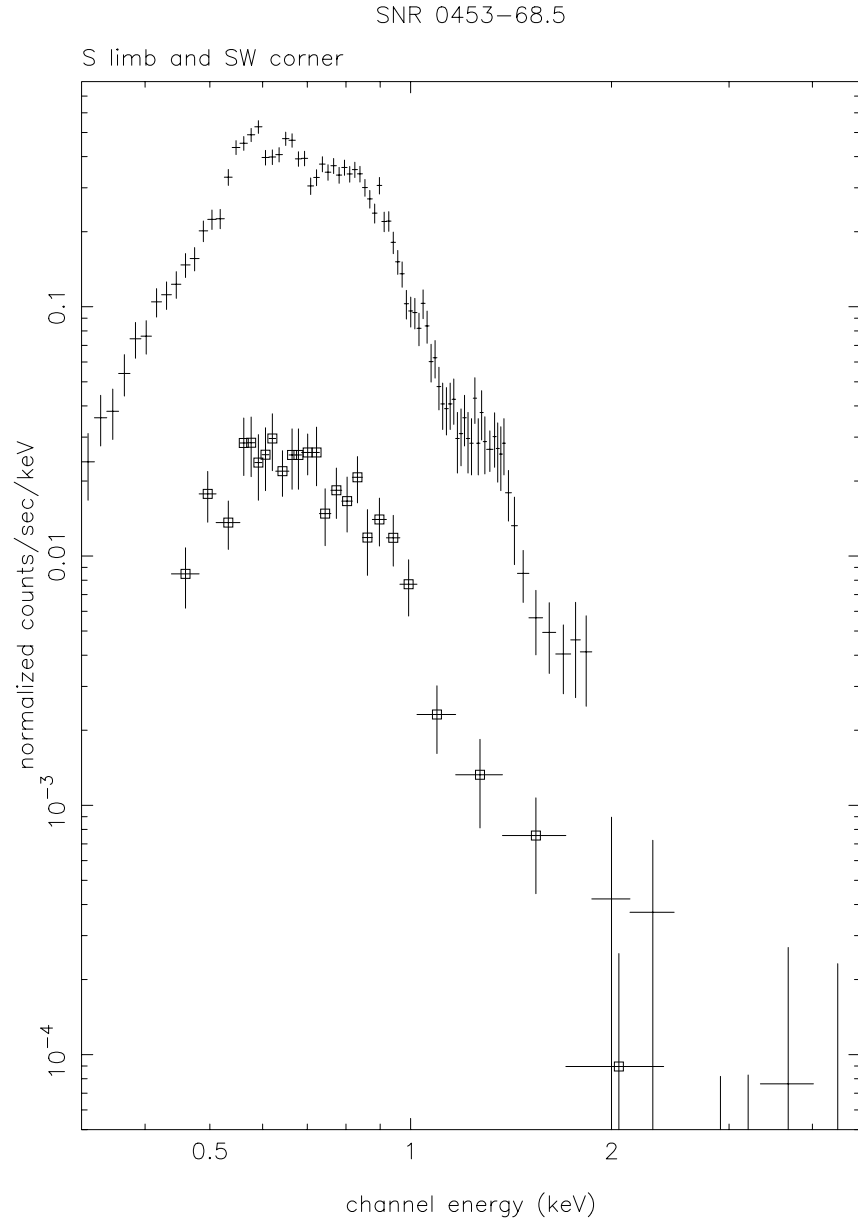


Figure 7.8: Spectral comparison of the S limb and the SW region (boxes).

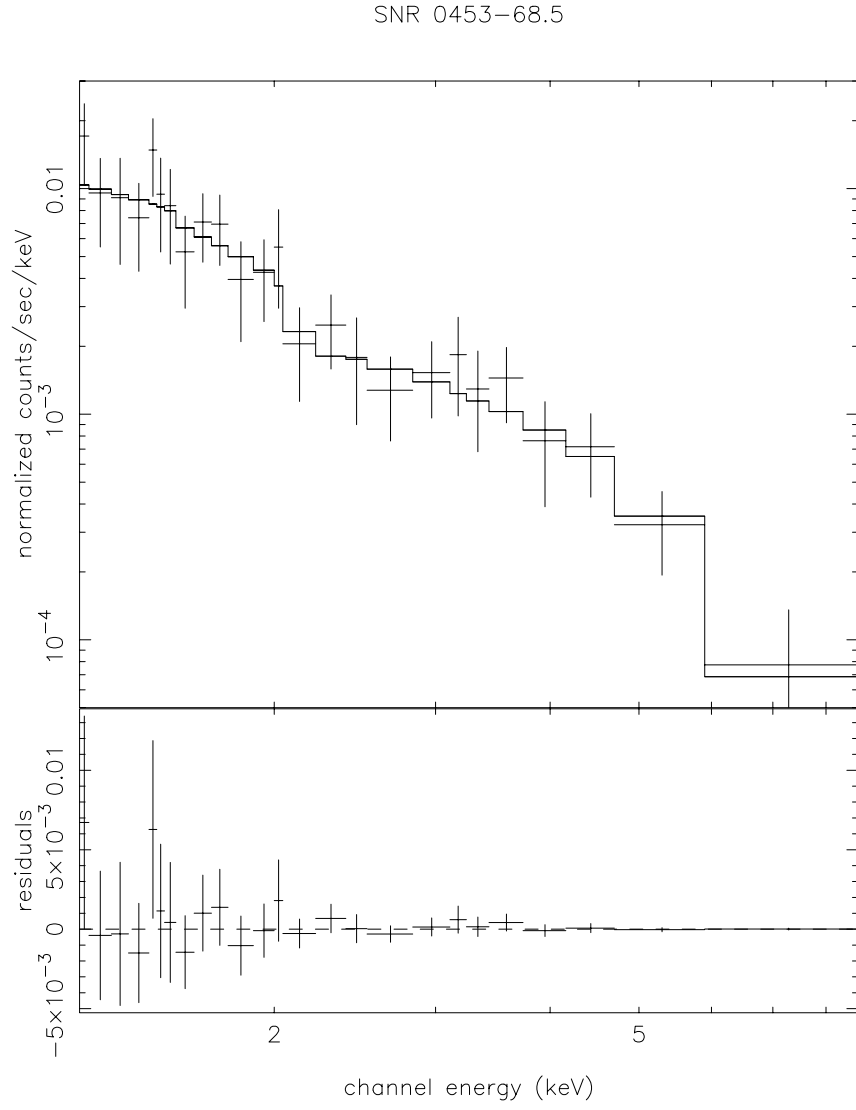


Figure 7.9: Spectrum of 0453-68.5 above 1.2 keV, fit with a power law.

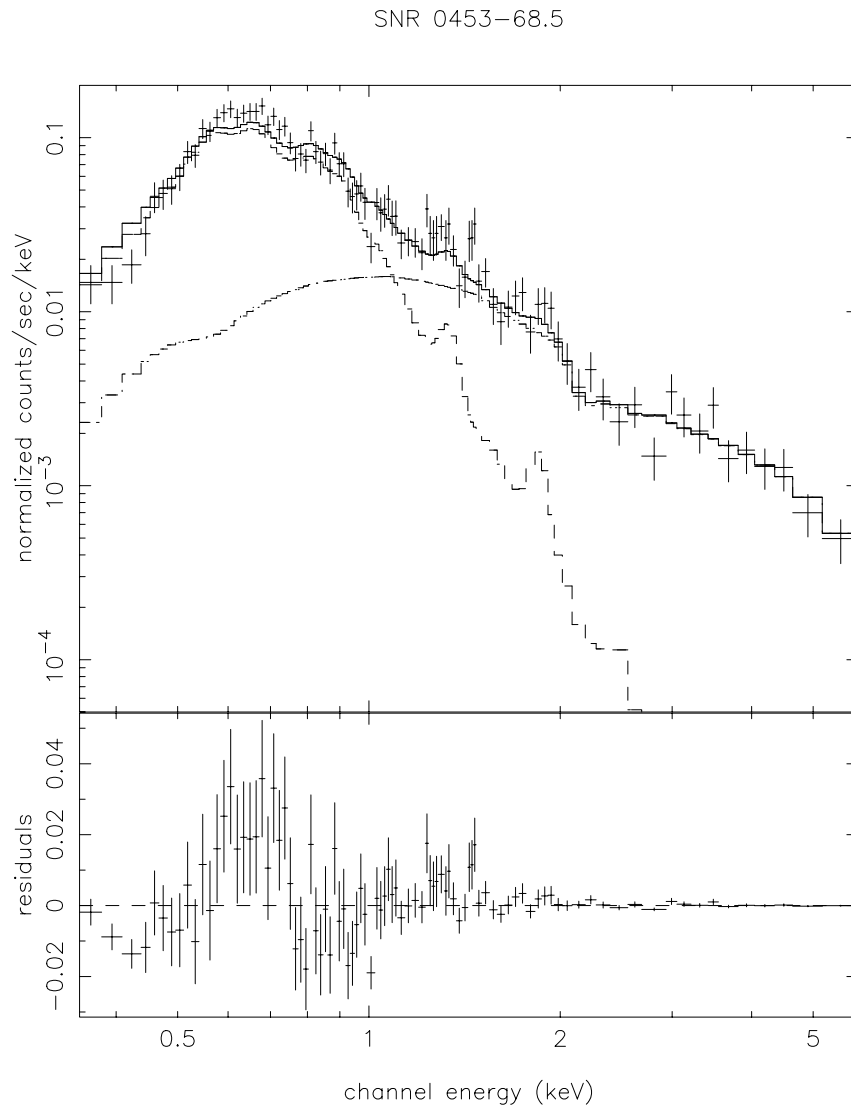


Figure 7.10: Two component fit to entire SNR. The curve through the data is the fit, with the two components shown separately as the two lower curves. The flat curve is the power law component, and the other represents the Sedov model component.

Table 7.1: Sedov model fits to limbs

Parameters	NW	E	S
χ^2/DOF	131/82	82/67	112/79
$N_H/10^{21} \text{ cm}^{-2}$ (LMC)	$0.8^{+0.4}_{-0.2}$	$0.4^{+0.5}_{-0.3}$	$1.9^{+0.6}_{-0.5}$
Abundances	0.4	0.4	0.4
Ne	$0.43^{+0.05}_{-0.05}$	$0.52^{+0.09}_{-0.08}$	$0.43^{+0.05}_{-0.06}$
kT_s (keV)	$0.24^{+0.02}_{-0.02}$	$0.34^{+0.01}_{-0.09}$	$0.34^{+0.01}_{-0.09}$
kT_e (keV)	$0.0^{+0.11}$	$0.0^{+0.19}$	$0.0^{+0.11}$
$\tau/10^{11} \text{ cm}^{-3}\text{s}$	$4.79^{+2.01}_{-1.29}$	$4.55^{+1.65}_{-0.95}$	$2.64^{+0.76}_{-0.59}$
$(EM/4\pi d^2)/10^{10} \text{ cm}^{-5}$	19.5	4.76	15.3

Table 7.2: Sedov model fit summary of NW features

Parameter	Horiz	Vert
χ^2/DOF	125/70	83/55
$N_H/10^{21} \text{ cm}^{-2}$ (LMC)	$1.6^{+0.4}_{-0.5}$	$1.2^{+0.8}_{-0.4}$
Abund	0.4	0.4
Ne	$0.44^{+0.07}_{-0.07}$	$0.28^{+0.08}_{-0.07}$
kT_s (keV)	$0.34^{+0.01}_{-0.04}$	$0.34^{+0.03}_{-0.08}$
kT_e (keV)	$0.0^{+0.12}$	$0.0^{+0.14}$
$\tau/10^{11} \text{ cm}^{-3}\text{s}$	$3.91^{+1.26}_{-0.66}$	$3.43^{+0.77}_{-1.13}$
$(EM/4\pi d^2)/10^{10} \text{ cm}^{-5}$	5.4	4.2

Chapter 8

Conclusions and Future Work

8.1 Summary

Supernova explosions are both the end and the beginning of the stellar lifecycle. Large, massive stars (Type II, Ib/c) and old white dwarf stars in binary pairs (Type Ia) explode, sending shock waves into the ISM, and a reverse shock into their ejecta. The hot material observed in X-rays expands, eventually cooling down and condensing to restart the process of star (and planetary) formation. X-ray observations can help us understand details in this important step of stellar evolution. A new tool has been developed for X-ray spectral analysis in supernova remnants. The HE shock model introduced in this work includes detailed nonequilibrium ionization calculations of the electron temperature in the shock heated plasmas, ejecta and ISM, that are responsible for X-ray emission. These calculations include contributions of the heavy elements to the electron density not included in the current generation of models, where hydrogen and helium are considered the sole source of electrons. SN ejecta have elevated levels of heavy elements from the stellar fusion processes that must be accounted for in spectral analysis. Chapter 3 reviewed the testing and development of this tool for use with the XSPEC spectral analysis software.

The HE shock model is an important refinement in existing theory, especially in light of detailed observations now being performed by the *Chandra* X-ray satellite. This work presents the results on observations of LMC SNRs 0548-70.4, 0534-69.9, and 0453-

68.5. Our investigations began with the outer limb regions of each, where we found spectra consistent with Sedov phase emission from shock heated ISM. When we turned to the central regions, we made some interesting discoveries. In the first two targets, we found evidence for ejecta emission, and used our new code to examine the elemental abundances of the ejecta. We believe our results indicate that 0548-70.4 and 0534-69.9 were created in Type Ia supernova explosions. Our third target contained no detectable heavy element ejecta, but we did discover a new pulsar wind nebula that proves SNR 0453-68.5 came from the gravitational collapse of a massive star. We have also examined data, from the ASCA satellite, for a sample of 11 supernova remnants to examine any non-thermal emission and placed an upper limit on electron acceleration.

8.2 The Future

With dozens of supernova remnants present in the Large and Small Magellanic Clouds, our methods can be applied to a growing sample of observed objects. This is an excellent sample due to the known distances and little extinction to the MCs. We have continued to request new targets with this purpose in mind, receiving new observations of SMC SNR 0049-73.6 (with *Chandra*) and SNR DEM L238 (joint *Chandra*-XMM target) in the LMC. An image of the recently completed 0049-73.6 observation (OBSID 3907) is in Figure 8.1. The central emission is quite striking, much brighter than the limbs. Nearby SNRs 0047-73.5 and 0050-72.8 are also present in that observation (we hope to find LMC SNR DEM L249 off-axis in our DEM L238 observation scheduled for July 2003). We have also submitted recent proposals for *Chandra* observations of LMC SNRs 0520-69.4, 0528-69.2, N11L, and the brightest LMC SNR N132D. The first three targets are not well studied, so they offer a chance to make the type of discoveries we have made with our current sample. N132D is a young, ejecta-dominated remnant where the HE shock model should be very useful in understanding the ejecta emission.

New observations are always important in astronomy, but theoretical advances must also be developed. The HE shock model is a good tool, but it can be improved upon. The obvious extension would be to generalize our calculations to spherical shocks. This would provide a correction factor to existing Sedov models that could include heavy element contributions in ISM plasma where H and He dominate. Plane shocks are still excellent

approximations, but they are limited by the assumption of fully mixed heavy elements with a uniform density. There is growing evidence that ejecta exist in dense clumps, so perhaps the most important advance needs to be understanding the ejecta structure. This requires a combination of spectral emission calculations and hydrodynamical modeling to fully describe the X-ray emission. Efforts will continue to come up with the theoretical background necessary to understand the growing observations of the extremely energetic explosions known as supernovae.

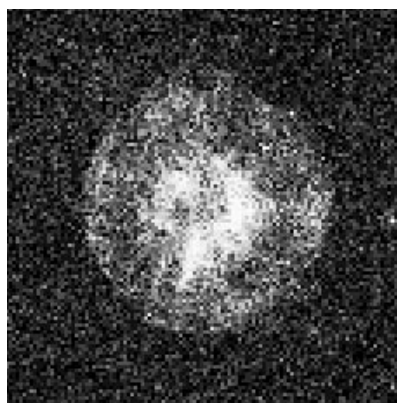


Figure 8.1: Raw image of SMC SNR 0049-73.6, from OBSID 3907.

Appendix A

8.2.1 ASCA Reduction

ASCA Data reduction guidelines

Step 1: Retrieve the data files from archive

Goto the HEASARC web page at <http://heasarc.gsfc.nasa.gov/> and click on archive, then choose an archive browser (my preference is w3browse). You can search by object name or coordinates. Select an observation and choose to list the data sets for that observation. From the retrieved list you only need the ASCA Spectra directory, ASCA Screened Events Files directory, ASCA Make Filter File, and the ASCA Orbit File. Select these four and retrieve data products. Download the tar file. Place the tar file in your data directory and unpack using this command:

```
tar xvf *.tar
```

This will create a subdirectory with the 8 digit sequence number in you data directory and 3 subdirectories which still need to be unzipped. Go into each subdirectory: spectra, screened, and aux, then gunzip each file:

```
gunzip *.gz
```

Now return to your work directory and begin reducing the data.

Step 2: Using Xselect

Type `xselect` to get started, on some machines you may have to type in `xanadu` first. Follow the steps below to apply filters to the screened events files to produce a spectrum of the desired object and create a background spectrum for the region removing any sources.

1- Set data directory with the command, `set datadir` , followed with the path to your screened subdirectory of the observation. This should set the `mkf` directory to the `aux` subdirectory, if not then simply use the command, `set mkfdir` , again followed by the path.

2- Next `list datadir` will display the data files in the screened directory. Data file names all have a similiar format - `ad <seq num><instrument><5 digit code for datamode><bit rate>.evt` - for example:

`ad12345678g200270h.evt` is a GIS2 observation with PH

datamode and high bit rate, or

ad12345678s000201m.evt is a SIS0 observation with Faint datamode and medium bit rate
Read in a data file with the command `read e` followed by the filename, the `e` is necessary to indicate the file is a screened events file (.evt).

3- Next `extract image`, which will start up SAOImage and display the entire field of view for the observation. Click on `Color` in SAOImage and use your mouse to adjust the brightness of the pixels in the image, makes sure the edge is clearly seen.

4- Now click on `Region` and select a shape best suited to include the object you wish to observe. Use your mouse to position the region cursor over the entire object: left mouse button moves the center of the region, the center mouse button changes size, and the right mouse button changes orientation. When all of the object is within the region type `s` in SAOImage to freeze the region then click on the `write` button in SAOImage to save the region file.

5- Remain in SAOImage to create region filters to remove the object, and any other sources in the field of view. Create a region around the object as before, but typing `e` in SAOImage will make the region an exclusive one. Save with the `region write` button. Repeat for any other sources. Finally, create an inclusive region about the entire field of view. For SIS observation use a square, and for GIS use a circle. Type `s` in SAOImage, and the use `region write` button to save the region. Exit SAOImage.

6- Return to `xselect`, apply the filter that included the object to the data set in use with the command `filter region <filename>`. A spectrum can now be extracted from the screened events file that only includes that object by typing `extract spectrum`. Then `save spectrum` will ask to rebin the data before saving, choose yes, and save it to your work directory as a .pha file. To double check that only the object is included you could use the `show status` command that displays `xselect`'s current status including applied filters, or you could `extract image` after the filter is applied to start up SAOImage with what should be an image of only the object.

7- Create background spectrum file by first clearing the

filter you just used with the command `clear region all`. Then filter region `<filename>` with the exclusive region and the region surrounding the entire field of view. `extract image` should display a image in SAOImage that has the object removed. Adjust the color as above to ensure that no photons from the source spill over into the background region, make a new region if necessary. Then `extract spectrum` and `save spectrum` to create a background spectrum for subtraction. Repeat for each instrument. Exit Xselect.

Step 3: Creating Response Files

First you need to create response matrix files (.rmf) for SIS observations, and then the ancillary response files (.arf). Response matrix files for the GIS observations are standard, and they can be found in the spectra subdirectory of you data directory. The GIS rmf files should be moved into your work directory. Now use the FTOOL `sisrmg` on each SIS spectra. This will ask for the pha filename, the arf filename, and the output rmf filename. You don't need an arf file to run `sisrmg`, so you can choose none for the arf filename. After `sisrmg` is complete, you can now create your arf files.

The FTOOL `ascaarf` will ask for the pha filename, the rmf filename, and the output arf filename. You need to create and arf file for each instrument, so run `ascaarf` for SIS0, SIS1, GIS2, and GIS 3 spectra. Now your work directory should include four files for each instrument: pha spectra, pha background, rmf file, and an arf file.

Step 4: FTOOL `grppha`

Before going into XSPEC to fit the spectra, you should use the FTOOL `grppha`. Type `grppha` and then you will be prompted for the input pha filename and the output pha filename. This FTOOL can rebin the data so that each data bin contains a minimum of 25 counts with the command `group min 25`. After the data has been rebinned, you can set the keywords to the arf, rmf, and background files you have also created. This will save time in XSPEC reading in the response files. In `grppha` the command `chkey <keyword> <filename>` will reset the keyword to the file you choose, for example:

```
chkey res snr_s0.rmf      -- this changes the response
keyword to the file snr_s0.rmf
The keywords to change are: res (for rmf files), anc (for arf
```

files), and backfile (for background pha file). When the keywords have been set, use the `write` command to save the rebinned data file with the adjusted keywords to a new file.

Step 5: Using XSPEC

Now you are ready to use XSPEC to fit models to your data. Start by typing `xspec` and then read in your rebinned data file with the command `data <filename>`, for simultaneous fits `data <file1> <file2> ...`. With the correct keyword settings the arf file, rmf file, and background spectrum for each data file will be located by XSPEC. There maybe a problem with the background file for the GIS fields, if the number of pha bins is not the same, background subtraction cannot take place. If this occurs, record the number of bins for the data file and then return to `xselect`. Repeat the procedure for creating a background spectrum. Before saving the spectra, use the command `set binsize`, followed by the number of bins you recorded for the data file. Now save the spectra, but answer no when `xselect` asks to rebin the spectra before saving. You will also need to reset the backfile keyword with `FTOOL grppha`.

8.2.2 *Chandra* Reduction

Data reduction steps for observations in the back-illuminated S3 chip. These notes are to be used in concert with the reduction threads available a the CXC's web page -- <http://cxc.harvard.edu>

1) Download data

the files needed for reduction are:

```
-events file *_evt1.fits
-aspect offset file *_asol1.fits
-bad pixel file *_bpix1.fits
-filter file *_flt1.fits
```

2) Prior to any processing we must update the `geom.par` file, to assure the most recent values are used. This pixel size is slightly changed, so check that parameter. Observation specific bad pixels file must be set in the `arbiblib.par` file rather than using standard CALDB values.

3) Run `acis_process_events`. Follow the Clean Very Faint Mode data thread for the proper parameter values.

4) Filtering (using `dmcopy`)

- Grade values of 0, 2, 3, 4, and 6 should be kept
- Good Time Interval, use the filter file to remove undesirable time frames from the data
- Chip, to only examine S3 (`ccd_id=7`) chip
- Energy, can choose any range, but for spectral extraction choose 0.3 keV for lower limit

5) Use `SAOimage` to create source and background regions

```
>ds9 *_evt2.fits &
```

- Draw region around feature you wish to investigate
- Before saving the region file set `FORMAT` to `CIAO` and `COORDINATE SYSTEM` to `PHYSICAL`
- Save `source.reg`
- Draw source-free region for background extraction

6) Run `psextract` script

- First you must set the parameters of this script, this can be done as input when running the script, or using the `pset` command before running it

- To set the parameters first you need to do the following:

```
>punlearn psextract
>pset psextract events "*.evt2.fits[sky=region(source.reg)]"
>pset psextract bgevents "*.evt2.fits[sky=region(bkgd.reg)]"
>pset root = spec
>pset aoff = "*.asol1.fits"
>pset bgoff = ""
```

- Now you can run `psextract` (you can set `verbosity = 2` to follow the scripts' progress)

```
>psextract verbose=2
```

- Output of `psextract`
 - `spec.pi` The source spectrum
 - `spec_bg.pi` Background spectrum
 - `spec_bg_grp.pi` Grouped background spectra

spec.rmf Response matrix file
 spec.arf Auxiliary response matrix
 spec_bg.arf Background auxiliary response

7) ACISABS

-A correction must be applied to the *.arf file. The ACISABS model accounts for contamination since launch.
 -grppha can then be used to reset the keywords

8) Analyze the data

-Processing is complete, and the data is extracted, turn to XSPEC.

8.2.3 *Chandra* CTI Reduction

Charge transfer inefficiency, CTI, is a serious problem for Chandra. Although effects are not as large in the S3 chip, it is still important to consider. Below are the steps to extract corrected spectra using the CTI corrector created by Townsley et al.(2000). Script and supporting files available at <http://www.astro.psu.edu/users/townsley/cti/install.html>

1) Run the correctit script

-This uses IDL to process the level 1 events file for CTI. If the script fails, try entering each command by hand replacing "#1" in the script with your events filename. This takes some time, so filter the events file so that correctit only runs on the chip you are interested in.

2) Get PSU *.rmf and *.qeu files

-Depending on focal plane temperature, -110 C or -120 C, the PSU team has generated files appropriate for each chip. At this step you must also update the ardlb.par file, for both bad pixels and the *.qeu file.

3) Run acis_process_events

-The important parameters are:
 doevtgrade = 0
 calculate_pi=no

```
rand_pix_size=0
stop=sky
```

4) Filter events file

```
-Filter GTI with the *.flt1 file with dmcoppy
> dmcoppy "*evt1.fits[@*flt1.fits]" new-evt1.fits
```

```
-Filter status and grades to get level 2 events file
```

5) Extract spectrum

-psextract generates the rmf files automatically, so it cannot be used for CTI. The steps of psextract must be done by hand to pick up the new PSU rmfs.

- 5.1) dmextract, need to match binning with the rmfs so add [bin pi=1:685:1] after [sky=region(source.reg)]
- 5.2) dmstat, for each region record mean x and y values
- 5.3) asp_apply_sim, generated *soff file from *aoff
- 5.4) asphist, creates aspect histogram
- 5.5) mkarf, generates arf file - then run ACISABS
- 5.6) grppha, bin counts and set keywords to background file, the newly created arf, and the PSU rmf.

Appendix B

8.2.4 Heavy Element Code

```

      subroutine heavyel(ear,ne,param,ifl,photar)

      integer ne, ifl
      real ear(0:ne),param(*),photar(ne)

c
c
c Aug 23,01 -- Calculates electron temperature as a function
c of ionization timescale, for a parallel plane shock
c where the contribution of the heavy elements to the
c electron density is included
c
c
c
c
c      Aug 99      Plane-parallel shock models with unequal ion and electron
c                  temperatures. May be used to model a shock section.
c additive model for XSPEC to use with general models.
c parameters (param):
c          (1):      postshock temperature t(keV)
c          (2):      postshock electron temperature t_e(keV)
c          (3):      hydrogen density in cm^-3
c (4..14): abundances of He,C,N,O,Ne,Mg,Si,S,Ca,Fe,Ni
c                  with respect to solar values
c          (15):      ionization parameter taul(cm^-3 s)
c          (16):      ionization parameter tauu(cm^-3 s)
c (17): redshift z
c
c          Parameter (3) should be usually set to one. Its only use
c          in the code is to set hydrogen abundance to zero; in this
c          case this parameter should be set to zero. It is not the
c          density parameter.
c
      real ts, tprevs
      real tau0, tau0p
      real taulow, taulp
      real tse, tsep
      real z, zf, zprev
      real pj
      integer nem, iem, ntt
      parameter ( nem = 400 )

```

```

parameter ( ntt = 5 )
integer ntempr, npt
parameter ( ntempr = 200 )
real tempr(ntempr), tau(ntempr)
*      Conversion factors.
real keV, boltz, tconv
parameter ( keV = 1.60219e-9, boltz = 1.38062e-16 )
parameter ( tconv = keV/boltz )

c
c      the lower and upper limiting values of log temperature,
c      the increment in log temp., and the number of
c      temperature values in the ionization tables.
c

real tll,tlu,dtl
integer nt, lt, it
parameter (tll=4.0e0,tlu=9.0e0,dtl=4.0e-3,nt=1251)
integer ion, iion
integer nion,nionp
parameter (nion=157,nionp=nion+2)
real ei(nionp,nt)
real delarr(nt), fout(nionp), fin(nionp)
real eitot, sumem, sumt, dela
integer ntemp, ntmp
parameter ( ntemp = 300 )
real t(ntemp)
real emb, eme, emz, delem, deleml
real eisum(nionp,ntemp)
real tt

*
*      nelem          - number of heavy elements
*      ielem          - curent index of elements (2 for He^++,
*                      3-12 for heavy elements
*      istart, iend   - dummy integer indexes
integer nelem, nelemp
parameter (nelem=10,nelemp=nelem+2)
integer nlines(3:nelemp), ielem, istart, iend
c      nmax = maximum allowed number of bins
c      workphot(nmax,ielem) = bin fluxes for each element ielem
c      (lines + continuum, in photons/sec) for solar abundances.
integer nmax
parameter ( nmax = 10000 )
real earold(0:nmax)
real workphot(nmax,nelemp)

```

```

*      nline - number of lines, el(iline) - energy of line iline in keV
      integer nline, iline
      parameter (nline=566)
      real el(nline)

c
c xline(iline) = line luminosity of line iline
c xline(iline) = lines
c
      real xline(nline), xlinetot(nline)
      integer nmaxfe
      parameter ( nmaxfe = 146000 )
      real elfe(nmaxfe), xlinefe(nmaxfe), xlinefetot(nmaxfe)
      real tauim1, tau1, tausiz
      integer nfe, n
      integer ifreq, ibine, l, j, i, k, ione

c
c
integer MAX, step, ionnum
parameter (MAX=100000)
integer zz, zp, zion(nelemp)
data zion/1,2,6,7,8,10,12,14,16,20,26,28/
real temps(MAX), taus(MAX), chie(MAX)
real wtempl, wtemp, wchi, tempavg, chiout
real taul, tauh, taumin, taumax, deltau, dchidt
real chiscal, taustep, delmin, tolerance, taulimit
real abundfr(nelemp)
real vshock, mu, neni, iondens
real gsum, gh, ghe, massum, abusum, mintau
real alog10
real g, c1, c2, denom
real pi, echarge, me, lnlambda
real mp, beta, taustart, tshock
parameter(mp=1.67e-24)
parameter (pi=3.1415926536)
parameter (echarge=4.8e-10)
parameter (me=9.11e-28)
parameter (lnlambda= 30.)
real TINY, delta, range
parameter (TINY=1.e-30)
real mass(nelemp)
logical qtmp
data nlines / 15, 15, 15, 20, 22, 48, 70, 71, 157, 133 /
      data qtmp /.true./

```

```

data mass/1.0079, 4.0026, 12.011, 14.00674, 15.9994, 20.179,
. 24.305, 28.086, 32.06, 40.08, 55.847, 58.70/

c
c elin and el values
c
cccccccccccccccccccccccccccccccccccccccccccccccccccccccccccc
c
c line energies in /elin/ (9/81)
c
cccccccccccccccccccccccccccccccccccccccccccccccccccccccccccc
c photon energies for lines in kev
data (el(i), i=1,15)/.364,.436,.463,.470,.477,.069,.091,.024,.299,
1 .304,.308,.355,.335,.371,.297/
data (el(i), i=16,30)/.500,.593,.626,.639,.649,.094,.126,.033,.420,.426,.431,
1 .498,.475,.522,.418/
data (el(i), i=31,45)/.654,.775,.816,.837,.849,.124,.163,.041,.561,.568,.574,
1 .666,.638,.698,.559/
data (el(i), i=46,65)/1.022,1.216,1.278,1.308,1.325,0.190,.262,.062,.905,.914,
1 .921,1.073,1.038,1.127,.120,.126,.141,.902,.127,.892/
data (el(i), i=66,87)/1.473,1.747,1.851,1.879,1.908,.273,.375,.104,1.331,
1 1.343,1.352,1.580,1.537,1.660,.188,.196,.214,1.328,.198,1.314,
2 .165,1.299/
data (el(i), i=88,135)/2.007,2.376,2.505,2.567,2.600,0.371,.498,.129,1.840,
1 1.856,1.865,2.183,2.132,2.292,2.348,.272,.281,.303,.400,.096,
2 1.837,1.829,1.828,.237,.252,.258,.268,.283,.380,.086,1.820,
3 .226,.227,.217,.245,.249,.218,.261,.318,1.802,.201,.224,.225,
4 1.785,.203,1.770,.180,1.755/
data (el(i), i=136,205)/2.623,3.108,3.280,3.351,3.397,.485,.657,.172,2.430,
1 2.447,2.461,2.884,2.825,.450,.420,.436,3.033,3.101,.370,.381,.408,
a .506,.512,
2 .538,.124,.124,.130,.138,.591,2.428,2.417,2.417,.330,.347,.347,
3 .367,.383,.462,.470,.537,.118,.124,.127,2.408,2.378,2.390,.328,
4 .292,.340,.344,.372,.446,2.387,2.408,2.355,.302,.316,.316,2.349,
5 .251,.291,.348,2.348,2.338,.281,2.331,.238,2.315,.204,2.315/
data (el(i), i=206,276)/4.106,4.882,5.167,5.391,5.536,.795,1.061,.285,3.861,
1 3.883,3.902,4.583,4.518,.716,.675,.699,4.822,4.932,.611,.629,
a .660,
2 .827,.838,.886,.207,.207,.218,.225,.969,3.864,3.861,3.923,.564,
3 .579,.574,.608,.636,.756,.765,.800,.191,.200,.204,3.837,3.788,
4 3.808,.551,.348,.579,.564,.614,.761,3.810,3.807,3.759,.451,.549,
5 .549,.534,.605,3.782,3.818,3.754,.520,3.742,.477,3.733,.431,
6 3.711,.393,3.708/
data (el(i), i=277,433)/6.966,8.251,8.700,8.908,9.021,1.287,1.738,.449,6.637,

```

```

1  6.675,6.700,7.881,7.799,1.235,1.171,1.212,8.295,8.487,1.090,
w  1.107,
2  1.148,1.480,1.488,1.533,.379,.379,.386,.397,1.715,6.662,6.636,
3  6.617,1.000,1.051,1.052,1.082,1.123,1.387,1.399,1.467,.350,.365,
4  .374,.074,6.628,6.544,6.579,1.000,0.892,1.040,1.033,1.078,1.378,
5  .074,.095,.100,6.583,6.597,6.554,.954,1.008,1.008,.992,1.069,
6  1.265,.083,.090,.115,6.542,6.553,6.502,.747,.874,.843,.967,
7  1.008,.958,.904,1.270,.102,6.502,6.541,6.464,.823,.819,.823,
8  .918,.918,.816,.898,1.148,1.148,.115,6.467,6.503,6.430,.785,
9  .857,.925,.805,.765,1.016,.119,6.430,6.467,.727,.739,.802,.812,
a  .826,.897,.113,.129,.133,.140,.976,1.023,6.434,.162,.187,
b  .195,.227,.246,.262,.292,.307,.334,.709,.725,.805,6.431,.177,
c  .175,.189,.209,.223,.235,.827,6.424,.210,.175,.161,.130,6.421,
d  .207,.200,.177,.174,6.414,.174,6.411,.154,6.404,.143,6.401,
e  .138,6.400/
data (el(i), i=434,566)/8.034,9.524,10.06,10.27,10.35,1.494,2.026,.536,7.732,
1  7.766,7.805,9.184,8.946,1.441,1.367,1.417,9.667,9.891,1.278,
w  1.302,
2  1.348,1.550,1.560,1.653,.477,.477,.477,.477,2.033,7.749,7.714,
3  7.712,1.242,1.242,1.278,1.222,1.640,1.653,1.734,0.413,0.431,
4  .442,.087,7.654,7.557,7.596,1.192,1.060,1.240,1.229,1.330,1.610,
5  .080,.102,.108,7.607,7.601,7.506,1.127,1.181,1.181,1.166,1.256,
6  1.459,.091,.098,.124,7.596,7.669,7.541,1.033,1.127,1.159,1.285,
7  1.259,1.512,.105,7.528,7.588,7.497,1.170,1.170,1.170,1.355,
8  1.355,1.393,.129,7.496,7.541,7.461,.984,1.069,1.192,1.279,.131,
9  7.461,7.462,.884,.900,.969,.981,.998,.998,1.148,1.240,7.462,
a  .207,.235,.244,.281,.297,.334,.336,.381,.411,.879,7.458,.221,
b  .221,.208,.272,.272,.294,0.935,7.450,.197,7.447,.225,7.439,
c  .202,7.435,.179,7.510/

c
c
c
*      save zion
* SAVE nlines, qtmp, mass, fin, fout, neni
* SAVE ei, xline, elfe, xlinefe, nfe
      SAVE

c
cccccccccccccccccccccccccccccccccccccccccccccccccccccccccccccc
c Start of program
c
c parameter assignment
c
beta=param(1)

```



```

zp = zz + 1
do 200 k=1,zp
fin(ionnum+k)=0.0
fout(ionnum+k)=0.0
200 continue
IF (abundfr(ielem).eq.0.) THEN
fin(ionnum+2)=0.0
ELSE
fin(ionnum+2)=1.0
ENDIF
ionnum=ionnum+zp
300 continue
call densityratio(abundfr, fin, neni)
abusum=0.0
massum=0.0
do 400 ielem=1,nelemp
      abusum = abusum+abundfr(ielem)
      massum = massum+(mass(ielem)*abundfr(ielem))
400 continue
mu=(massum*mp)/abusum
c
cccccccccccccccccccccccccccccccccccccccccccccccccccccccccccc
c
c First step is from CHie=0.0 to CHie=0.0001. The analytical approximation
c for small taus is used to determine the starting point for the main
c loop.
c This step is only done if beta=0.0
c
c
IF (beta.eq.0.0) THEN
temps(1)=0.0
chie(1)=0.0
chie(2)=0.0001
c2=neni
gh=zion(1)**2*abundfr(1)*fin(1)/(abusum*mass(1)*mp)
ghe=zion(2)**2*abundfr(2)*fin(2)/(abusum*mass(2)*mp)
gsum=gh+ghe
ionnum=2
do 500 ielem=3,nelemp
zz = zion(ielem)
zp = zz + 1
do 600 k=1,zp
600 gsum=gsum+(((k-1)**2.)*abundfr(ielem)*fin(ionnum+k))
/(abusum*mass(ielem)*mp)

```



```

        if( delarr(it) .gt. 0.0 ) eitot = 1.0
116    continue
        if( eitot .gt. 0.0 )then
            ntmp = ntmp + 1
            sumem = 0.0
            sumt = 0.0
            do 117 ion=1,nionp
                eisum(ion,ntmp) = 0.0
117    continue
            do 118 it=i,i+ntt-1
                tt=it-1
                tt=tll+dtl*tt
                tt=10.0e0**tt
                dela = delarr(it)
                sumem = dela + sumem
                sumt = dela*tt + sumt
                do 119 ion=1,nionp
                    eisum(ion,ntmp) = ei(ion,it) + eisum(ion,ntmp)
119    continue
118    continue
                t(ntmp) = sumt/sumem
            endif
115    continue
*        Calculate spectra.
*        Continua.
        call contm(t,eisum,ntmp,ear,ne,z,workphot)
*        Zero out lines.
        do 630 iline = 1,nline
            xlinetot(iline) = 0.0
630    continue
        do 640 iline = 1,nmaxfe
            xlinefetot(iline) = 0.0
640    continue
*        Line emission.
        do 110 it = 1, ntmp
            call emis1(t(it),eisum(1,it),xline)
*        Ne- to Li-like Fe ions
            call felines(t(it),eisum(104,it),elfe,xlinefe,nfe)
            do 120 iline=1,nline
                xlinetot(iline) = xlinetot(iline) + xline(iline)
120    continue
            do 130 iline=1,nfe
                xlinefetot(iline) = xlinefetot(iline) + xlinefe(iline)
130    continue

```

```

110      continue
*          Add lines into energy bins.
      irstart = 1
      iend = 0
      do 2000 ielem=3,nelemp
          iend = iend + nlines(ielem)
          do 2200 iline=irstart,iend
c          print *, iline, el(iline)
              call hunt(zf*el(iline),ear,ne,ifreq)
              if( ifreq .ne. 0) workphot(ifreq,ielem) =
.              workphot(ifreq,ielem) + xlinetot(iline)/(keV*el(iline))
2200          continue
              irstart = iend + 1
2000      continue

*          Add Fe L-shell lines.
      do 3200 iline=1,nfe
          call hunt(zf*elfe(iline),ear,ne,ifreq)
          if( ifreq .ne. 0) workphot(ifreq,11) =
.          workphot(ifreq,11) + xlinefetot(iline)/(keV*elfe(iline))
3200      continue
          tprevs = ts
          taulp = tau_low
          tau0p = tau0
          tsep = tse
          zprev = z
          do 20 l = 0, ne
              earold(l) = ear(l)
20      continue
          qtmp = .false.
c          Zero out the spectrum
          DO l = 1, ne
              photar(l) = 0.0
          ENDDO
          DO 3300 j = 1, nelemp
              DO 3400 l = 1,ne
                  photar(l)=photar(l)+workphot(l,j)
3400          continue
3300      continue
*      Normalize ion density to 1 and
*      Correct for redshift.
iondens=0.0
      DO 3500 ielem=1,nelemp

```

```

3500      iondens = iondens+abundfr(ielem)
      DO 1 = 1, ne
        photar(1) = 1e14*photar(1)*zf/(iondens*(taumax-mintau))
c   print *, 1,photar(1)
      ENDDO
      RETURN
      END

*
*
*
c
cccccccccccccccccccccccccccccccccccccccccccccccccccccccccccc
cccccccccccccccccccccccccccccccccccccccccccccccccccccccccccc
c
      subroutine ionfract (eltemp, taul, tauh, abunds, frin, frou)
c
c Calculates ionization fractions fout at electron temperature, temp, for the
c range of ionization timescales taul to tauh.
c
c
integer n, i, j
integer max, min
integer ion, ionout
integer nion, nionp
parameter (nion=157, nionp=nion+2)
real alog10
integer nz, nzp, nzm, nelm, ielm
real eltemp, taul, tauh, frin(nionp)
real frou(nionp)
c
c nz is Zof ion, nzp = z + 1, nzm = z - 1, nelm is the number of elements,
c ielem is the current element
c
parameter (nelm=10)
real abunds(12)
integer nzz(nelm), iun(nelm), iunit
data nzz/6,7,8,10,12,14,16,20,26,28/
data iun/50,51,52,53,54,55,56,57,58,59/
c
c
integer nzmax, nzpmax
parameter (nzmax=28, nzpmax=nzmax+1)
c
c

```



```
do 100 i=2,nzp
100   f(i)=frin(ionout+i)
c
ccccccccccccccccccccccccccccccccccccccccccccccccccccccccccccccccccccc
c
c read in data for correct temperature
c
ccccccccccccccccccccccccccccccccccccccccccccccccccccccccccccccccccccc
                                CALL ftgcvd(iunit, 1, it, 1, nzp, 0.d0, feq,
&                                     qanyf, status)
                                CALL ftgcvd(iunit, 2, it, 1, nz, 0.d0, eig,
&                                     qanyf, status)
                                ioff = 1
                                DO j = 1, nz
                                    CALL ftgcvd(iunit, 3, it, ioff, nz, 0.d0, vr(1,j),
&                                             qanyf, status)
                                    CALL ftgcvd(iunit, 4, it, ioff, nz, 0.d0, vl(1,j),
&                                             qanyf, status)
                                ioff = ioff + nz
                                ENDDO
c
c decompose into eigenvectors
c
do 200 i=1,nz
200 work(i)=fs(i)-feqs(i)
do 400 i=1,nz
fspec(i)=0.0d0
do 300 j=1,nz
300 fspec(i)=fspec(i)+vl(j,i)*work(j)
400 continue
c
c propagate forward
c
do 500 i=1,nz
500 work(i)=fspec(i)*dexp(eig(i)*dbple((tauh-taul)))
sum=0.0d0
do 700 i=1,nz
fs(i)=0.0d0
do 600 j=1,nz
fs(i)=fs(i)+work(j)*vr(i,j)
600 continue
fs(i)=fs(i)+feqs(i)
if (fs(i).lt.0.0d0) fs(i)=0.0d0
700 sum=sum+fs(i)
```

[illegible]


```

c
c Calculates the values of the derivative dchi/dtau
c
c
real tau, chi, dchidtau, difftemp, chitemp
integer ielem, nelem, k
parameter (nelem=12)
real abnds(nelem)
integer ions, ionnum
parameter (ions=159)
real fions(ions)
integer z, zp, zion(nelem)
data zion/1,2,6,7,8,10,12,14,16,20,26,28/
real vshock, mu, neni
real pi, echarge, me, ksb, lnlambda
parameter (pi=3.1415926536)
parameter (echarge=4.8e-10)
parameter (me=9.11e-28)
parameter (ksb=1.38e-16)
parameter (lnlambda= 30.)
real mass(nelem)
data mass/1.0079, 4.0026, 12.011, 14.00674, 15.9994, 20.179,
. 24.305, 28.086, 32.06, 40.08, 55.847, 58.70/
real mp, gh, ghe
parameter(mp=1.67e-24)
real c1, c2, g, gsum, musum, abusum
save
c
c
c
neni=0.0
chitemp=chi
call densityratio(abnds, fions, neni)
abusum=0.0
musum=0.0
do 100 ielem=1,nelem
    abusum = abusum+abnds(ielem)
    musum = musum+(mass(ielem)*abnds(ielem))
100 continue
gh=zion(1)**2*abnds(1)*fions(1)/(abusum*mass(1)*mp)
ghe=zion(2)**2*abnds(2)*fions(2)/(abusum*mass(2)*mp)
gsum=gh+ghe
ionnum=2
do 200 ielem=3,nelem

```

```

z = zion(ielem)
zp = z + 1
do 300 k=1,zp
300 gsum=gsum+(((k-1)**2.)*abnds(ielem)*fions(ionnum+k))/(abusum*mass(ielem)*mp)
ionnum=ionnum+zp
200 continue
mu=(musum*mp)/abusum
  g = 8.*((2.*pi)**0.5)*(echarge**4.)*lnlambda*(me**0.5)*gsum/(3.*(ksb**1.5))
denom=(mu**1.5)*(vshock**3.)
c1= (16.*ksb/3.)*1.5*g*(neni**0.5)/denom
c2=neni
difftemp=c1*(c2*(1-chitemp)-chitemp)/(chitemp**1.5)
dchidtau=difftemp
return
end

```

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