Nuclear Astrophysics

II.-V. Core-collapse supernova

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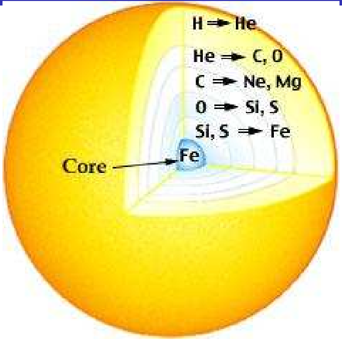
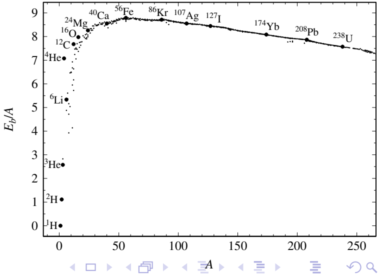
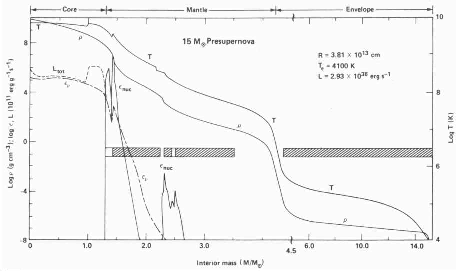
Presupernova star

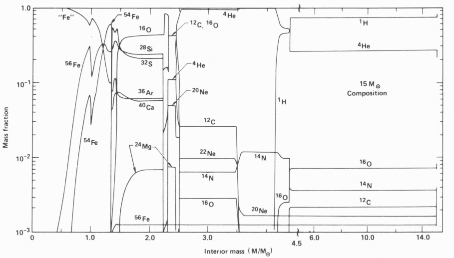
- Star has an onion-like structure.
- Iron is the final product of the different burning processes.
- As the mass of the iron core grows it becomes unstable and collapses once it grows above around 1.4 solar masses.





Presupernova structure





Early iron core

- The core is made of heavy nuclei (iron-mass range $A \sim 45-65$) and electrons. There are Y_e electrons per nucleon.
- The mass of the core M_c is determined by the nucleons.
- There is no nuclear energy source which adds to the pressure. Thus, the pressure is mainly due to the degenerate electrons, with a small correction from the electrostatic interaction between electrons and nuclei.
- As long as $M_c < M_{ch} = 1.44(2Y_e)^2 M_\odot$ (plus slight corrections for finite temperature), the core can be stabilized by the degeneracy pressure of the electrons.

Onset of collapse

However, there are two processes which make the situation unstable.

1. Silicon burning is continuing in a shell around the iron core. This adds mass to the iron core, thus M_c grows.
2. Electrons can be captured by nuclei.

$$e^- + (Z, A) \rightarrow (Z - 1, A) + \nu_e$$

   This reduces the pressure and cools the core, as the neutrinos leave. In other words, Y_e and hence the Chandrasekhar mass M_{ch} is reduced.

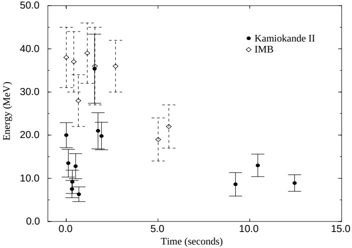
The core finally collapses.

SN1987A

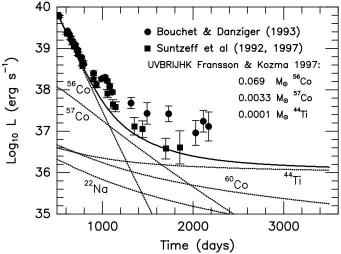
Type II supernova in LMC ($\sim$ 55 kpc)

- $E_{\rm grav} \approx 10^{53}$ erg
- $E_{\rm rad} \approx 8 \times 10^{49}$ erg
- $E_{\rm kin} \approx 10^{51}$ erg $= 1$ foe

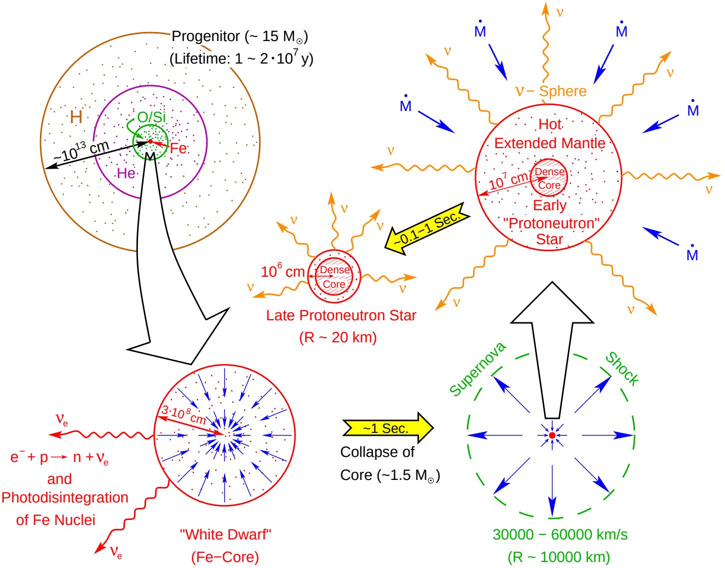
neutrinos $E_\nu \approx 2.7 \times 10^{53}$ erg



light curve



Schematic evolution

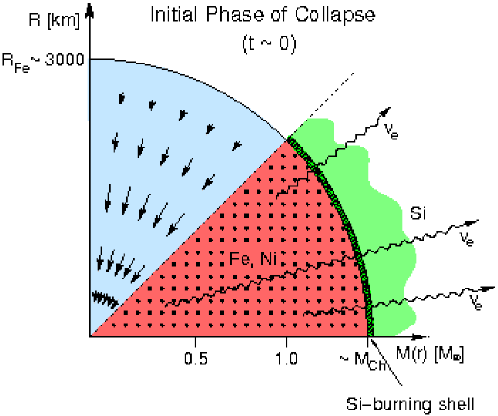


Presupernova and collapse models

Core-collapse supernova simulations are separated into:

1. **presupernova models**:
   - describes the stellar evolution through the various hydrostatic burning stages (H, He,...,Si) and follows the collapse of the central core until densities of order $\rho_9 = 10$ are reached
   - large nuclear networks are used to include the nuclear energy generation and the changes in composition
   - neutrinos, produced in weak-interaction reactions, can leave the star unhindered and are treated as energy loss
2. **collapse models**
   - describes the final collapse and the explosion phase
   - the temperature during these phases is high enough that all reactions mediated by the strong and electromagnetic interaction are in equilibrium; thus the matter composition is given by Nuclear Statistical Equilibrium (NSE)
   - reactions mediated by the weak interaction are not in equilibrium
   - neutrino interactions with matter have to be considered in details (Boltzmann transport)

Central evolution



(from H.-Th. Janka, habilitation thesis)

Initial collapse conditions

If we approximate the pressure by the one of a relativistic degenerate electron gas ($P \sim \rho^{4/3}$) one has

$$P/\rho \approx \frac{1}{4} Y_e \mu_e$$

$$\mu_e \approx 1.1(\rho_7 Y_e)^{1/3}\ \mathrm{MeV}$$

Hence P/ρ is given in MeV, with

$$1\ \text{MeV} = 0.96 \times 10^{18}\ \text{erg/g}$$

Note that the electron chemical potential μ_e is nearly 1 MeV at $\rho_7 = 1$ and hence reaches the nuclear energy scale. Hence it might be energetically favorable to capture high-energy electrons by nuclei.

Electron capture

The cross section for electron capture on *free protons* at rest is

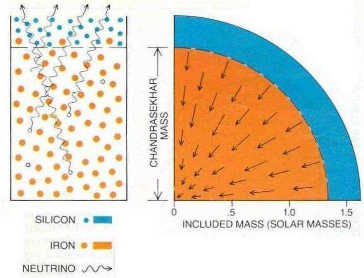
$$\sigma_p = 4.5 \times 10^{-44} E_\nu^2 \ \mathrm{cm}^2$$

where E_ν is the energy of the emitted neutrino in MeV. The rate of electron capture on free protons then is

$$r = \sigma_p N_A Y_p = 0.016 \rho_7 E_\nu^2 Y_p [s^{-1}]$$

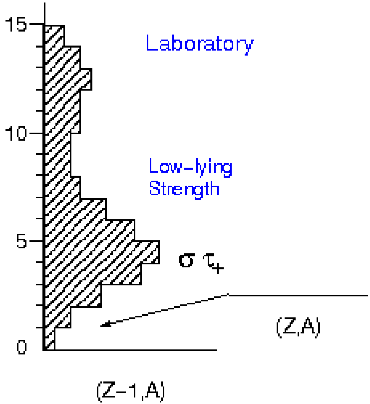
Although the capture cross section for those nuclei present in the core is usually smaller than the one of free protons (due to the larger energy threshold between parent and daughter nucleus), the abundance of free protons is quite low, so that the total electron capture rate is dominated by nuclei. This is an interesting nuclear structure problem which was first tackled within the Independent Particle Model (IPM) and then within the interacting shell model.

Presupernova evolution

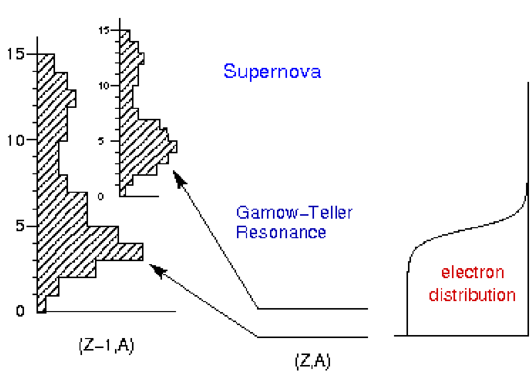


- $T = 0.1\text{–}0.8$ MeV, $\rho = 10^7\text{–}10^{10}$ g cm^{-3}. Composition of iron group nuclei (A = 45–65)
- Important processes:
  - electron capture:
    $e^- + (N, Z) \rightarrow (N+1, Z-1) + \nu_e$
  - β^- decay:
    $(N, Z) \rightarrow (N-1, Z+1) + e^- + \bar{\nu}_e$
- Dominated by allowed transitions (Fermi and Gamow-Teller)

Laboratory vs stellar electron capture



capture of K-shell electrons to tail of GT strength distribution; parent nucleus in ground state

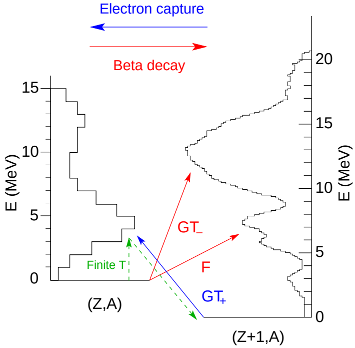


capture of electrons from high-energy tail of FD distribution; capture of strong GT transitions possible; thermal ensemble of initial states

Calculating weak-interaction processes in core-collapse supernova

- Recognition of electron capture during collapse; estimate of rates solely on basis of $f_{7/2} \rightarrow f_{5/2}$ transition
- Derivation of rate formalism; calculation of rates within the Independent Particle Model
  (Fuller, Fowler, Newman, 1982-85)
- Measurement of GT strength distributions by charge exchange reactions
  (n, p) at TRIUMF (1990's), $(d, {}^2He)$ at KVI Groningen (since 2002)
- Calculation of rates using microscopic nuclear models (Shell-Model)
  (Langanke and Martinez-Pinedo, 2001)

Beta-decay, electron capture, GT distributions



At supernova conditions, beta-decays have important contributions from back-rsonances (states with large GT transition in electron capture direction, which are thermally populated).

GT in charge exchange reactions

GT strength can be measured in charge-exchange reactions:

- GT_- measured by (p,n), $(^3\mathrm{He},t)$.
- GT_+ measured by (n,p), $(t,{}^3\mathrm{He})$, $(d,{}^2\mathrm{He})$.

Mathematical relationship ($E_p \geq 100$ MeV/nucleon):

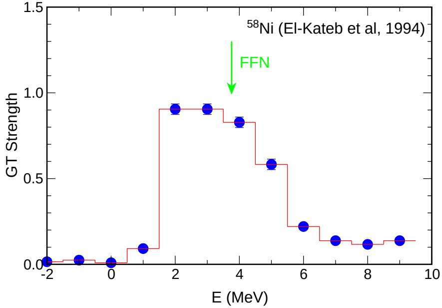
$$\frac{d\sigma}{d\Omega dE}(0^\circ) \approx NB(GT)$$

$$B(GT) = \left(\frac{g_A}{g_V}\right)^2 \frac{\langle f || \sum_k \boldsymbol{\sigma}^k \boldsymbol{t}^k_\pm || i \rangle^2}{2J_i + 1}$$

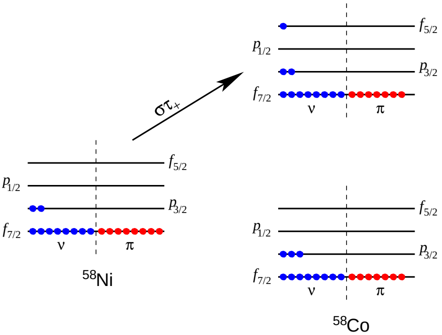
The normalization N can be adjusted to a known transition, e.g. from laboratory electron capture. N is slightly energy-dependent, which, however, can be calculated.

Independent Particle Model

GT$_+$ strength in ^{58}Ni measured by (n, p).



independent particle model (FFN).



The IPM allows for a single transition ($f_{7/2} \rightarrow f_{5/2}$) which is indicated by the green arrow. It does not correctly reproduce the fragmentation of the GT strength.

Shell Model.

INTERACTING **SHELL MODEL**

INDEPENDENT SHELL MODEL
SLATER DETERMINANT ϕ

EXTERNAL SPACE (ALWAYS EMPTY)

RESIDUAL INTERACTION H_{eff}

$O_\alpha = a_i^\dagger a_j$ { MOVES NUCLEON FROM STATE j TO STATE i

$[O_\alpha, O_\beta] \neq 0$

N_s^2 SUCH OPERATORS

HAMILTONIAN

$$\boxed{H = \varepsilon O - \tfrac{1}{2} V O^2}$$

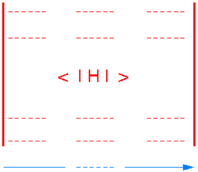
→ IF $V = 0$, H IS PURE I - BODY, SOLVABLE

$N_S \times N_S$ MATRIX ELEMENTS

→ IF $V \pi 0$, $\phi \xrightarrow{H_{eff}}$ {ALL POSSIBLE ϕ's}

FULL COMBINATORIAL DIMENSION

DIAGONALIZATION APPROACH



DIMENSION N

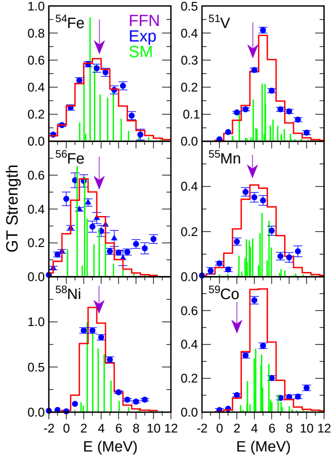
"GIAN

N =

≈ 1

Shell model and (n,p) Gamow-Teller strengths

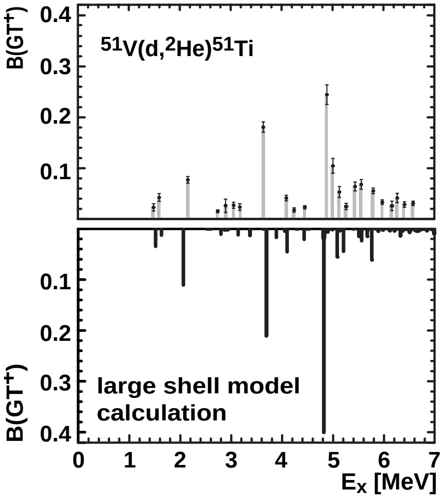
GT_+ strength from (n, p) experiments (TRIUMF).

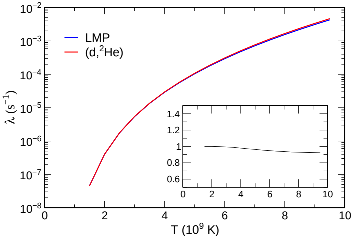


- good agreement with all measured data
- however: (n,p) data have an energy resolution of about 1 MeV

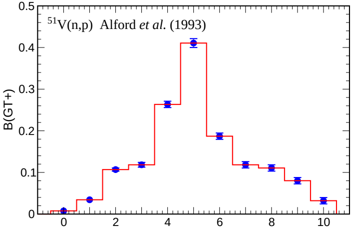
Shell model and $(d,{}^2\text{He})$ GT strengths

C. Bäumer *et al.* PRC **68**, 031303 (2003)

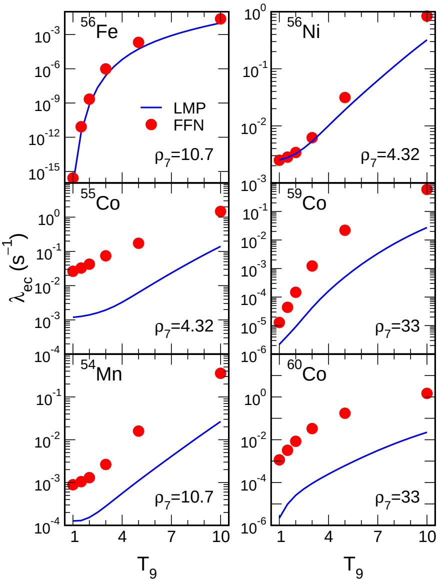




Old (n,p) data

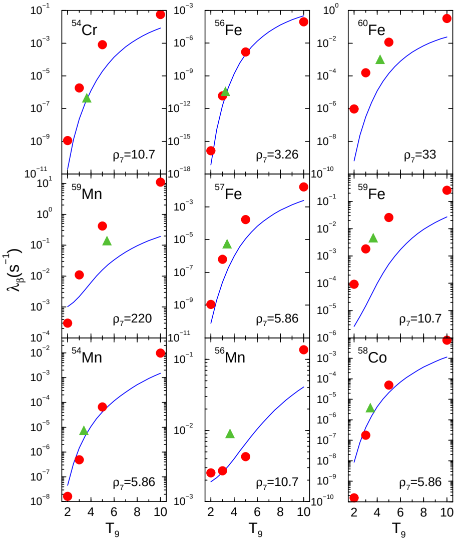


Comparison shell model vs IPM capture rates



- shell model rates on average more than an order of magnitude slower than FFN rates for $A \sim 45 - 65$ nuclei
- shell model β-decay rates similar to FFN rates

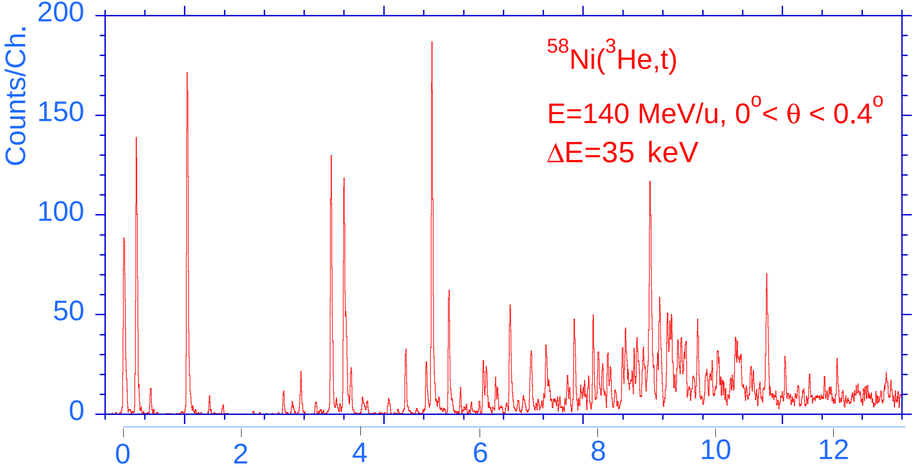
Beta-decay rates during presupernova stage



With increasing core density, β-decays get Pauli blocked.

GT- distribution from (^{3}He,t) data

Equivalent to β^-



Y. Fujita, *et al*

Capture rate during collapse evolution

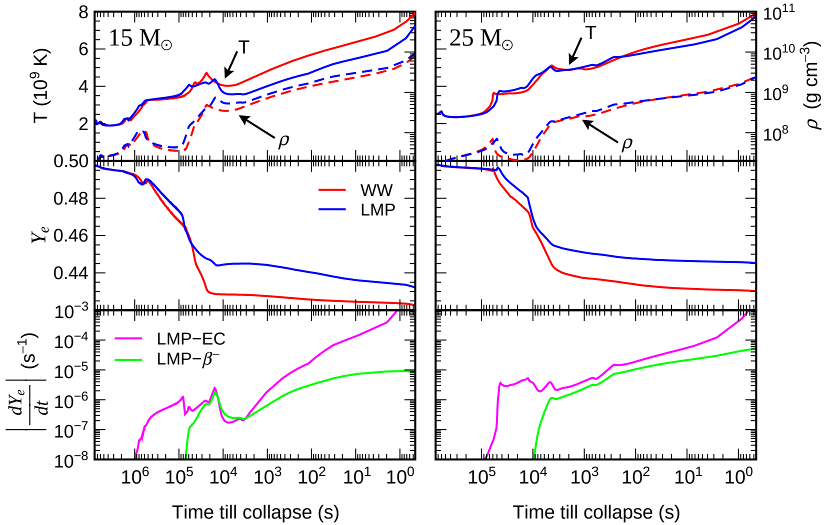
- Composition is determined by NSE. Important parameters (T, ρ, Y_e)
- Weak interactions are not in equilibrium. Y_e evolution has to be computed explicitly.

$$Y_e = \sum_i Y_i Z_i$$

$$\dot{Y}_e = -\sum_i \lambda^i_{ec} Y_i + \sum_i \lambda^i_{\beta^-} Y_i$$

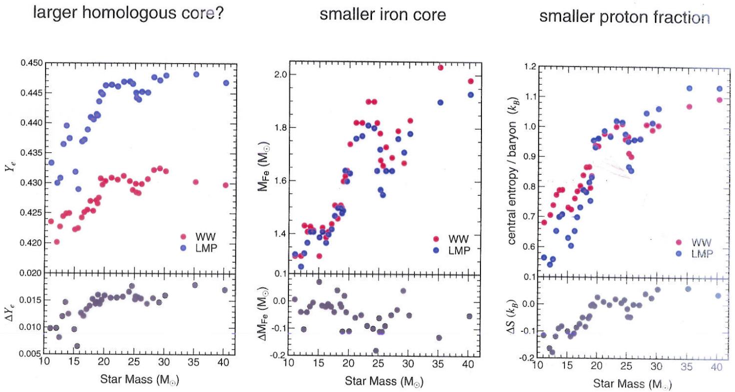
Consequences of weak rates

(A. Heger *et al.*, 2001)

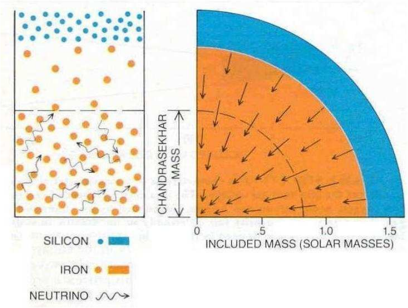


Consequences of weak rates

(A. Heger *et al.*, 2001)



Collapse phase



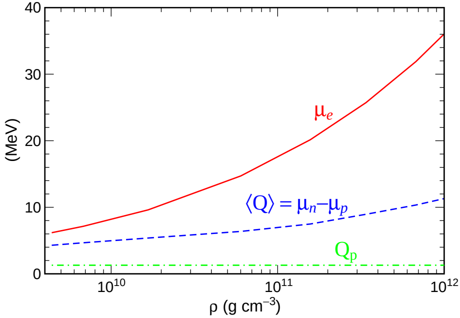
Important processes:

- Neutrino transport (Boltzmann equation):
  $\nu + A \rightleftarrows \nu + A$ (trapping)
  $\nu + e^- \rightleftarrows \nu + e^-$ (thermalization)

  cross sections $\sim E_\nu^2$
- electron capture on protons:
  $e^- + p \rightleftarrows n + \nu_e$

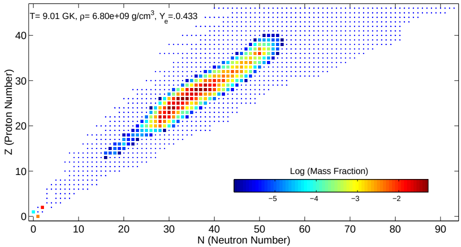
What is the role of electron capture on nuclei?
$e^- + (N, Z) \rightleftarrows (N+1, Z-1) + \nu_e$

What about weak processes on nuclei?

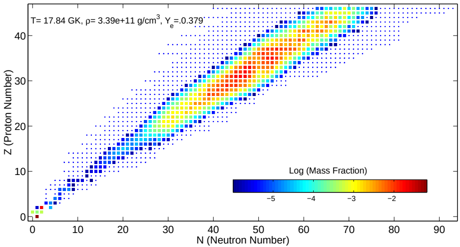


The electron chemical potential grows fast ($\sim \rho^{1/3}$).
This increasingly suppresses beta-decays as the phase space for the electrons gets blocked.
Electron captures increase as electron energies grow. Rates become less sensitive to the detailed reproduction of the GT distributions.
However, an interesting nuclear structure problem arizes which lead to the assumption that capture on heavy nuclei is strongly suppressed.

Collapse abundances



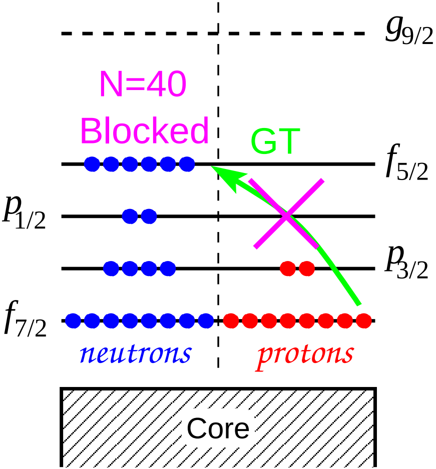
presupernova

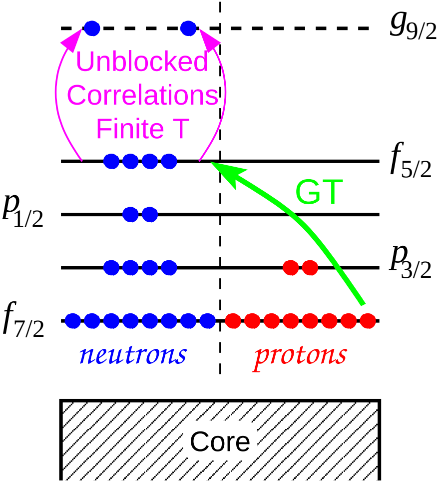


neutrino trapping

(Un)blocking electron capture at N=40

Independent particle model (Bruenn)

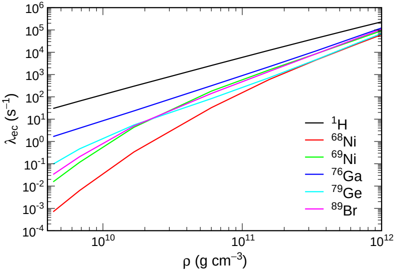




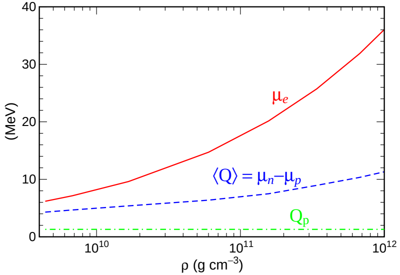
Weak Rates for nuclei with A=65–112 computed using the Shell Model Monte Carlo plus RPA approach

Electron capture: nuclei vs protons

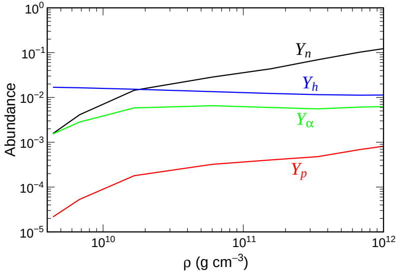
Electron capture rates



Energetics

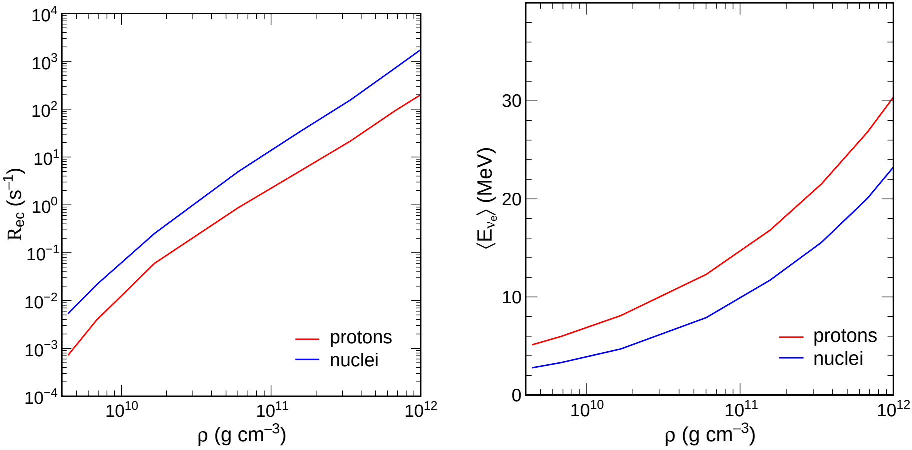


Abundances



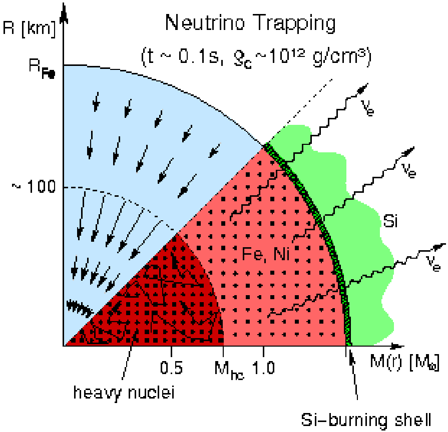
$$R_h = \sum_i Y_i \lambda_i = Y_h \langle \lambda_h \rangle$$
$$R_p = Y_p \lambda_p$$

Reaction rates



Electron capture on nuclei dominates over capture on protons

Neutrino trapping

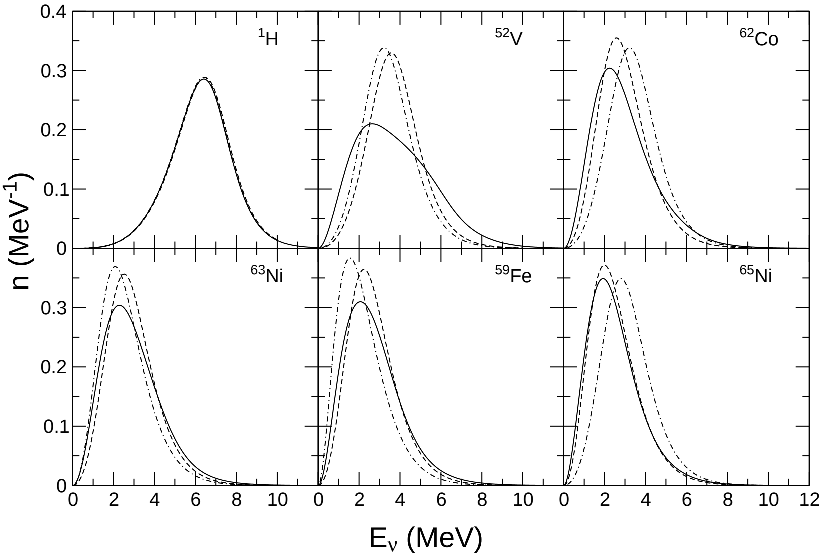


Neutrino-induced reactions

- $\nu + A \rightleftarrows \nu + A$ (trapping)
  elastic process, no energy, but momentum transfer
- $\nu + e^- \rightleftarrows \nu' + e^-$ (thermalization)
  inelastic scattering, energy transfer
- $\nu + (Z, A) \rightarrow \nu' + (Z, A)^*$ (thermalization)
  inelastic scattering, energy transfer
- cross sections $\sim E_\nu^2$

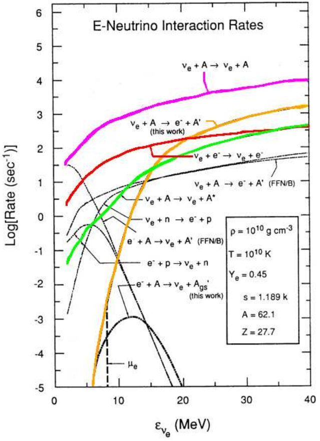
treatment by neutrino transport (Boltzmann equations) which consider all neutrino types and keep track of neutrino fluxes, energies at all space-time points

Typical neutrino spectrum



Neutrino interactions during the collapse

Bruenn and Haxton (1991)
based on results for ^{56}Fe



- Elastic scattering:
  $\nu + A \rightleftarrows \nu + A$ (trapping)
- Absorption:
  $\nu_e + (N, Z) \rightleftarrows e^- + (N-1, Z+1)$
- ν-e scattering:
  $\nu + e^- \rightleftarrows \nu + e^-$
- Inelastic ν-nucleus scattering:
  $\nu + A \rightleftarrows \nu + A^*$

Elastic neutrino-nucleus scattering

Elastic neutrino-nucleus scattering is mediated by the neutral current. The mean-free path of neutrinos in matter, composed of heavy nuclei (mass fraction X_h) and free neutrons (X_n), due to elastic scattering is

$$\lambda_\nu = 1.0 \times 10^8 \rho_{12}^{-1} [\frac{(N - 0.08\, Z)^2}{6A} X_h + X_n]^{-1} E_\nu^{-2} \mathrm{cm}$$

where E_ν is the neutrino energy in MeV.
Taking typical values at $\rho_{12} = 1$ ($N = 50, A = 82$, $E_\nu = 20$ MeV), one has $\lambda_\nu = 0.4$ km. The core radius at that moment is about $R \approx 30$ km. Thus, the neutrinos scatter often in the core and their way out should be treated as a diffusion process. Indeed: the diffusion time scale under these conditions is longer than the collapse time scale (of order 1.5 ms at $\rho_{12} = 1$).
The neutrinos are effectively trapped in the core during the final collapse.

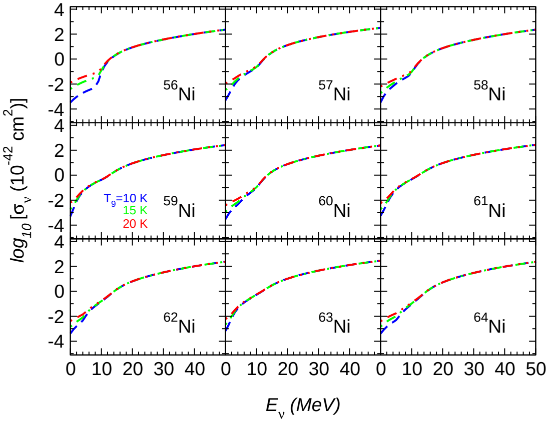
Inelastic neutrino scattering

Current collapse simulations consider inelastic $\nu_e + e^-$ scattering as means to exchange energy between neutrinos and matter. As the electrons are highly degenerate at this stage, they can only gain energy. Hence, neutrinos are down-scattered in energy by this process. This has two consequences:

- Neutrinos of lower energies have larger mean-free paths and may leave the core.
- Neutrinos will be thermalized with the rest of the matter. In simulations this is achieved when the core density is about $\rho_{12} = 1$.

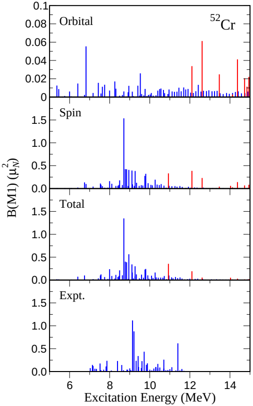
Inelastic neutrino-nucleus scattering is not yet included in simulations as the relevant rates have just been evaluated for the first time.

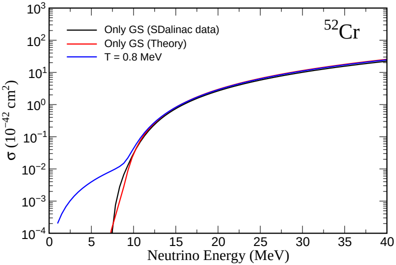
Inelastic neutrino-nucleus cross sections



- large-scale shell model (allowed transitions), finite-T effects
- random phase approximation (forbidden transitions)
- at low neutrino energies enhanced due to finite temperature effects

Inelastic neutrino cross sections from electron scattering



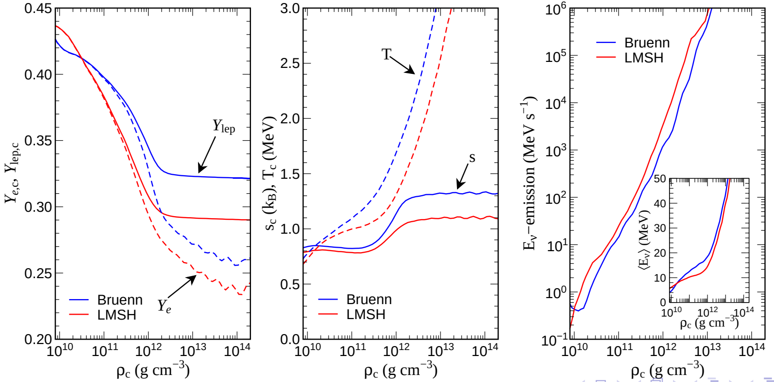


- high-precision SDalinac data
- large-scale shell model

- neutrino cross sections from (e, e') data
- validation of shell model

Collapse simulations

With Rampp & Janka (General Relativistic model)
15 $M_\odot$ presupernova model from A. Heger & S. Woosley



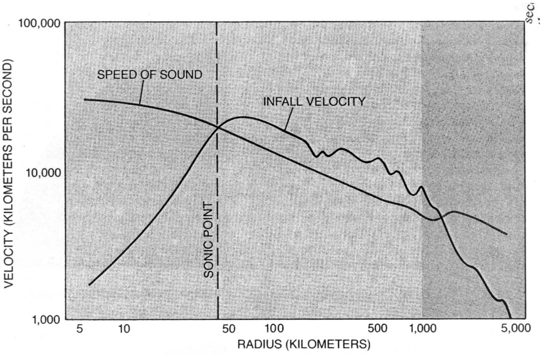
Importance of neutrino trapping

With neutrino trapping and thermalization, also neutrinos become degenerate. They are described by a Fermi-Dirac distribution with a neutrino chemical potential μ_ν which, after establishment of an effective equilibrium fulfills:

$$\mu_\nu = \mu_e - (\mu_n - \mu_p)$$

The presence of degenerate neutrinos also stop the electron capture process and guarantee that a sizable electron fraction (and proton fraction!) survives the collapse.
The inner core, which is effectively in weak equilibrium, collapses as a homologous unit.

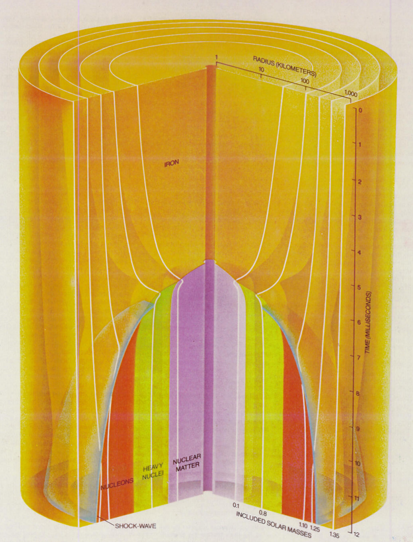
Homologous core

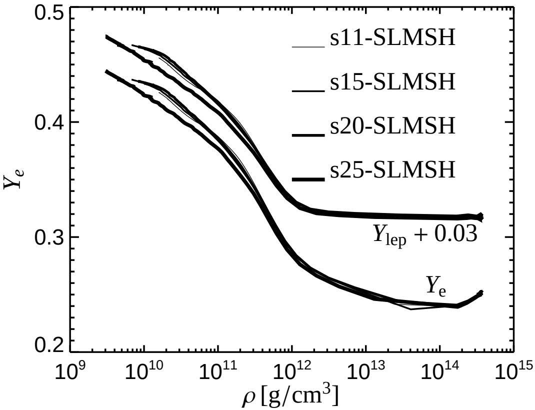


Homologous collapse (velocity proportional to radius)

After thermalization, an inner homologous core forms in which the local sound velocity is larger than the infall velocity. The outer core moves at supersonic velocities. A sound signal from the inside cannot get beyond the sonic point which is slightly further inside than the point where the velocity reaches a maximum.

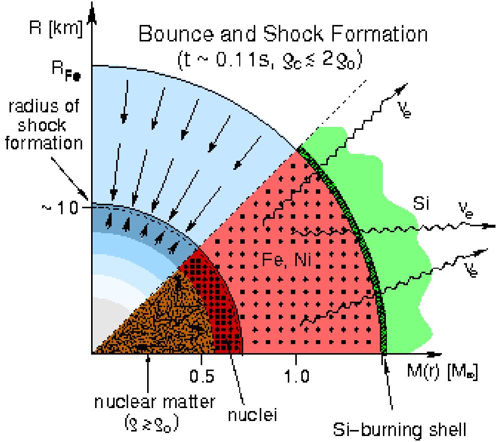
Collapse history





Electron captures on nuclei and protons are self-regulating leading to the same trajectories at bounce for different stellar masses.
(H.Th. Janka, A. Marek, G. Martinez-Pinedo)

Core bounce



R [km]
Bounce and Shock Formation
(t ~ 0.11s, ϱc ≲ 2ϱo)
RFe
radius of shock formation
~ 10
νe
Si
νe
Fe, Ni
νe
0.5
1.0
M(r) [M⊙]
nuclear matter (ϱ ≳ ϱo)
nuclei
Si–burning shell

The collapse is stopped

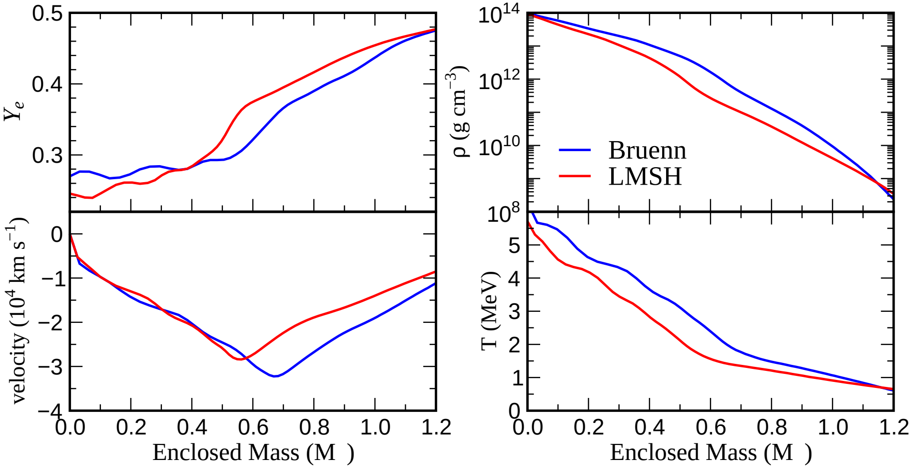
The collapse continues until the central density becomes substantially (by about a factor 2-4) larger than nuclear density ($\rho_{nm} \approx 2 \times 10^{14}$ g/cm^3. Then nuclear pressure slows down the infall and finally stops it. The inner core has reached its maximum density (*maximum scrunch*). After maximum scrunch, the core rebounds and a shock starts.
A decisive quantity for this stage of the collapse is the *Equation of State*. It is assumed that matter consists of nuclear and electron components, while neutrinos have negligible interactions, but are important for the determination of quantities like Y_e or temperature.

The Equation of State

The forces between nucleons and electrons are nuclear and electromagnetic. Most nucleons reside in heavy nuclei until trapping. With increasing density, nuclear matter forms in a two-phase system: nuclei surrounded by a low-density gas of alpha-particles and nucleons. At densities between $\approx 10^{13}$ g/cm^3 and saturation one finds the 'spaghetti', 'lasagna' or 'Swiss cheese' phases, (rods and slabs of nuclear matter, parts of space filled with uniform nuclear matter and holes in-between), and finally nuclear matter filling space uniformly.

Modern EOS are derived from nuclear mean-field models. However, the nuclear composition considers only proton, neutrons, α-particles and a 'representative' heavy nucleus whose (Z, A) values change with temperature and density.

Properties at bounce



(PRL **90**, 241102)

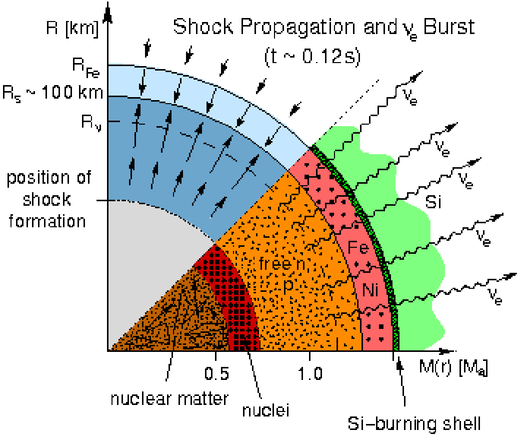
Formation of the shock

When the center of the star reaches and exceeds nuclear density, the material becomes very hard to compress. Pressure builds up and a 'mild' pressure wave propagates outwards. Near the sonic point close to the surface of the homologous core this pressure wave turns into a shock.
In the shock the temperature increases, as does the entropy. So the passage of the shock dissociates the nuclei into free nucleons which costs the shock energy (about 8-9 MeV/nucleon).

Prompt shock scenario

In the **prompt shock scenario** the shock is energetic enough to go through the entire star and expel most of it. However, in most cases (perhaps with exception of the lightest stars around 10 $M_\odot$) this scenario fails. For a success the shock has to travel from the surface of the homologous core (M_{hom}) to the surface of the iron core (M_{Fe}) where it reaches the region of silicon burning and hence a fresh energy source (which would even be more effective if heated). Thus, the shock has to traverse the infalling matter of the core envelope with a size of about $M_{env} = M_{Fe} - M_{hom}$. The shock will dissociate the traversed matter loosing energy. Current simulations estimate $M_{env} \approx 1\ M_\odot$, while the shock's energy (about 10^{51} ergs) is sufficient to pass through about $0.6 M_\odot$.
At a radius of about 200 km, the shock turns into an *accretion shock* in which additional infalling material accretes to the existing core and the outward motion has stopped.

Shock stagnation



Shock stagnation

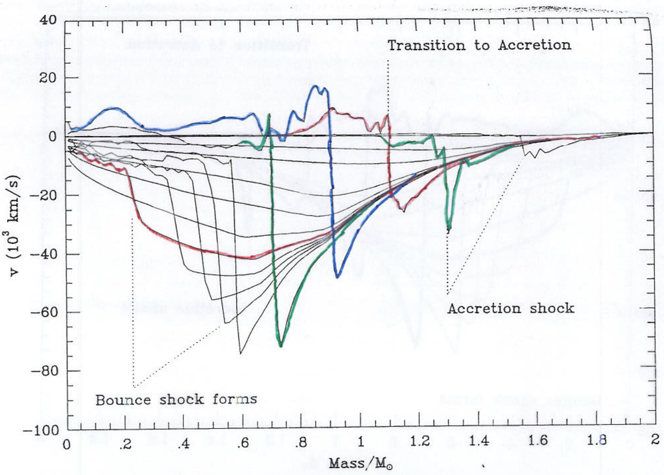
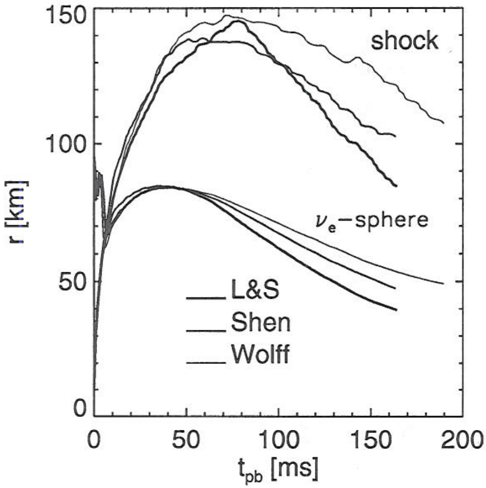


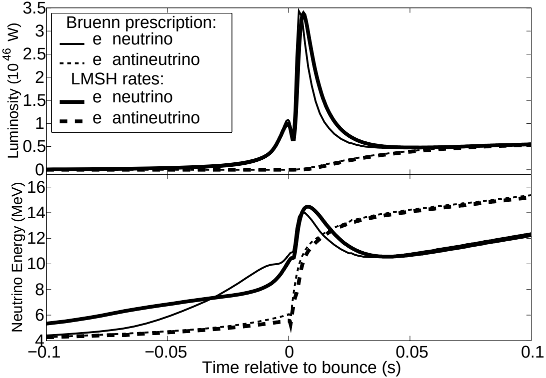
Fig. 12.5. Bounce Shock: Velocity

Dependence of shock on EOS



Neutrinos help the prompt shock to fail

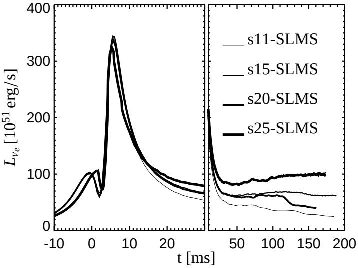
With the slowing down of the shock, (trapped) neutrinos can be emitted further reducing the energy.
Additionally, the shock has changed the matter from heavy nuclei to free nucleons. During collapse protons existed in neutronrich nuclei with large thresholds for electron captures. The later is strongly enhanced once the remaining electrons can be captured on free protons. This leads to a burst in electron neutrinos and costs the shock energy.
At this point, the proton-to-nucleon ratio is reduced to the value of the neutron star which is formed as the remnant in the center.



(PRL **91 201102**)

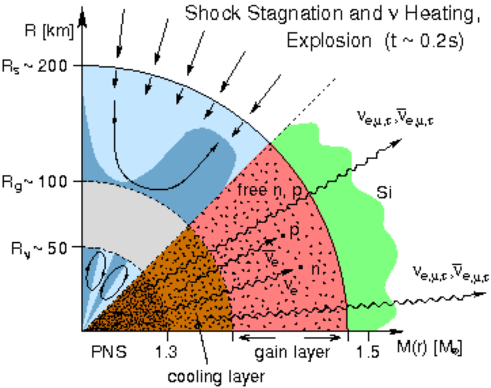
The neutrino burst is very similar for stars of different masses!

'Standard' neutrino burst



- shock dissociates matter into free protons and neutrons
- fast electron captures on free protons create ν_e neutrino burst
- 'standard' ν_e bursts
- future observation by supernova neutrino detectors
- 'standard neutrino candles'?

Shock revival



Supernova neutrinos

Neutrinos play an essential role in a supernova explosion. In fact, most of the energy ($\approx 99\%$) liberated in the explosion is carried away by neutrinos. Besides the ν_e neutrinos produced by electron captures, neutrinos of all types can be produced.
The temperature in the core behind the shock is quite high (10 MeV or more) and many electron-positron pairs exist. These can couple to neutrino-antineutrino pairs where the rate of energy transfer from e^+e^- pairs to $(\nu\bar{\nu})$ pairs is appreciable

$$r \approx 10^{25} T^9 \times (1 + 0.19n)\mathrm{ergs/cm^3/s}$$

where $n = 2$ is the number of non-electron neutrino generations which cannot couple to e^+e^- pairs via charged currents.
Note the strong temperature dependence. Thus neutrino pair production is only important in the inner part of the core.

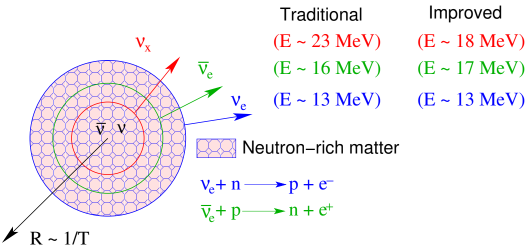
Neutrinospheres

At high material densities, neutrinos initially continue to be trapped. But as the shock proceeds, the density at the shock front decreases and becomes less than the trapping density. Neutrinos can be released, first those directly behind the shock, then those from farther inside. The radius from which they stream out 'freely' is called neutrinosphere. It is defined by

$$\int_{R_\nu} \alpha dr = \frac{2}{3}$$

where α is the number of collisions per cm.
The neutrinosphere for ν_e, $\bar{\nu}_e$ coincides with $\rho_{11}(R_\nu) \approx 1$. At the early stage of the collapse this translates to $R_\nu = 50$ km; later, when the material is less dense, R_ν moves to smaller radii, around 20 km.
However, the different neutrino types have different neutrinospheres!

Hierarchy of neutrino spectra



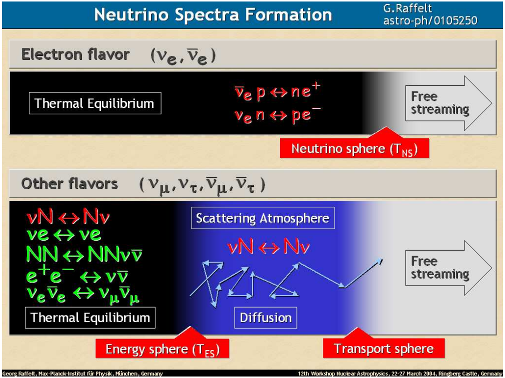
ν_x neutrinos (i.e. ν_μ, ν_τ and their antiparticles) interact only via neutral current; while ν_e and $\bar{\nu}_e$ additionally interact via charged current. The most important reactions are

$$\nu_e + n \rightarrow p + e^-; \quad \bar{\nu}_e + p \rightarrow n + e^+$$

As the matter behind the shock is neutronrich, the opacity for ν_e is larger than for $\bar{\nu}_e$. thus, there is a natural hierarchy for the neutrinospheres:

$$R_{\nu_x} < R_{\bar{\nu}_e} < R_{\nu_e}; \; E_{\nu_x} > E_{\bar{\nu}_e} > E_{\nu_e}$$

Supernova neutrino-spectrum formation



As $\sigma_\nu \sim E_\nu^2$, the neutrinospheres actually depend on the neutrino energies and are diffuse.

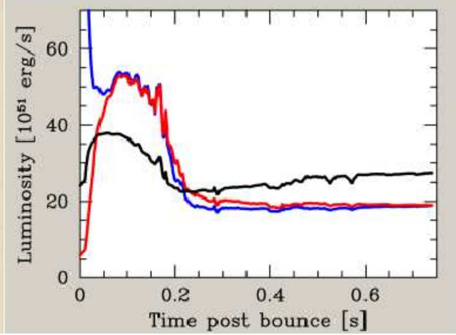
Neutrino spectra

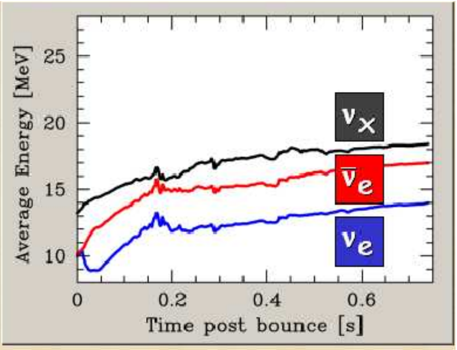
The gravitational energy of the collapsed core (a few 10^{53} ergs) is radiated away in neutrinos of all types. There is a large luminosity in neutrinos ($L_\nu > 10^{52}$ ergs/s) for nearly 10 seconds, before it decreases. The luminosity is nearly the same for all neutrino types and is maintained by mass accretion onto the proto-neutron star where the kinetic energy of infall is converted into thermal energy. The neutrinos have approximately the Fermi-Dirac spectra with zero chemical potential. Then

$$\langle E_\nu \rangle = \pi T_\nu; \; \langle E_\nu^2 \rangle \approx 6 T_\nu^2$$

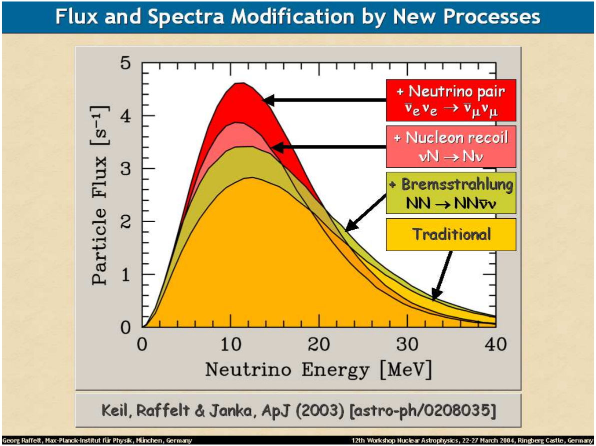
The average energy of the emitted neutrinos ($\sim$ 15 MeV) is much less than the energy of neutrinos produced in the high-density core ($\sim$ 150 MeV). When the neutrinos diffuse out of this core, they are down-scattered in energy. As they carry away the entire energy, there are about 10 neutrinos emitted for every one produced in the center.

Neutrino-spectrum evolution

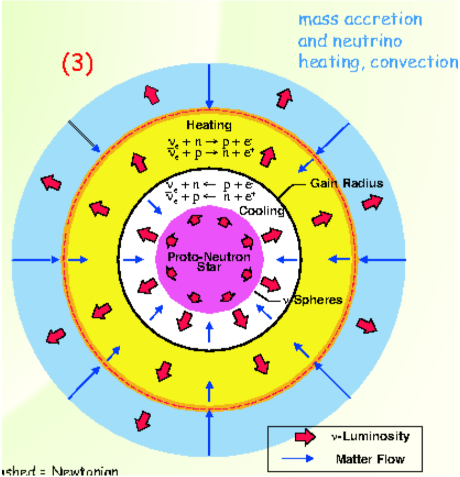




Influence of different reactions on the spectra



Gain radius



- The important reactions directly behind the shock are:

$$\nu_e + n \leftrightarrow p + e^-;$$
$$\bar{\nu}_e + p \leftrightarrow n + e^+$$

- Competition between emission (cooling) and absorption (heating) by neutrinos.
- Due to different (ρ, T) profiles there exists a gain radius R_g behind the shock such that emission dominates for $R < R_g$ and absorption for $R > R_g$.
- Thus the material directly behind the shock gets heated.

Shock revival by neutrino absorption

The cross section for neutrino absorption on a nucleon is

$$\sigma_{abs} = 9 \times 10^{-44} E_\nu^2 \mathrm{cm}^2$$

with E_ν in MeV. If L is the neutrino luminosity, the energy gain per nucleon at distance R_m is

$$[\frac{dE}{dt}] = \frac{0.5 L \sigma_{abs}}{4\pi R_m^2} X_n \approx 25 \mathrm{MeV/s}$$

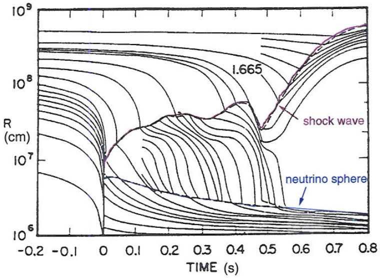
if one uses typical values ($R_m = 200$ km, $L = 5 \times 10^{52}$ ergs/s, $E_\nu = 10$ MeV, $X_n = 1$).
This is to be compared to the energy required to move a nucleon out of the gravitational well of the mass inside the shock $M(R_m)$:

$$E_{grav} = \frac{GM(R_m) m_n}{R_m} \approx 10 \mathrm{MeV}$$

assuming $M(R_m) = 1.5 M_\odot$ and $R_m = 200$ km.
Thus it takes about 0.4 s for the neutrinos to deliver this energy.

The delayed shock model

it is currently believed that neutrino absorption on nucleons indeed revives the shock which then moves outwards and expels the rest of the star. This scenario, in which a prompt shock is initiated, but get stalled and is later revived by neutrinos is called the **delayed shock model**. Simulations show that neutrinos indeed transport energy from hotter zones to the shock and some of this energy is translated into kinetic energy of matter.



Wilson' simulation which led to the discovery of the delayed shock mechanism (1982).

Quest for detailed neutrino transport

The neutrino heating in the gain region

$$[\frac{dE}{dt}] = \frac{X_n}{\lambda_{\nu_e}} \frac{L_{\nu_e}}{4\pi r^2} \langle E^2_{\nu_e} \rangle \frac{1}{F} + \frac{X_p}{\lambda_{\bar{\nu}_e}} \frac{L_{\bar{\nu}_e}}{4\pi r^2} \langle E^2_{\bar{\nu}_e} \rangle \frac{1}{\bar{F}}$$

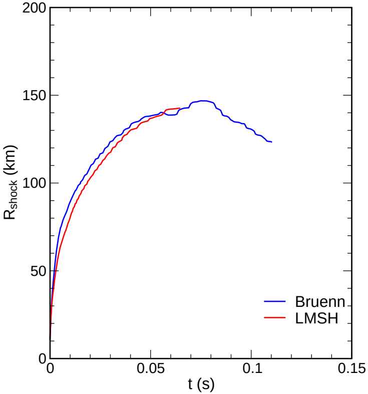
which depends inversely on the flux factor F (which is a measure of the isotropy of the neutrino distribution), linearly on the luminosity and quadratically on the neutrino spectrum (energy). These are the key neutrino ingredients.

The shock revival is a complex interplay of heating and mass accretion through the shock and through the gain radius where the difference determines the mass in the gain region which can be heated. Mass accretion through the gain radius helps to sustain the neutrino luminosity, but it reduces the pressure in the gain region - thus it is good and bad.

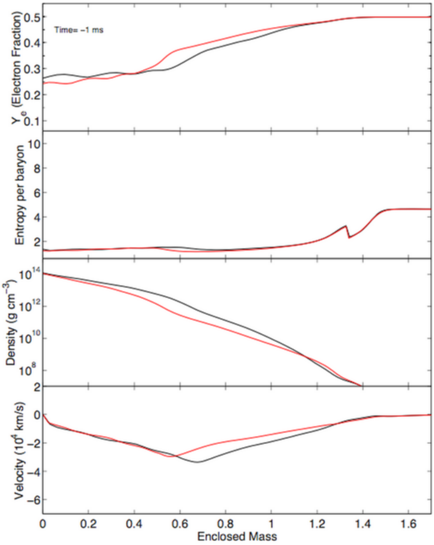
In modern supernova simulations this complex interplay is described by neutrino Boltzmann transport equations which consider also other neutrino-induced reactions.

Shock propagation in one-dimensional models

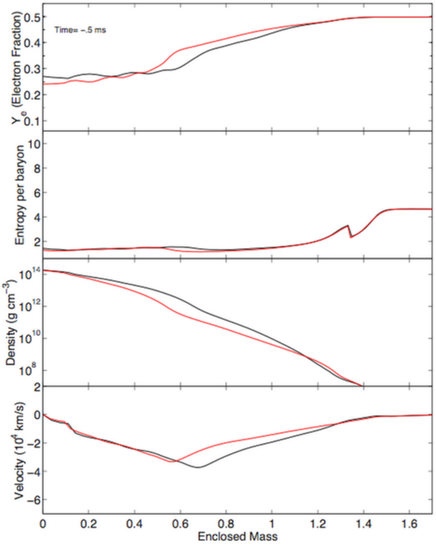
Unfortunately the most sophisticated one-dimensional simulations (with 'best' nuclear physics and neutrino transport) fail in showing successful explosions.



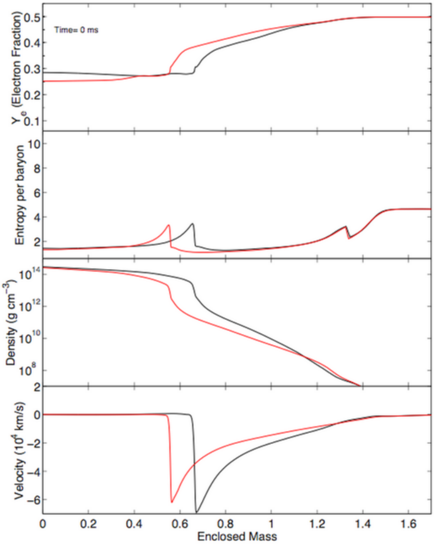
Spherical simulations of shock wave evolution



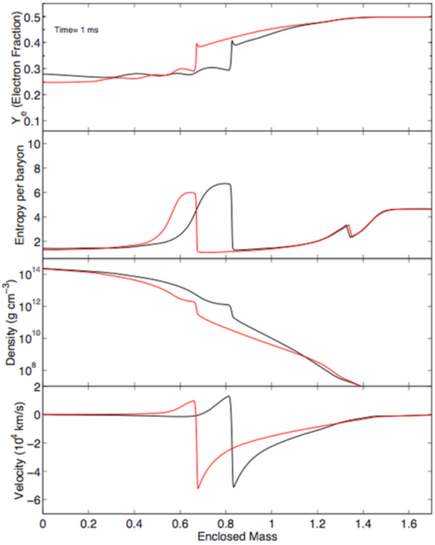
Spherical simulations of shock wave evolution



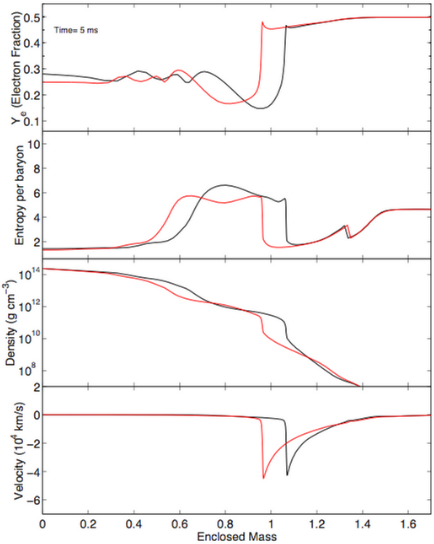
Spherical simulations of shock wave evolution



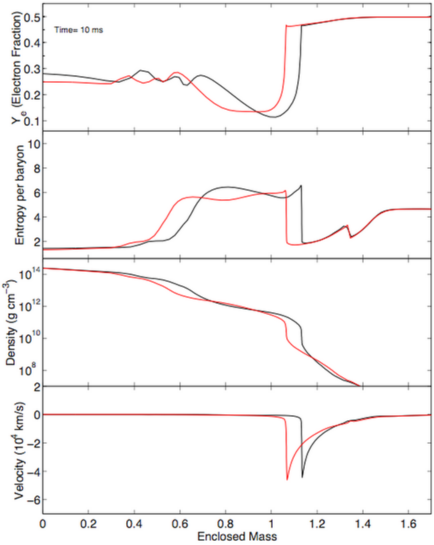
Spherical simulations of shock wave evolution



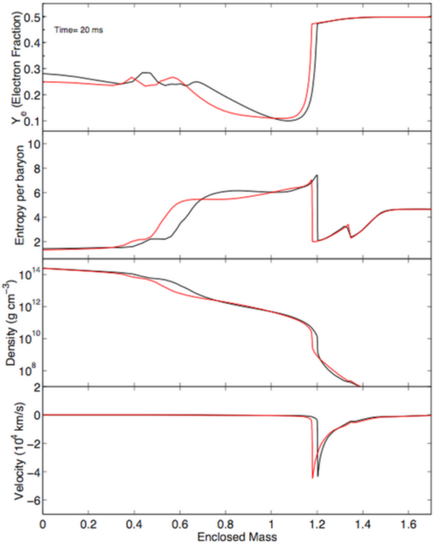
Spherical simulations of shock wave evolution



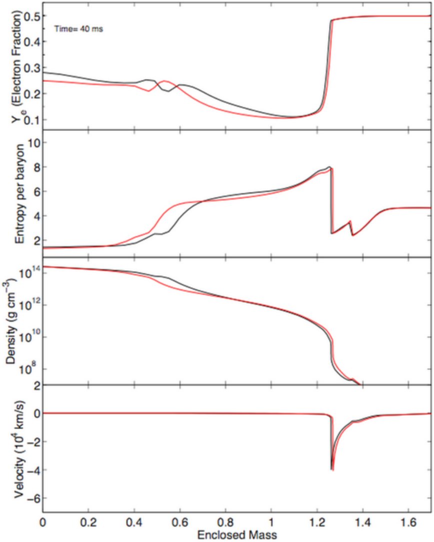
Spherical simulations of shock wave evolution



Spherical simulations of shock wave evolution



Spherical simulations of shock wave evolution



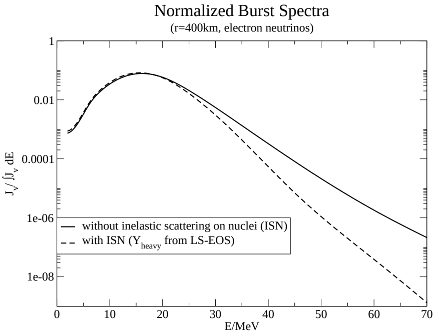
Inelastic ν-nucleus scattering in supernovae

- thermalization of neutrinos during collapse
- preheating of matter before passing of shock
- nucleosynthesis, νp-process
- supernova neutrino signal

This process has been incorporated into supernova simulations only very recently (Garching-Darmstadt collaboration). The effect is negligible, except for the distribution from the neutrino burst.

Supernova neutrino signal

inelastic ν-nucleus scattering adds to the opacity for high-energy neutrinos



B. Müller, H.-Th. Janka, G. Martinez-Pinedo, A. Juodagalvis, J. Sampaio

Which nuclear physics should be improved

- Despite contrary believe, it turned out that electron capture on nuclei is quite important. A more reliable description of the rate is desirable; although a change by a factor 2-3 will probably not matter too much.
- The Equation of State should be improved, where in particular the nuclear compression modulus is important (current accepted value is $K = 9[\frac{dP}{d\rho}]_{\rho_0} = 210 \pm 30$ MeV for symmetric nuclear matter), further the symmetry energy and their temperature and density dependence.
- The nuclear level density and partition function
- Neutrino interactions (opacities) at densities around 10^{13} g cm^{-3}, which are important for shock revival, and at higher densities, which might influence the luminosities.

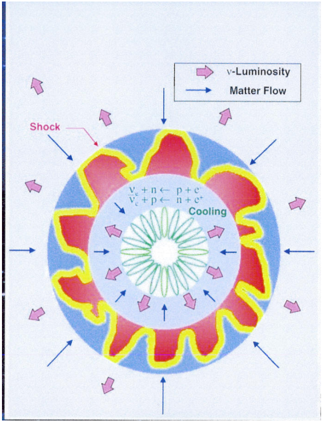
Whatelse is missing?

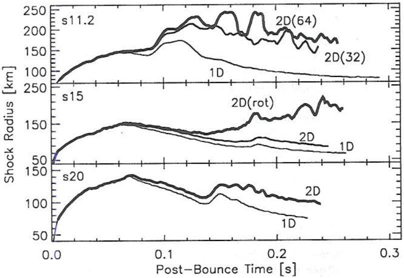
Two effects are currently intensively studied:

- **Multidimensional effects (convection).**
  This should bring neutrinos from deeper (hotter) layers to the shock and increases the energy transfer effectiveness.
- **Magnetic fields**
  The field should help to drive and collimate outflows from the stellar core. Some supernovae show jets and bipolar flows which might point to the importance of magnetic fields.

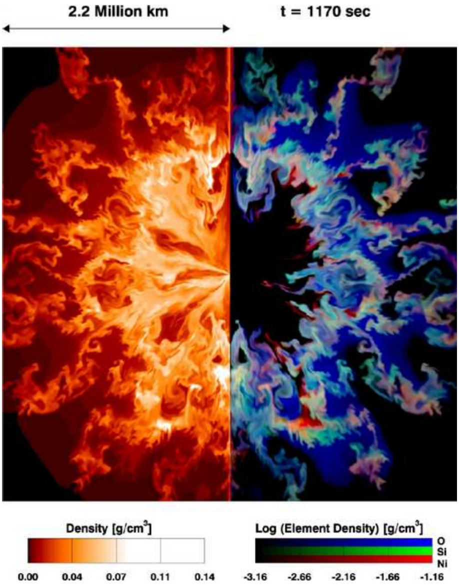
Convection

There exist now two-dimensional simulations (with neutrino transport, but not the latest nuclear physics like electron capture on nuclei). They also fail to explode, but show shocks which move to larger radii.

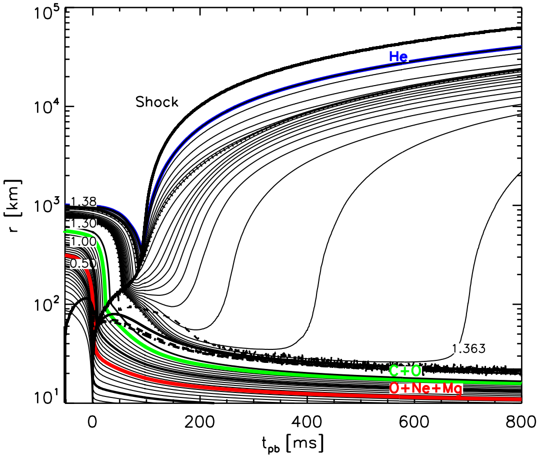




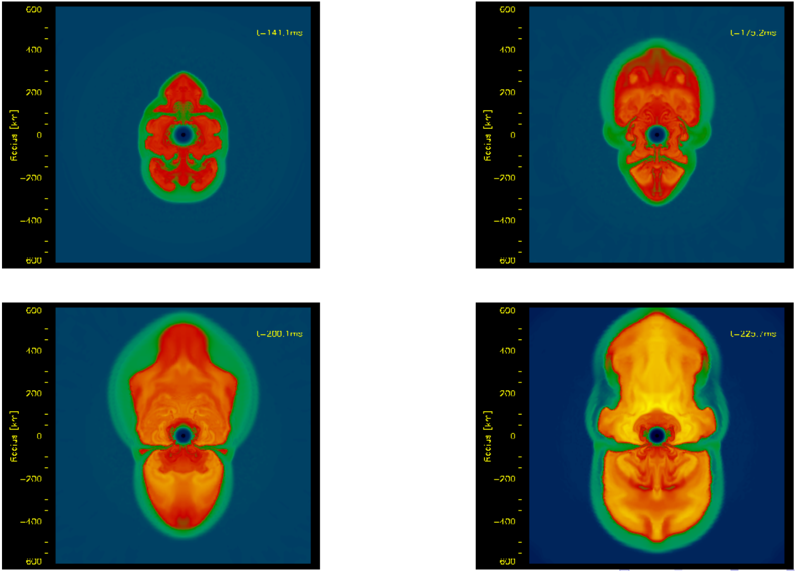
Mixing in the explosion



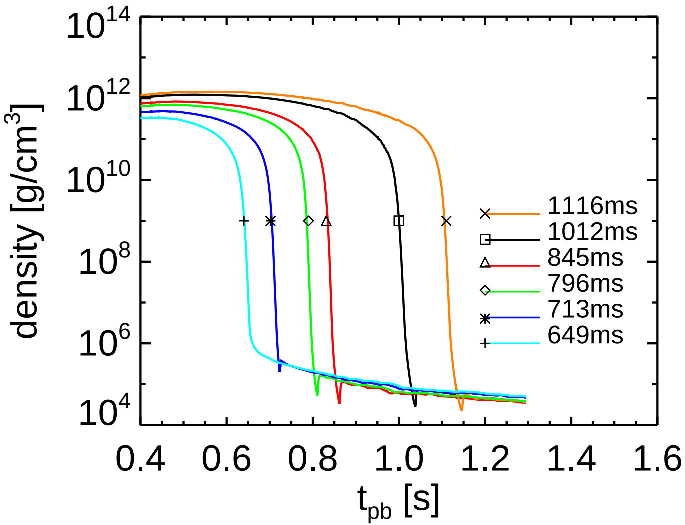
Successful 2-dimensional explosion of $11 M_\odot$ star with ONeMg core (H.-Th. Janka)



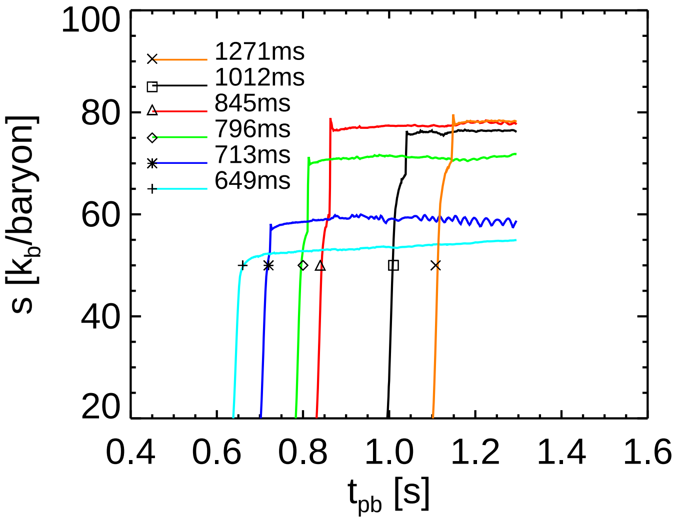
Successful two-dimensional supernova



Properties of ejecta: density profile

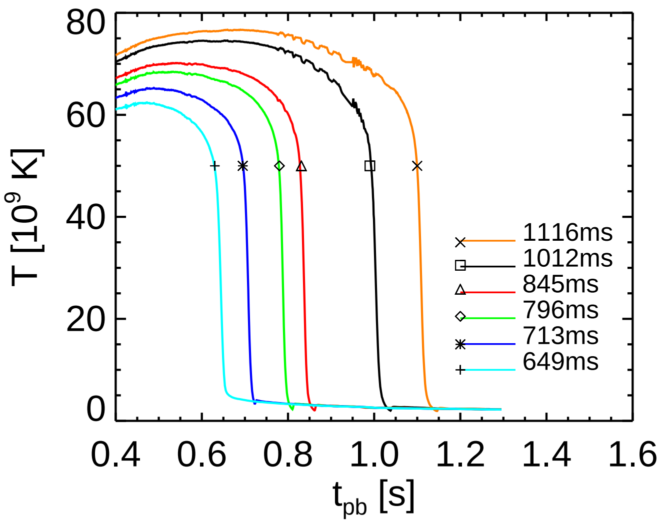


Properties of ejecta: entropy profile

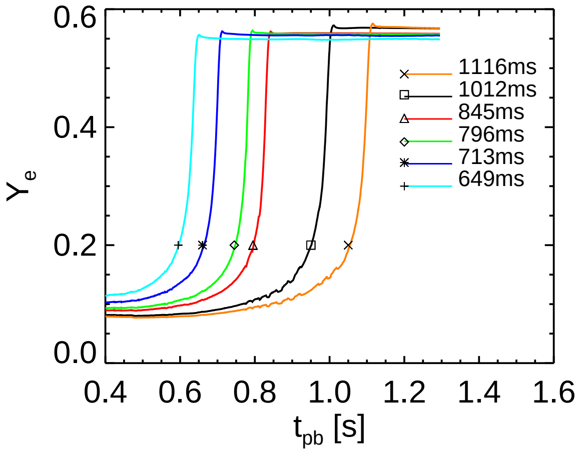


Ejection is adiabatic (constant entropy)

Properties of ejecta: temperature profile



Early-on ejecta are very hot → matter is dissociated into free nucleons

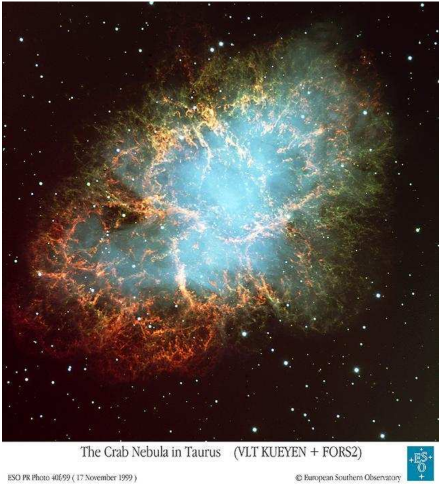


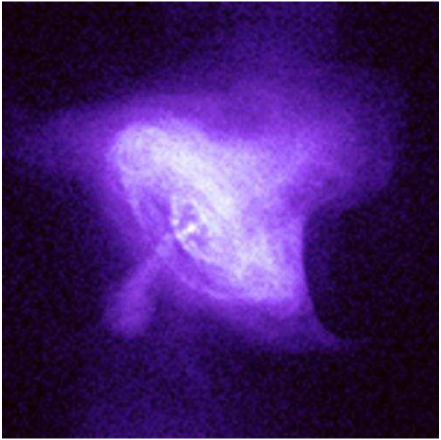
Early ejected matter is protonrich ($Y_e > 0.5$)!! This will have exciting consequences for nucleosynthesis.

Supernova remnants

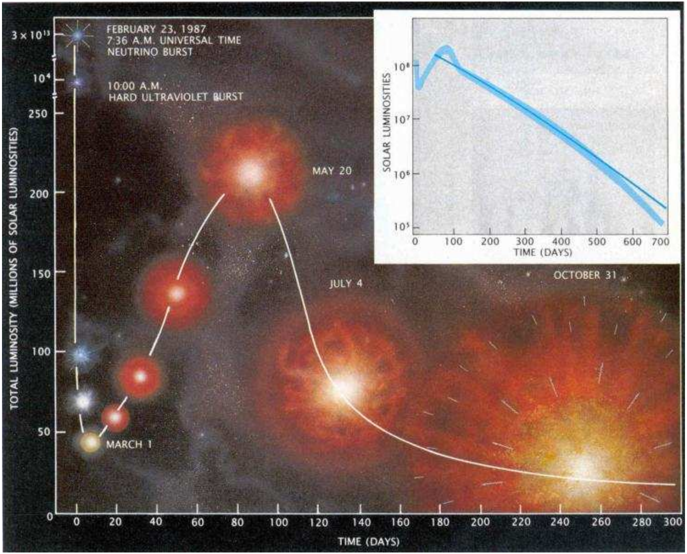
The remnant left over in the explosion depends on the main-sequence mass M_{ms} and on the maximum mass for neutron stars. The later is not quite well known. Most neutron stars, whose masses are well determined (they are in binaries), have masses around 1.4 $M_{\odot}$, however, recent observations might imply masses up to 2.1 $M_{\odot}$.
It is generally assumed that the collapse of stars with $M_{ms} > 20 - 25 M_{\odot}$ leads to a black hole in the center, while stars with $8 M_{\odot} < M_{ms} < 20 - 25 M_{\odot}$ yield a supernova with a neutron star remnant.
It is also possible that accretion during the explosion might put the remnant over the neutron star mass limit. It is speculated that this happened in the case of the SN87A.

Supernova remnants





Light curve



FEBRUARY 23, 1987
7:36 A.M. UNIVERSAL TIME
NEUTRINO BURST
10:00 A.M.
HARD ULTRAVIOLET BURST
MAY 20
JULY 4
OCTOBER 31
MARCH 1
TOTAL LUMINOSITY (MILLIONS OF SOLAR LUMINOSITIES)
TIME (DAYS)
SOLAR LUMINOSITIES
TIME (DAYS)

Supernovae are very, very bright!

Supernovae are observed to outshine an entire galaxy, having luminosities of order $10^{10}\ L_\odot$. They are optically visible and have surface temperatures of the same order as the Sun. Then they must have large surface areas!

Supernovae have radii of order 10^{15} cm.

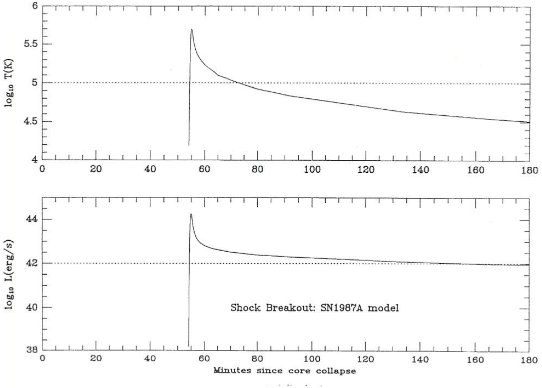
An object which radiates such large luminosities over several weeks ($\approx 10^6$ s) emits about 10^{50} ergs. From Doppler-shifts the expansion velocities are determined as $2000 - 10000$ km/s; this translates into akinetic energy of about 10^{51} ergs. The observed neutrinos of SN87A corresponds to an energy release of about 10^{53} ergs.

The explosion of an initially dense star cannot describe the observed lightcurves, as by the time they expand to become big, they have been cooled by just that expansion. The energy source which powers the lightcurve must be unaffected by the expansion.

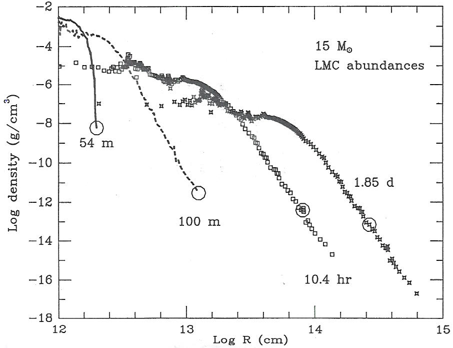
The energy source of the lightcurves is the decay energy released by radioactive nuclei which have lifetimes just long enough to release the energy after the expansion.

Shock emergence

Neutrinos, travelling at the speed of light, reach the surface first, after about 1 minute. The shock, travelling with an average speed of $\bar{v} = 6250$ km/s, needs about 60 minutes to reach the surface, which is at $R \approx 2 \times 10^{12}$ cm then. This leads to a burst in the luminosity.



Supernova envelope expansion

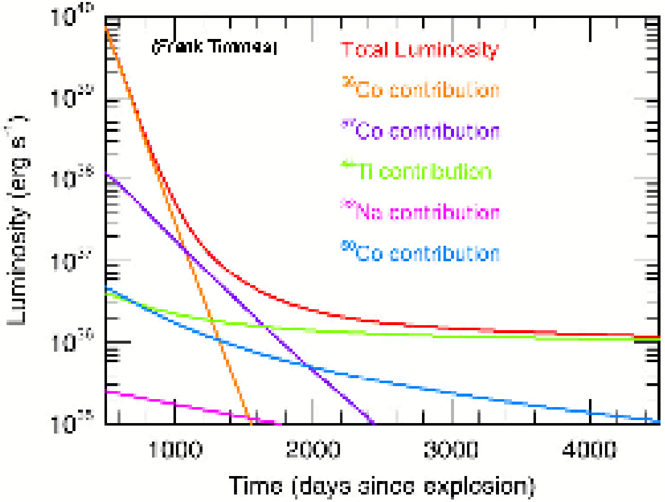


The expansion is adiabatic ($T \sim \frac{1}{R}$), thus the luminosity ($\sim T^4$) decreases. Also, the radius from which photons are emitted (photosphere) moves to larger radii, but inside the matter.

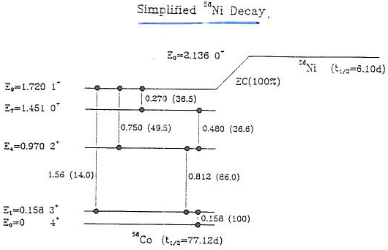
Energy from radioactive decays

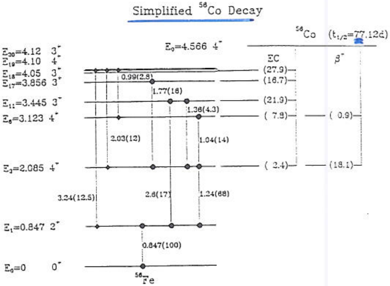
A core-collapse supernova produces about $0.15 - 0.2\ M_{\odot}$ ^{56}Ni. This is made in the outer layers of the star ($Y_e = 0.5$, mainly ^{16}O) when the shock wave passes through and brings this matter into NSE by fast reactions. Supernova also produce other radioactive nuclides (for example ^{57}Ni and ^{44}Ti). ^{44}Ti is only barely made (about $10^{-4}\ M_{\odot}$), but has a lifetime of about 60 years. It dominates the lightcurve of SN87A today.
These radiactive nuclides decay, producing γ radiation in the MeV range. By scattering with electrons, these photons are thermalized and then radiated away as infrared, visible, and ultraviolet light.

Radioactivity powers the lightcurve



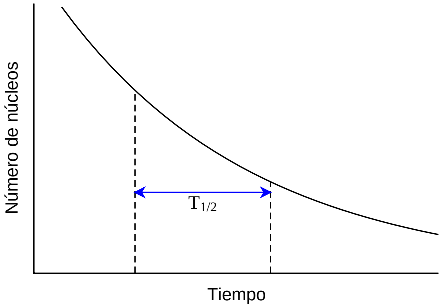
^{56}Ni and ^{56}Co decay schemes

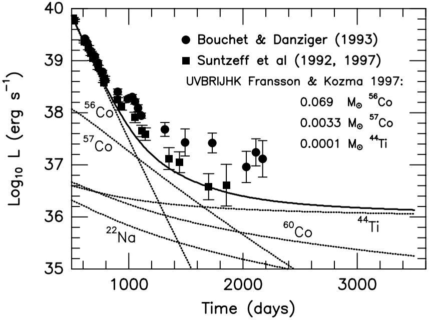




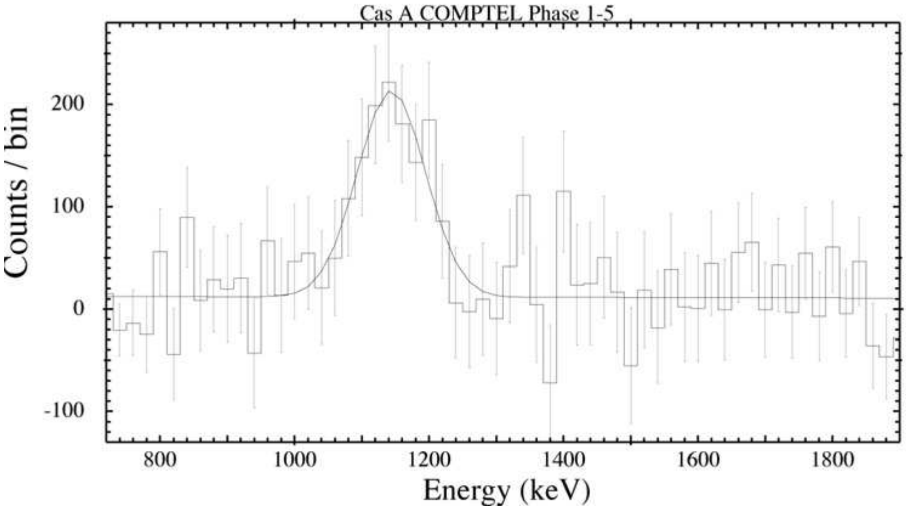
Light curve follows the decay of Nickel.

$$^{56}\mathrm{Ni} \xrightarrow{6\ \mathrm{days}} {}^{56}\mathrm{Co} \xrightarrow{77\ \mathrm{days}} {}^{56}\mathrm{Fe}$$

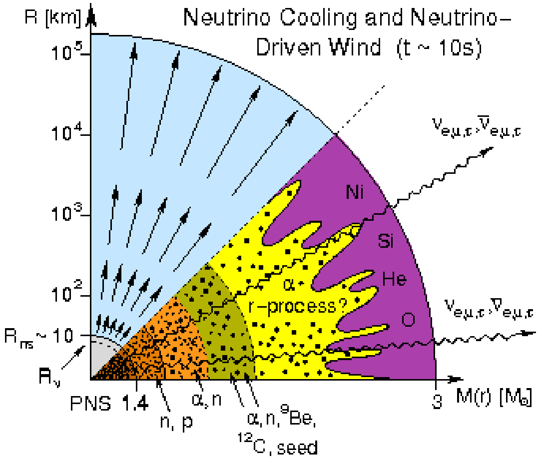




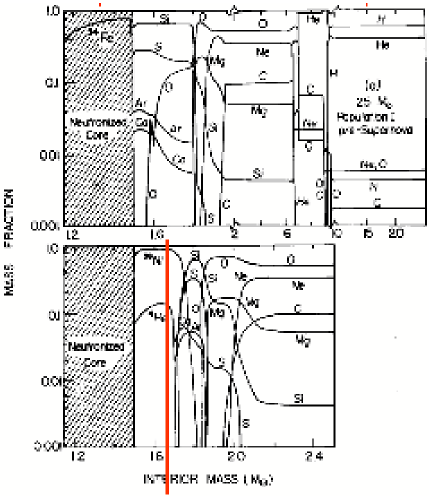
^{44}Ti from CasA



Explosion and explosive nucleosynthesis



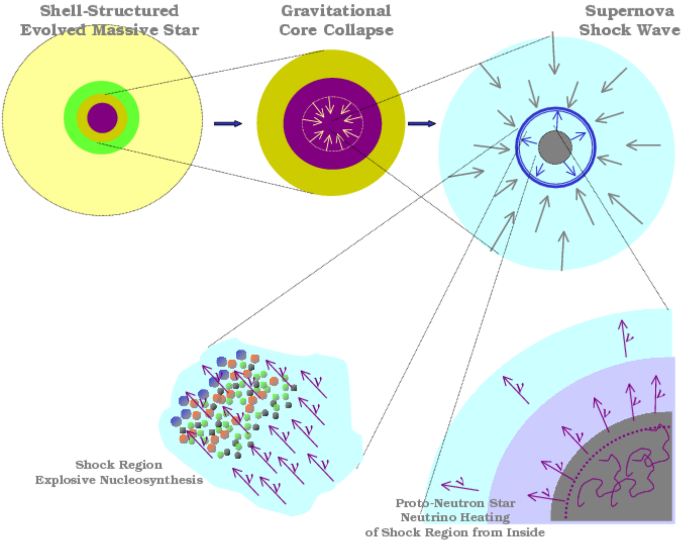
Masscut



mass cut somewhere here

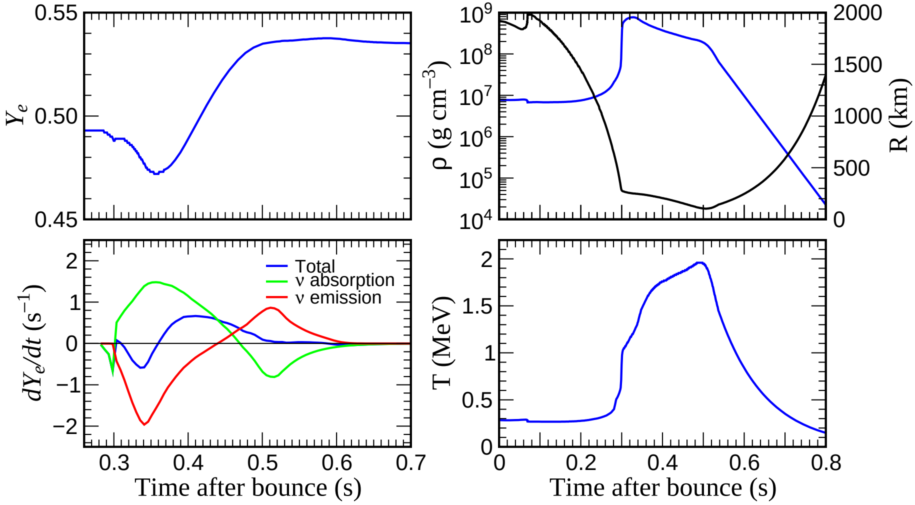
not ejected | ejected

Explosive nucleosynthesis



Neutrino interactions determine Y_e value

M. Liebendörfer *et al*



Why is $Y_e > 0.5$ for early ejecta?

The change of Y_e is determined by various neutrino reactions on free nucleons:

- electron capture (EC): $e^- + p \rightarrow n + \nu$
- electron neutrino capture (NC): $\nu + n \rightarrow e^- + p$
- positron capture (PC): $e^+ + n \rightarrow p + \bar{\nu}$
- anti electron-neutrino capture (AC): $\bar{\nu} + p \rightarrow e^+ + n$

$$\frac{1}{c}\frac{dY_e}{dt} = K \int dE E^2 [h(E+Q)(-EC+NC) \\ +\Theta(E-Q-m_e)h(E-Q)(PC-AC)$$

$$h(x) = x^2 \left[1 - \left(\frac{m_e c^2}{x}\right)^2\right]^{1/2}$$

Step function $\Theta(x)$ considers thresholds.

Various neutrino reaction rates

Considering a Fermi-Dirac distribution for electrons and positrons, the various reaction rates are given by:

$$
\begin{aligned}
EC &= \frac{1}{1+e^{\beta(E+Q-\mu_e)}} n_p (1-f_\nu) \\
NC &= \frac{e^{\beta(E+Q-\mu_e)}}{1+e^{\beta(E+Q-\mu_e)}} \; n_n \; f_\nu \\
PC &= \frac{1}{1+e^{\beta(E-Q+\mu_e)}} \; n_n \; (1-f_{\bar{\nu}}) \\
AC &= \frac{e^{\beta(E-Q+\mu_e)}}{1+e^{\beta(E-Q+\mu_e)}} \; n_p \; f_{\bar{\nu}}
\end{aligned}
$$

n_n, n_p are neutron and proton number densities; f_ν defines neutrino distribution (not necessarily FD). E is the neutrino energy.

Competition of E, Q, μ_e

The balance of the 4 reactions is determined by n_n, n_p and the exponential $\exp(\beta(E \pm Q \mp \mu_e))$. The energy integral depends on the competition between the neutrino energy E (distributed like f in NC and AC), the neutron-proton mass difference $Q = 1.29$ MeV and the electron chemical potential μ_e. Depending on the various conditions, any of these 3 quantities can dominate and hence determine the balance of protons and neutrons.

Very early ejecta

At very early times ejecta are close to the neutron star and all 4 reactions are active. The distributions f are non-trivial functions of the accretion rate, the distance from the neutrinosphere, and the local weak interactions. It becomes important how able the matter is to accept electrons and neutrinos produced in the processes (Pauli blocking). In particular, if electrons are degenerate, μ_e can dominate over the neutrino energy; hence $\exp(\beta(E + Q - \mu_e))$ is small, while $\exp(\beta(E - Q + \mu_e))$ is large. Thus, neutrino absorption (NC) and positron capture (PC) are suppressed. The electron fraction Y_e decreases due to dominance of electron capture (EC) and anti-neutrino absorption (AC). A balance can only be achieved when the ratio between proton and neutron number densities is decreased to compensate the dominance of EC and AC. This leads to $n_p < n_n$ and matter with $Y_e < 0.5$.

Protonrich matter in supernova ejecta

For high-energy neutrinos, i.e. if the neutrino energy E is sufficiently larger than $|Q - \mu_e|$, $\exp(\beta E)$ is large and the neutrino absorption reactions (NC and AC) dominate over their inverse reactions. Thus, if the abundance of such high-energy neutrinos is large, equilibrium is determined by the competition of neutrino and anti-neutrino absorption, and depends on the unspecified neutrino distributions. The rate of change of Y_e is given by

$$\frac{Y_e}{dt} \approx \lambda_{\nu n} - Y_e(\lambda_{\nu n} + \lambda_{\bar{\nu} p})$$

Protonrich matter in supernova ejecta

The neutrino and antineutrino absorption rates are given by

$$\lambda_{\nu n} = \frac{L_\nu}{4\pi r^2}\sigma_0 \left(E_\nu + 2\Delta + \frac{\Delta^2}{\langle E_\nu \rangle} \right)$$
$$\lambda_{\bar{\nu} p} = \frac{L_{\bar{\nu}}}{4\pi r^2}\sigma_0 \left(E_{\bar{\nu}} - 2\Delta + \frac{\Delta^2}{\langle E_{\bar{\nu}} \rangle} \right)$$

with $\Delta = 1.29$ MeV, $\sigma_0 = 9.8385 \times 10^{-44}$ cm^2 MeV^{-2}. The luminosities at the neutrinosphere L are defined in MeV s^{-1} and E is the average neutrino and anti-neutrino energy in MeV.
If we assume $L_\nu \approx L_{\bar{\nu}}$, we find

$$\frac{dY_e}{dt} = \frac{L_\nu}{4\pi r^2}\sigma_0 \left\{ E_\nu + 2\Delta \frac{\Delta^2}{\langle E_\nu \rangle} - Y_e(E_\nu + E_{\bar{\nu}}) - Y_e \Delta^2 \left(\frac{1}{\langle E_\nu \rangle} + \frac{1}{\langle E_{\bar{\nu}} \rangle} \right) \right\}$$

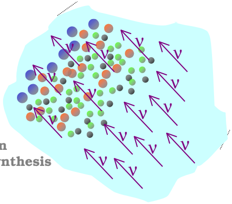
Protonrich matter in supernova ejecta

If we assume $Y_e = 0.5$ as a typical value, we find

$$\begin{aligned}\frac{dY_e}{dt}\Big|_{Y_e=0.5} &= \frac{L_\nu}{4\pi r^2}\sigma_0 \left\{2\Delta - \frac{1}{2}(E_{\bar{\nu}} - E_\nu) + \frac{\Delta^2}{2}\left(\frac{1}{\langle E_\nu \rangle} - \frac{1}{\langle E_{\bar{\nu}} \rangle}\right)\right\} \\ &\approx \frac{L_\nu}{4\pi r^2}\sigma_0 \left\{2\Delta - \frac{1}{2}(E_{\bar{\nu}} - E_\nu)\right\}\end{aligned}$$

Thus $\frac{dY_e}{dt} > 0$, if $4\Delta > (E_{\bar{\nu}} - E_\nu)$. We find that matter gets protonrich, if the average anti-neutrino energy exceeds the average neutrino energy by less than about 5 MeV. This is fulfilled for the early ejecta!

Consistent supernova nucleosynthesis



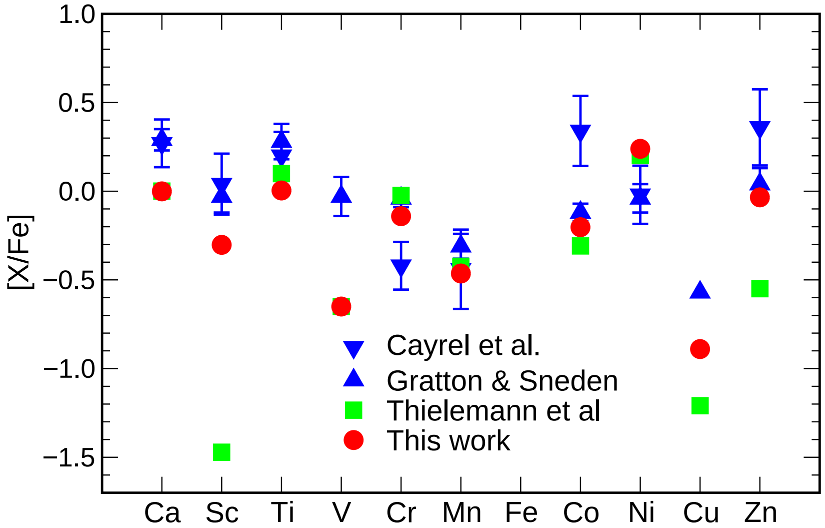
- Consistent treatment of supernova dynamics coupled with a nuclear network.
- Essential neutrino reactions in the shock heated region

$$\nu_e + n \rightleftarrows p + e^-$$
$$\bar{\nu}_e + p \rightleftarrows n + e^+$$

- early ($\sim$ 1 s): matter protonrich $\rightarrow$ νp-process
- later: matter neutronrich $\rightarrow$ r-process

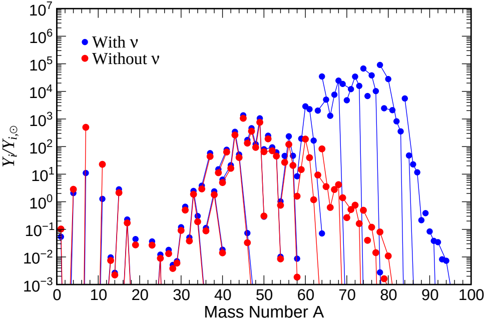
Comparison with observations.

Carla Fröhlich, G. Martinez-Pinedo *et al.*



The νp-process: basic idea

- Protonrich matter is ejected under the influence of neutrino reactions
- Nuclei form at distance where a substantial antineutrino flux is present

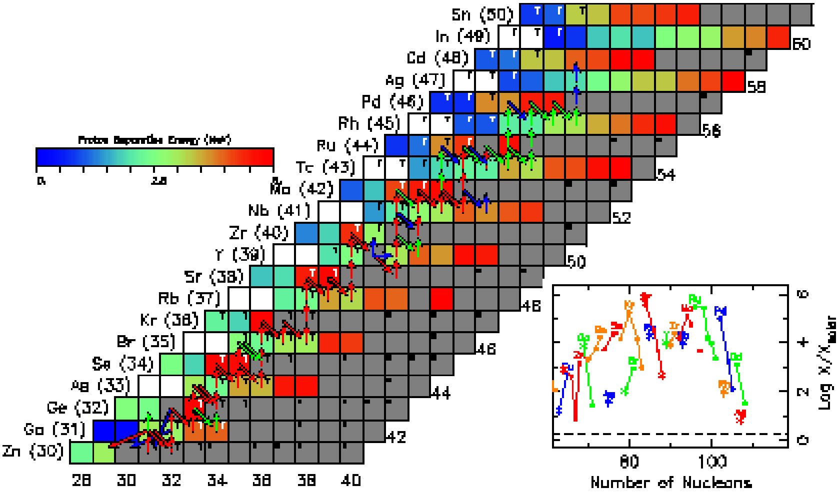


Antineutrinos help to bridge long waiting points via (n,p) reactions

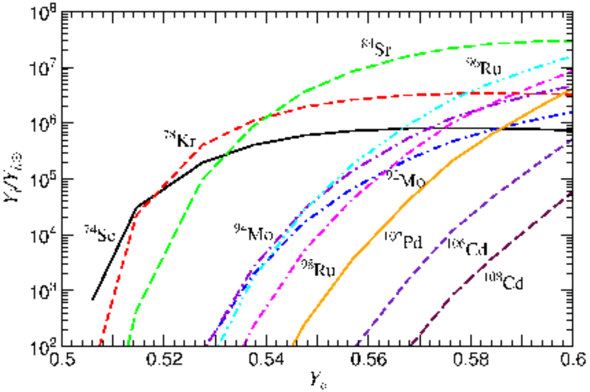
$$\bar{\nu}_e + p \rightarrow e^+ + n; \quad n + {}^{64}\text{Ge} \rightarrow {}^{64}\text{Ga} + p; \quad {}^{64}\text{Ga} + p \rightarrow {}^{65}\text{Ge}; \ldots$$

C. Fröhlich, G. Martinez-Pinedo, et al., PRL 96 (2006) 142502

Matter flow in the νp process



Pruet, Woosley, Janka, Rampp

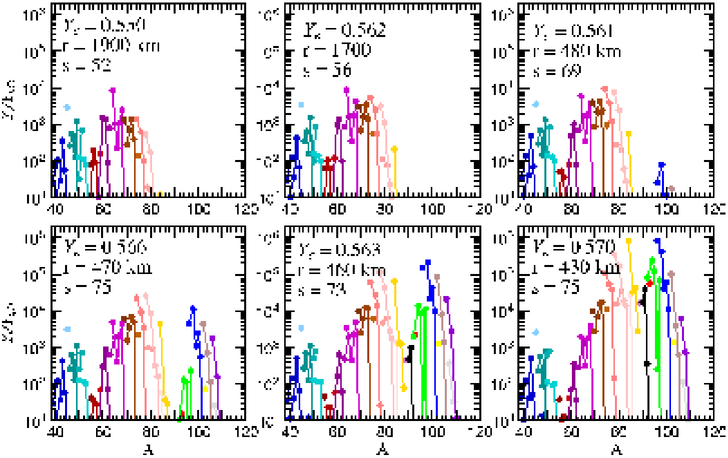


Y_e: electron-to-nucleon ratio

the larger Y_e, the more protons exist and can be transformed into neutrons by antineutrino capture

The νp-process: abundance yields

νp-process nucleosynthesis for realistic supernova trajectories (from Hans-Thomas Janka)

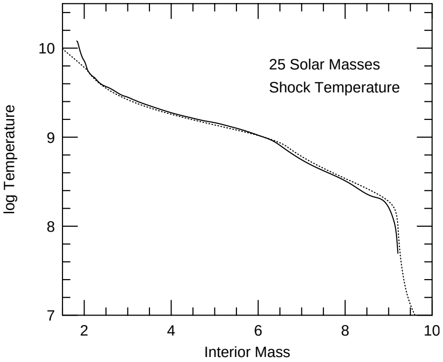


Explosive nucleosynthesis

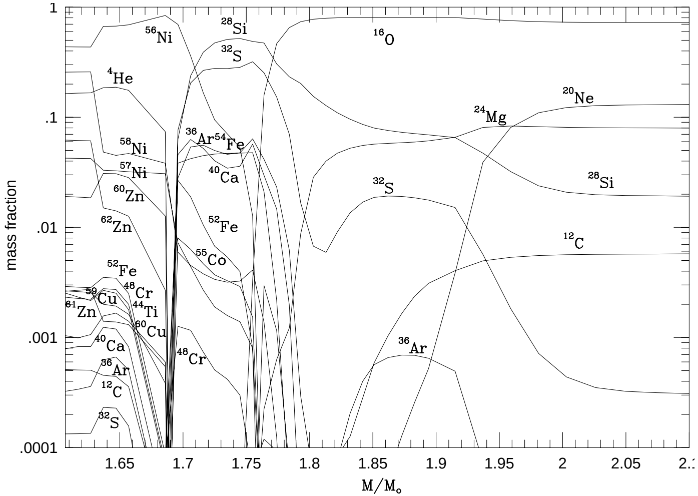
- Determined mainly by temperature.
- Radiation dominated inner zone $(E/V = aT^4)$

$$\frac{4}{3}\pi r^3 aT^4 = E_{kin} \approx 10^{51} \text{ erg}$$

- Interior 3000 km, $T \sim 5$ billion K. Matter converted to NSE to give mainly Iron group elements.
- 5000 km, temperature below 4 billion K.
- 13000 km, below 2 billion K.

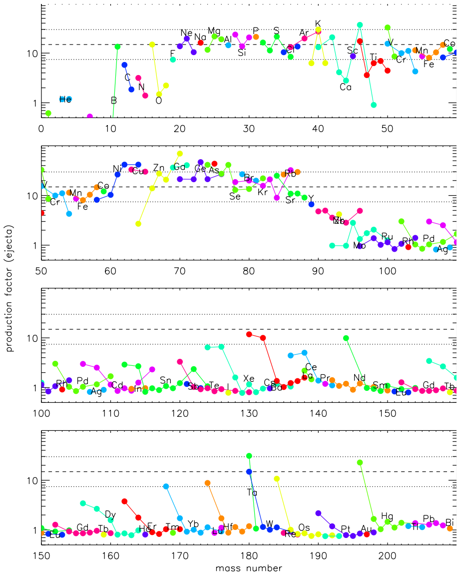


Explosive nucleosynthesis

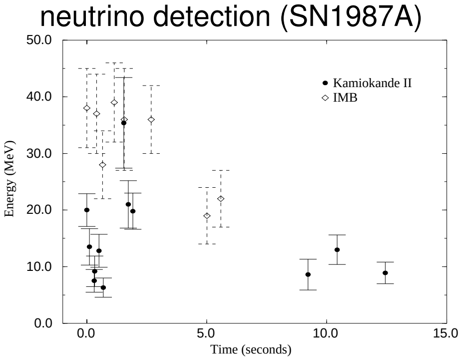


- Inner layers, α-rich freeze out.
- Outer layers, explosive Si burning.

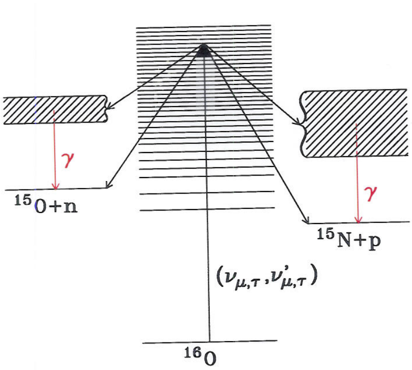
Explosive nucleosynthesis

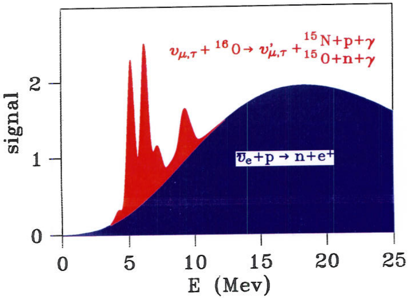


Supernova neutrino detection



The Kamioka and IMB detectors are water Cerenkov detectors. Observed have been $\bar{\nu}_e$ neutrinos via there interaction on protons (in the water molecule). The detection of the other neutrino types is the main goal for the next nearby supernova to test the predicted neutrino hierarchy.



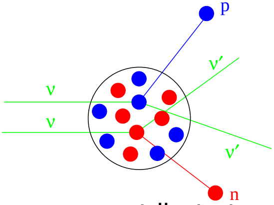


SNO will also observe ν_x neutrinos by dissociation of deuterons, as well as ν_e neutrinos.

Neutrino nucleosynthesis

When the neutrinos pass through the outer shells of the star, they can interact with nuclei exciting them above particle thresholds. Examples are:

$^{12}\mathrm{C}(\nu,\nu'\mathrm{p})^{11}\mathrm{B}$; $^{12}\mathrm{C}(\nu,\nu'\mathrm{n})^{11}\mathrm{C}$

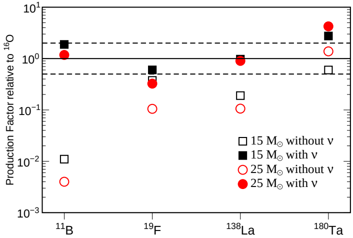


In this way, neutrino-induced reactions can contribute to nucleosynthesis. As a rule of thumb, neutrino nucleosynthesis is important if the abundance ratio of parent and daughter is 1000 or larger. Simulations of neutrino nucleosynthesis abundances require reliable nucleosynthesis calculations during hydrostatic burning (including s-process) and during the shock passage.

Neutrino nucleosynthesis

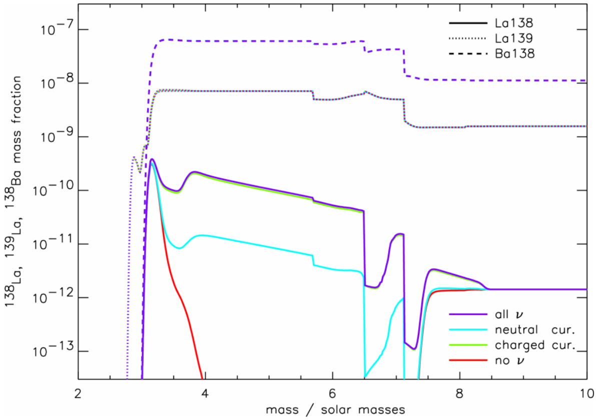
A. Heger *et al*, PLB 606 (2005) 258

Product	Parent	Reaction
^{11}B	^{12}C	$(\nu,\nu' n), (\nu,\nu' p)$
^{19}F	^{20}Ne	$(\nu,\nu' n), (\nu,\nu' p)$
^{138}La	^{138}Ba	(ν_e, e^-)
	^{139}La	$(\nu,\nu' n)$
^{180}Ta	^{180}Hf	(ν_e, e^-)
	^{181}Ta	$(\nu,\nu' n)$



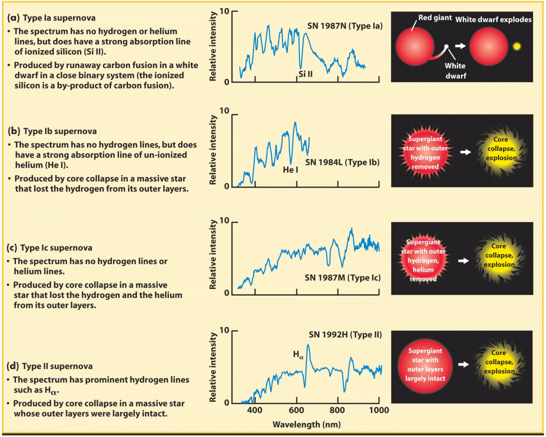
Neutrino nucleosynthesis is sensitive to those neutrino types, which have not been observed from SN87A.

Making ^{138}La by neutrino nucleosynthesis

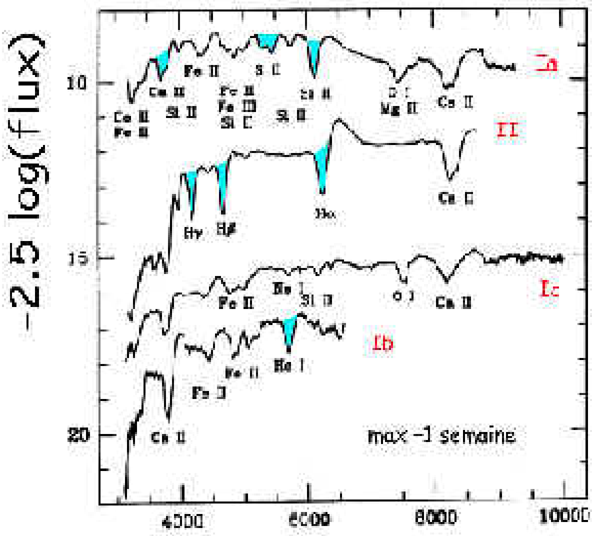


^{138}La is being made by charged-current (ν_e, e^-) reactions on ^{138}Ba, which in turn has been produced by the s-process before. The relevant GT_- cross section has been recently measured in Osaka, including researchers from TUD.

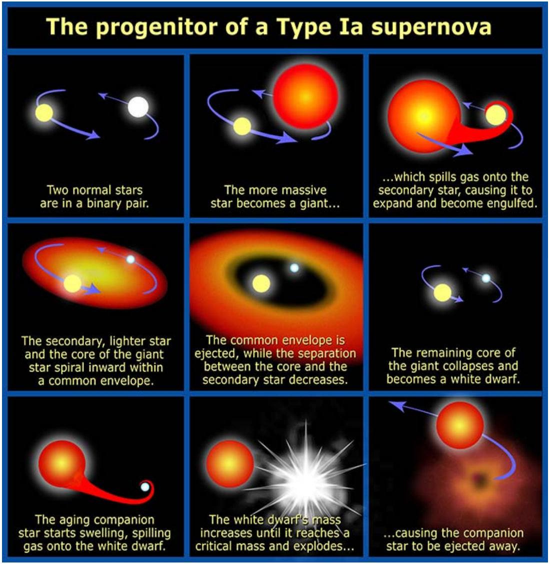
Supernova classes



Spectra of different supernova classes



Progenitor evolution of a type Ia supernova



Type Ia supernovae: general properties

- there are no hydrogen lines in the spectra, but prominent Si lines
- the spectra are dominated by intermediate-mass elements (early: Si, Ca, Mg, S, O; later: Fe, Co)
- typical velocities of the ejecta are a few 10^4 km/s
- there are no neutron star remnants
- they produce a few tenth of $M_\odot$ of ^{56}Ni which powers the lightcurve
- there is not much variations between different type Ia's

Correlations and standard candles

Inhomogeneities among type Ia observables are strongly interco rrelated. The most important one is the correlation between the width of the light curve around maximum and the peak brightness (Phillips relation). This makes type Ia's to **standard candles**.