MODELS FOR THE EARLY AND LATE SPECTRA OF SUPERNOVAE

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ABSTRACT: The ionizing and heating effects of the radiation from the supernova on the circumstellar matter are discussed. In particular, it is shown how observations of UV absorption lines can give information about the ionization structure and radiative acceleration of the gas. In the second part, the physics of the late phases of supernovae is studied with special emphasis on the metal rich interiors of massive core-collapse supernovae. These results are then applied to SN 1985f.

1. INTRODUCTION

Mass loss, as inferred from radio observations of recent Type II and Type I peculiars, is in many ways cruical for both the structure of the progenitor and the environment into which the supernova expands. For the various aspects of this subject I refer to reviews by Chevalier (1984) and Fransson (1986), as well as the contribution by Chevalier in this volume. In this contribution I will discuss two recently studied consequences of mass loss, first the observability of the circumstellar gas around the supernova and in the second part the late spectrum of a supernova, which has lost its hydrogen envelope.

2. EARLY UV SPECTRA AND THE SUPERNOVA ENVIRONMENT

2.1 Heating and ionization by the supernova shock wave

The interaction between the expanding supernova ejecta and the circumstellar matter, ejected during previous mass loss, creates a region of shocked gas, with a blast wave propagating into the circumstellar gas with a velocity of ~ 10^4 km/s and a reverse shock heating the ejecta (Chevalier 1982). Hydrodynamic models indicate that for Type II supernovae the velocity of the reverse shock with respect to the ejecta is $(1-2)\times10^3$ km/s, depending on the density gradient of the ejecta. Consequently, the temperature behind the blast wave and the reverse shock are $(1-2)\times10^9$ K and $(1-2)\times10^7$ K, respectively. The blast

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wave is cooled mainly by Compton scattering of soft photons from the supernova photosphere by the nearly relativistic electrons in the shocked gas. The result is an up-scattering of the photons, known as Comptonization, creating a power-law tail of the flux into the UV and EUV (Fransson, 1982). The reverse shock, on the other hand, radiates most of its energy as soft X-rays. Depending on the density of the circumstellar gas, and thus the mass loss rate of the progenitor, the Comptonized UV flux dominates the ionizing radiation from the supernova for at least the first months, and probably during the first year, after the explosion.

Half of this flux will be emitted inward from the shock and will there be absorbed by the supernova ejecta. This energy will then be reemitted as UV emission lines in the same way as in QSO emission line regions (Fransson, 1984). The outgoing flux will heat and ionize the circumstellar gas, which has several interesting consequences. First, the mass loss rate, as inferred from the free-free absorption of the radio emission (e.g. Sramek, this volume), is proportional to $T_e^{3/4}$, where T_e is the gas temperature. In the past a temperature of 10^4 K has been assumed without any strong justification. However, in the early phases the heating by the radiation from the shock may give a considerably higher temperature. In the late phases the gas starts to recombine, and if unity optical depth is reached before ~ $10^3\ \dot{M}_{-4}/u_6$ days, the gas will only be partially ionized. (Here $\dot{M}_{-4}$ is the mass loss rate in $10^{-4}\ M_\odot$/yr and u_6 the wind velocity in 10 km/s.) Both effects tend to decrease the optical depth. Another interesting effect is that the ionizing flux from the shock may give rise to observable absorption lines in the UV from the circumstellar gas, which may be an important diagnostic of this medium. Lundqvist and Fransson (1986) have calculated the flux from the shock wave as realistic as possible and studied the effects of this radiation on the surrounding gas. A complication is here that the recombination time of the gas for the relevant ions is long compared to both the decay time scale of the flux and the expansion time scale of the shock wave, and a time dependent calculation is necessary. Since the column density of the circumstellar gas decreases as r^{-3}, the gas close to the shock is most important for the absorption. For $\dot{M}_{-4}/u_6$ ~ 1 the temperature is ~ 2×10^5 K after 20 days and even after one year ~ 2×10^4 K. When the free-free optical depth is unity, the gas in front of the shock is already partially neutral. Taking these effects into account we derive a mass loss rate of 1.3×10^{-4} $M_\odot$/yr for SN 1979c and for SN 1980k 2×10^{-5} $M_\odot$/yr, both for a wind velocity of 10 km/s.

The strong flux of photospheric and Comptonized EUV photons immediately after the shock has formed essentially fully strips the atoms of electrons, and the abundances of ions like C IV, N V and O VI are very low. As the flux decreases and the ions have time to recombine, the abundances increase and thus also the column densities. For $\dot{M}_{-4}/u_6$ ~ 1, the column densities increase from less than 10^{14} cm^{-2}, 20 days after the explosion, to ~ 3×10^{18} cm^{-2} for C IV and ~ 1×10^{18}cm^{-2} for N V after 50 days. They then decrease only slowly as the shock expands, due to the

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frozen-in ionization structure. The high column densities found means that it is very likely that these ions should be observable as absorption lines in the UV resonance transitions. To estimate the strengths of the lines it is, however, necessary to know the velocity of the absorbing gas, which brings us to the next section.

2.2 Radiative acceleration of the circumstellar gas

The fact that the optical depth of both the continuum and lines may be appreciable, means that a large fraction of the momentum in the radiation from the supernova is deposited in the circumstellar gas. There are several processes, which may be important for this. The simplest to treat is electron scattering, which gives an acceleration

$$\Delta v = \frac{\sigma_e \int L dt}{4\pi\, m_p\, c\, r^2} = 1.1\times10^2 \, \frac{E_{49}}{r_{15}^2} \quad \text{km/s} \qquad (1)$$

where E_{49} is the total energy in 10^{49} ergs emitted in photons and r_{15} the distance in 10^{15} cm. As an example we take the well-observed SN 1979c where the <u>observed</u> emitted energy was ~ 1×10^{49} erg/s. Although at least twice as much could have been missed before the supernova was discovered, it is unlikely that the maximum imparted velocity of the gas was more than ~ 200 km/s, if electron scattering dominated. Chevalier (1981) suggested that the radiative acceleration in the resonance lines increased the acceleration by a large factor, and made a rough estimate of the effect. Using the results of Abbott (1982) for the radiative acceleration of ordinary stellar winds, it is possible to make a fairly accurate estimate of the velocity of the gas. As we have seen, the UV lines have large column densities and are therefore likely to have large high optical depths in their resonance lines. Once a line has become optically thick it can, however, not absorb more momentum from the radiation field, and from the work by Castor et al. (1975) we know that the total acceleration is dominated by a large number of lines with small or moderate optical depths, in a wavelength range where the flux is high. Abbott (1982) has studied the acceleration as a function of the effective temperature of the radiation and the density, and finds that over the range 6000-50000 K the acceleration is surprisingly insensitive to the state of ionization. It is therefore possible to evaluate the 'force multiplier', Q, defined as the ratio of the calculated force to that of pure electron scattering, which is a simple function of the density and instantaneous velocity gradient. For $\dot{M}_{-4}/u_6 \sim 1$ and a velocity of ~ 10^3 km/s at 10^{15} cm, we find Q ~ 4-10 , falling rapidly below T_{eff} = 6000 K. This insensitivity to the effective temperature makes it fairly easy to evaluate the effect of acceleration for supernovae, since the acceleration then only depends on the total radiated luminosity from the outbreak of the shock until the effective temperature has fallen below 6000 K, which occurs after about one month. For SN 1979c, with E_{49} ~ 1, we find $v(r) \sim 1\times10^3/r_{15}^{3.3}$ km/s. As was remarked above, the energy radiated before the discovery could, however, be of the same magnitude, and the velocity a factor of 2 higher.

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Conversely, since the force multiplier is probably known within a factor of two, it may be possible to estimate the total radiated energy from the supernova in this way. An uncertain factor is the effect of the energetic EUV and X-ray flux from the shock. If the gas is pre-accelerated to the ejecta velocity immediately after the explosion, there will be no viscous shock and only the photosperic flux is important (Epstein, 1981). Not until after one expansion time scale will the shock develop (Fransson, 1982). Conversely, an early burst of hard flux (Chevalier and Klein, 1979) will completely strip all atoms of electrons, and only electron scattering with is small crossection will be important, so the velocity will be small.

Assuming that hard flux initially is small, we can now calculate the optical depth of the lines using the Sobolev approximation, and it is found that both the C IV and N V lines will have optical depths $> 10^2$. The maximum velocity of the lines should after ~ 10 days be $\sim 1.2\times10^3$ km/s (for $E_{49} = 2$). Because of the large line width and high optical depths, these lines should be easy to observe with the Space Telescope. In fact, they may already have been seen with IUE (Fransson et al., 1984, Panagia, 1982), since both SN 1979c and SN 1980k had very strong absorption components in both C IV, Si IV and N V (Fransson et al., 1984, Panagia, 1982), consistent with the strengths expected from a circumstellar gas. Unfortunately, the resolution of IUE is only 500-1000 km/s, and it was impossible to separate the lines from the components arising in the galactic haloes of our Galaxy and the parent galaxy. Therefore, this type of observation has to await the Space Telescope, but should then be an important diagnostic of the structure of the circumstellar gas around supernovae.

There is, however, one other source of information, which is not hampered by low resolution, in the optical spectra. The 1984 supernova in NGC 3169 observed by Dopita et al. (1984) displayed an unusual Hα line. On top of a broad ($\sim 1.55\times10^4$ km/s) emission component arising in the supernova envelope, was a narrow ($\sim 3\times10^3$ km/s) P-Cygni component. Also SN 1979c (Branch et al., 1981) had a similar line profile, although less extreme. Balancing the recombination rate with the escape of Ly α photons it can be shown that the population of the n=2 level is high enough to make Hα optically thick even in the low density stellar wind, and it is therefore likely that these P-Cygni components arise as a result of scattering by the accelerated circumstellar gas. This will also bring the mass loss rate of the 'super wind', determined by Dopita et al., down to a reasonable level of $\sim 10^{-4}$ $M_\odot$/yr, since the relevant parameter $\dot{M}/u$ and the outflow velocity of the pre-supernova wind could then be ~ 10 km/s, instead of 3000 km/s.

3. LATE TIME SPECTRA OF CORE-COLLAPSE SUPERNOVAE

3.1. Physical processes

For an understanding of the nucleosynthesis, late time observations of supernovae are probably the best source of information. Late time

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here means from ~ 100 days to several years, the latter limit determined by the increasing difficulty of observing the fading supernova. At this time the supernova envelope, if present at all, has become transparent and it is possible to directly observe the synthesized material. The remarkable observations of the oxygen dominated spectra by Filippenko and Sargent (1985) and Kirshner, discussed in the contributions by Chevalier and Filippenko, are good illustrations, but also late time spectra of ordinary Type II:s should show a similar behaviour. In this section I will therefore discuss the physics of this phase in some detail, mainly with application to massive, core-collapsing supernovae. For similar discussions of low mass Type I:s see Meyerott (1980) and Axelrod (1980).

Stars more massive than ~ 10 $M_\odot$ are in general assumed to give rise to Type II supernovae. This classification is, however, based only on the presence of hydrogen lines in their spectra. Since the dynamics of the collapsing core is independent of the outer envelope, this classification is with regard to the nucleosyntesis rather misleading. It is therefore likely that there is continuity in this respect between ordinary Type II:s and what Roger Chevalier calls the Type III:s, consisting of massive stars which have lost their envelopes. In his contribution Chevalier argues that the statistical properties of the Type III:s are compatible with these originating from massive stars that have lost their hydrogen envelopes, i.e. Wolf-Rayet stars. Also their radio emission is consistent with this scenario.

The core/mantle region of massive stars develops an onion skin structure with increasingly heavy elements towards the center (for a review see Wilson et al., 1986). In the outer parts of the mantle helium is the dominant element with a fairly high fraction also of carbon and neon. Next, helium burning and the subsequent $^{12}C(\alpha,\gamma)^{16}O$ reaction result in a shell dominated by oxygen, but also with a high fraction of carbon and magnesium. Inside this in the core region we find elements synthesized during the collapse and explosion, with silicon, sulphur, argon and calcium as the most abundant. Finally, close to the mass cut there is a small amount of ^{56}Ni, which will be very important for the rest of the discussion, since the decay of this radioactive isotop provides a source of energy even at late times. For a star without any extended envelope this is also at early times the only source of radiation, due to the adiabatic expansion losses. The exact amount of ^{56}Ni produced in the collapse of a massive star is very sensitive to the mass cut, but from calculations Weaver and Woosley (1980) find 0.1-0.4 $M_\odot$. Strong evidence for the presence of ^{56}Ni in Type II supernovae comes from photometric observations of SN 1979c and SN 1980k by Barbon et al. (1982a, b). After ~ 100 days, the decay follows a perfect straight line with a slope of 114 days, within the observational errors, indicating that all γ-rays are trapped and thermalized. For SN 1979c and SN 1980k, the late time flux observed by Barbon et al. requires 0.06 and 0.02 $M_\odot$ ^{56}Ni, respectively, to explain the observed level. For this estimate H_0=50 has been assumed. Also the ^{56}Ni mass of ~ 0.55 $M_\odot$ determined from bolometric observations of the Type I peculiar SN 1983n (Panagia et al.,

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1986), is consistent with this scenario. In addition, the observations by Kirshner and Uomoto (1986) of the decay of the Hα line in Type II:s indicate excitation by radioactivity.

During the first month the supernova envelope is optically thick in the continuum and the observed spectrum is a complicated function of the density structure and the expansion. The radiative transfer is made difficult by NLTE effects and the velocity field, and is at present an unsolved problem. However, after ~ 100 days the density is low enough to make the continuum transparent and most ions are in their ground states, a condition referred to as the nebular stage. The radiative transfer then becomes relatively simple and the atomic physics is better understood, and I think that this is the most promising of all supernova stages for determining the nuclear abundancies.

The main difference between the models for these massive remnants and the late time Type I models by Meyerott (1980) and Axelrod (1980), is the chemical composition, with iron-group elements dominating the physics in the latter case. Also the smaller mass of the Type I:s reduces the γ-ray trapping, and positron heating is more important. Since both the thermalization of the high energy electrons and the cooling of the gas are sensitive to the composition, I will in the rest of the paper only discuss massive remnants, dominated by oxygen and possibly helium.

More than ~ 15 days after the explosion the main ionizing and heating source due to radioactivity is the $^{56}Co \rightarrow {}^{56}Fe$ decay. In this process 96% of the energy is emitted as MeV γ-rays and 4% as positrons. For the simple case of a uniform expanding sphere, the optical depth of the γ-rays to Compton scattering is $\tau_\gamma = 0.84\ M_{10}\ t_{100}^{-2}$, where M_{10} is the mass of the envelope, normalized to $10\ M_\odot$, and t_{100} the time since the explosion in 100 days. A typical expansion velocity of 1.5×10^4 km/s has been used. Thus, after ~ 100 days the envelope will be optically thin to the γ-rays, and a uniform deposition of the energy is a good approximation. The total γ-ray luminosity is given by $L_\gamma = 1.36 \times 10^{42} (M_{Ni}/0.1)\ e^{-t/114}$ erg/s where M_{Ni} is the mass of ^{56}Ni. Of this a fraction τ_γ will be absorbed in the envelope. The scattering of the γ-rays by the bound and free electrons in the envelope is highly inelastic, and most of the energy of the γ-rays is transferred to the electrons . These primary electrons, with energy of 0.1-0.5 MeV, will subsequently loose their energy in ionization, excitation of discrete levels and by Coulomb scattering of thermal (kT < 1 eV) electrons. At epochs later than ~ 450 M_{10} days, heating by the positrons dominate if these are trapped, which is probably the case unless the magnetic field is strictly radial. They will then directly act as primary electrons and the rest of the thermalization will be the same.

The ionizations by the primaries produce secondary electrons with energy of 10-100 eV, which in turn loose their energy by the same processes as the primaries. Therefore, the primary electrons, created by the γ-ray scattering, give rise to a cascade of non-thermal secondary electrons in a way very similar to that responsible for the excitation of aurorae. Since the ionization and most excitation crossections has an ln(E)/E behaviour at high energy and the Coulomb crossection is proportional to E^{-2}, the former will dominate the losses of the primaries with high energies, whereas heating of the thermal electrons is the most important process for the low energy secondaries. The relative fraction of the energy going into ionizations, excitations and electron heating is rather sensitive to the fraction of free electrons. To determine the fate of the primary and secondary electrons in a gas dominated by oxygen, a Monte-Carlo program has been used, and it is found that direct excitation of discrete levels of oxygen is small compared to the heat input in the thermal electrons for an electron concentration of more than 10^{-4}. This number is roughly the ratio of the excitation and the electron collision crossections at low energy. The energy, χ, needed for each oxygen ionization also depends on the electron fraction, x_e, but for $x_e < 0.1$ it is close to 30 eV, and the ionization rate is then given by $\kappa_\gamma \rho L_\gamma/(4\pi r^2 \chi)$, where ρ is the total density and κ_γ the energy averaged Klein-Nishina opacity (~ 0.03 cm^2/g).

Recombination takes place by both radiative and dielectric recombination and by charge-exchange with other species. Whereas the former two rates are well known, the charge-exchange rates are largely undetermined for reactions like $O^+ + He^o \rightarrow O^o + He^+$. As an example, the $O^+ + Ca^o \rightarrow O^o + Ca^+$ reaction rate is 3.1×10^{-9} cm^3/s at 5000 K, compared to 5.0×10^{-13} cm^3/s for radiative recombination. Thus, if the density of Ca I is more than ~ 2×10^{-4} that of the electron density, charge exchange dominates. If we nevertheless, for simplicity, neglect this process, it is easy to determine the ionization state of the gas, and in analogy with the case in eg. QSO emission line regions, it is found that this is determined by a parameter $\Gamma_\gamma = n_\gamma/n = L_\gamma/(4\pi r^2 cn)$, where n_γ is the effective density of ionizing particles and n the total number density. The electron fraction is proportional to $\Gamma_\gamma^{1/2}$ and for the simple oxygen homogeneous sphere approximately given by

$$x_e = 0.16\left(\frac{M_{Ni}}{0.1}\,\frac{\langle A\rangle}{16}\,\frac{10M_o}{M}\right)^{1/2} t_{100}^{1/2}\, e^{-t/228}\, \frac{R}{r} \qquad (2)$$

where R is the total radius of the supernova and <A> the mean atomic weight of the gas. From this expression it is apparent that the ionization is likely to vary significantly within the envelope, but rather slowly with time.

The energy going into ionizations is reemitted as recombination emission. The fraction of this emitted as visible line emission varies greatly from element to element. For O I the recombination cascade has been calculated by Julienne et al. (1974), who finds that the lines at 7774 Å and 8446 Å accounts for ~ 10% of the total emission. For He I, on the other hand, the important 5876 Å line is only responsible for 1.6%, while most of the emission is coming out in radiation with $h\nu > 13.6$ eV, which will ionize oxygen. Thus, a comparatively large amount of He can be "hidden".

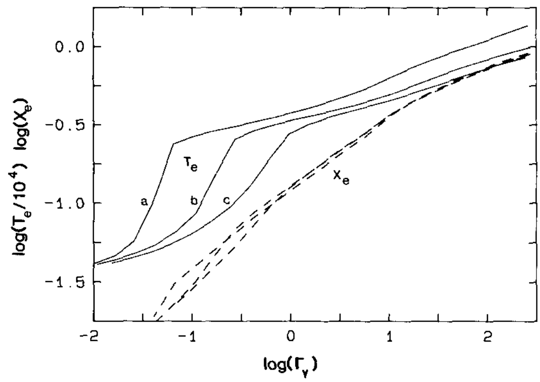
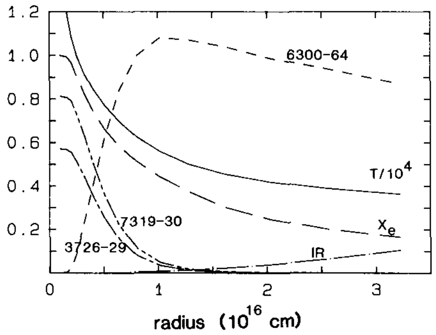


Fig. 1. Gas temperature and electron fraction as a function of the γ-ray ionization parameter, for a pure oxygen gas with a total oxygen density of a 4.2×10^7 cm^{-3}, b 5.2×10^6 cm^{-3} and c 6.5×10^5 cm^{-3}.

The energy absorbed by the thermal electron gas will lead to collisional excitation of levels close to the ground state. Since the typical electron densities are less than ~ 10^7 cm^{-3} at the epochs of interest, the cooling is dominated by forbidden and intercombination lines. In an oxygen rich gas the strongest lines will be the $^3P-^1D$ 6300-64 Å doublet. Because of the high excitation potential of He I, thermal excitation of this ion is unlikely. Whereas the electron concentration is determined by Γ_γ, the temperature is also a function of the density, since collisional deexcitation of the forbidden and metastable lines is important in this density range. In general, when the density increases the temperature has to follow to compensate for the thermalization. Exactly which lines will dominate the cooling is strongly dependent on the relative abundancies. In particular , Mg I, Mg II, Ca II and Na I all have easily excited optical and near-UV levels with large collision crossections. The temperature and spectrum of the remnant will thus be sensitive to the composition.

In Fig. 1 the temperature and electron fraction are shown as a function of Γ_γ for a pure oxygen gas and for three different densities. For high values of Γ_γ the temperature decreases steadily with Γ_γ, and is also sensitive to the density. Above 7000 K ($\Gamma_\gamma > 10^2$) cooling is dominated by forbidden O II lines and below by O I. An interesting feature is the sharp drop in temperature from 3000 K to less than 1000 K at Γ_γ ~ 0.1-0.3. At this temperature cooling by the O I fine structure transitions at 64µ and 140µ become dominant. Since their excitation temperatures are only 200-300 K, the cooling is independent of temperature, which means that a small decrease in the heating, i.e. in Γ_γ, leads to a large drop in temperature. This effect is analogous to that found for iron by Axelrod (1980). Because of the low temperature, it is likely that formation of molecules and dust may take place, unless the γ-rays will dissociate these. This situation could arise in the late phases in the outer part of the supernova, especially if there are density concentrations due to eg. instabilities.



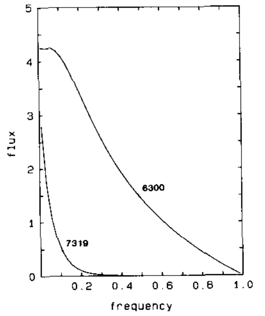


Fig. 2a. Temperature and electron fraction for a 10 M_o remnant at 250 days, when $n = 5.2\times10^6$ cm^{-3}. The contributions to the cooling due to line emission is indicated.

Fig. 2b. Line profiles, for the same model, of the [O I] 6300 Å line and the [O II] 7319 Å line.

Although it is probably premature to do any detailed modelling, some general properties may be seen from simplified models. We have therefore calculated the temperature and ionization structure for a pure oxygen remnant of constant density, heated by the central source of ^{56}Ni. Specifically, we have taken $M = 10$ M_o, $M_{Ni} = 0.1$ M_o and an expansion velocity of 1.5×10^4 km/s. The result of one of these calculations 250 days after the explosion is shown in Fig. 2a. The most interesting result is the highly non-uniform structure in both the temperature and the electron concentration, as was anticipated above. This has the consequence that the emissivity of the different lines vary strongly with radius, so that a one-zone model is a poor approximation. In the same figure, the energy emitted per radial interval, $dL/dr = 4\pi r^2 n^2 j_\nu$, for the [O I] 6300-64 Å and the [O II] 3726-29 Å and 7319-31 Å lines are shown. Close to the center, the O II lines dominate, but further out all the cooling is due to O I. The auroral O I line at 5577 Å is weak because of the low electron density and temperature. A consequence of this non-uniformity is that the line profiles should have quite different widths, depending on the ionization stage and the element. It is also worth noting the increasing importance of the IR lines in the outer parts of the remnant.

The widths and shapes of the line profiles is a rich source of information about the distribution of the individual ions and elements, as well as the total density. For an expanding remnant the velocity will quickly relax to a $v\propto r$ distribution. The flux, F_ν, at frequency ν from a line with rest frequency ν_o is the given by

$$F_\nu = 2\pi \int_{r_{min}}^{R} j_\nu \, n^2 \, r \, dr \qquad (3)$$

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where $r_{min} = ct(\nu-\nu_o)/\nu_o$. In Fig. 2b the profiles of the 6300 Å O I line and the 7319 Å O II line for the model in Fig. 2a are shown, and we note the very different shapes of the O I and O II lines, reflecting the inhomogenous structure. Conversely, from the observations one can, independent of any assumptions about the density distribution, derive the emission per volume in the various lines as a function of the radius, since

$$j_\nu n^2 = \frac{1}{2\pi r}\frac{dF_\nu(\Delta\nu)}{d\nu} \tag{4}$$

where $\Delta\nu = \nu_o r/ct$. If the line, in addition, is dominating the cooling, radiative equilibrium implies that $j_\nu n^2 \propto L_\gamma \rho/r^2$, and the density distribution, $\rho(r)$, can be calculated.

Finally, the time evolution is determined by the decrease in $\Gamma_\gamma \propto t\, e^{-t/114}$, with increasingly low excitation lines of neutral atoms and molecules becoming important. The total flux should, however, basically follow the γ-ray deposition and vary as $\tau_\gamma L_\gamma \propto t^{-2} e^{-t/114}$.

3.2 Application to SN 1985f

To get a feeling for what can be done, we apply these ideas to the observations of SN 1985f (see the contribution by A. Filippenko). We stress, however, that these conclusions are preliminary and mainly illustrative and need more detailed modelling. Unfortunately, the most critical parameter, namely the age of the supernova, is not known. Since this enters most expressions for quantities like the ^{56}Ni mass and the oxygen mass from the line fluxes, little can at present be said about these parameters. There are, however, a number of conclusions, which can be drawn even without this information.

From the calculation of the slowing down of the secondary electrons, we find that the ratio between the energy going into ionization and into heating decreases from 0.37 to 0.10 for $0.1 < x_e < 0.5$. This ratio should be reflected in the relative strengths of the recombination and collisionally excited lines. If a fraction ε of the total recombination emission is coming out in eg. the 7774 Å O I line and a fraction δ of the thermal energy in the [O I] 6300-64 Å lines, we expect that

$$\frac{E_{ion}}{E_{heat}} = \frac{\delta\, F(7774)}{\varepsilon\, F(6300\text{-}64)} \tag{5}$$

Julienne et al. (1974) find that ε ~ 0.06 for the 7774 Å line. From the March/April spectra of SN 1985f, we estimate that F(6300-64)/F(7774) ~ 28 and that 64% of the total flux in the lines is in the [O I] 6300-64 line. Thus δ ~ 0.64 and we obtain E_{ion}/E_{heat} ~ 0.38, which is in good agreement with the theoretical ratio for a gas with electron fraction

~ 0.1, showing that a consistent picture can be obtained.

An important discriminator between different types of progenitors is the helium to oxygen mass. An upper limit for the abundance of helium can be obtained from the O I 7774 and He I 5876 lines, since the fraction of the γ-ray energy going into ionizations is roughly the same for both oxygen and helium. The fraction of He I recombination energy emerging in the 5876 Å line is 1.6%, so we get

$$\frac{F(5876)}{F(7774)} = \frac{\varepsilon(7774)\ X(\mathrm{HeI})}{\varepsilon(5876)\ X(\mathrm{OI})} \sim 4\,\frac{X(\mathrm{He})}{X(\mathrm{O})} \qquad (6)$$

if the medium is mainly neutral, as is indicated by Eq (2). Since the line at 5905 Å is probably due to Na I, we use the strength of this line as an upper limit, and find F(5876)/F(7774) < 1.3 and thus X(He) < 0.3 X(O). This limit should, however, be taken with caution, since the argument rests on the assumption that all the ionization energy emerges as radiative recombination emission. As we have already discussed, charge exchange may be very important, in which case the relative emission in the lines can be very different from the purely radiative value. The relative intensities of the 7774, 8446 and 9264 Å recombination lines are, however, in agreement with those calculated by Julienne et al. for optically thick resonance lines (Case B). With a more detailed knowledge of the charge transfer, similar arguments can, however, be used to obtain abundancies in both this case and for other elements.

Applying Eq. (4) to the blue wing of the [O I] 6300 Å line profile, we find that $j_\nu n^2 r^2$ must be relatively constant inside half the total radius, and then rapidly falling off. If the 6300-64 lines indeed dominate the cooling at all radii, then the density must also decrease, which is somewhat surprising, since most hydrodynamic models evolve towards a shell structure (Chevalier, 1976). That the decrease in emission at large radii is due to a decrease in the γ-ray heating is unlikely, since the remnant should be optically thin to this radiation, unless the remnant is less than 100 days. Possibly, it could be explained by the fact that other lines outside the optical window contribute appreciably to the cooling. If there really was a shell then Γ_γ would be much smaller than in the interior and, as we saw earlier, a thermal instability may take place, with most of the radiation coming out in the IR. A structure different from a shell may also be the result of the Rayleigh-Taylor instability. In particular, the rarefaction wave from the surface after the break-out of the shock, can make the shell unstable (Chevalier, 1976). This would also mix the elements of the different zones. A more detailed study of the line profiles could possibly give some clues to this important problem.

More detailed calculations, incorporating other important elements indicate that especially Ca II and Mg I and Mg II may be very important as coolants. We identify the wide line at ~ 8700 Å with the Ca II IR triplet, rather than N I, which is consistent with the presence of the

7291-7323 Å metastable line and also with the fact that nitrogen is not expected in these zones. The absence of the H and K lines is a result of line trapping, which converts these photons into the other branch. The relative strengths of the Ca II lines indicate that the temperature is more than 4000 K in the line forming region.

These remarks are only meant to illustrate some of the possibilities for obtaining new information about the result of the supernova explosion, and in the future more detailed models are needed, both for the hydrodynamics and for the spectral synthesis. I also think that on the observational side, we have only seen 'the tip of the iceberg'. Late time observations both from the ground and by the Space Telescope open up unique possibilities to directly confront the observations with realistic nucleosynthesis models.

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