Lecture 14

Neutrino-Powered Explosions, Rotation, and Mixing

Baade and Zwicky, *Proceedings of the National Academy of Sciences,* (1934)

"With all reserve we advance the view that a supernova represents the transition of an ordinary star into a neutron star consisting mainly of neutrons. Such a star may possess a very small radius and an extremely high density. As neutrons can be packed much more closely than ordinary nuclei and electrons, the gravitational packing energy in a cold neutron star may become very large, and under certain conditions, may far exceed the ordinary nuclear packing fractions ... "

Chadwick discovered the neutron in 1932 though the idea of a neutral massive particle had been around since Rutherford, 1920.

For the next 30 years little progress was made though there were speculations:

Hoyle (1946) - *supernovae are due to a rotational bounce!!*

Hoyle and Fowler (1960) – *Type I supernovae are due to* the explosions of white dwarf stars

Fowler and Hoyle (1964) – *other supernovae are due to thermonuclear burning in massive stars – aided by rotation and magnetic fields*

The explosion is mediated by neutrino energy transport ....

THE HYDRODYNAMIC BEHAVIOR OF SUPERNOVAE EXPLOSIONS*

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Received June 29, 1965

ABSTRACT

We regard the release of gravitational energy attending a dynamic change in configuration to be the primary energy source in supernovae explosions. Although we were initially inspired by and agree in detail with the mechanism for initiating gravitational instability proposed by Burbidge, Burbidge, Fowler, and Hoyle, we find that the dynamical implosion is so violent that an energy many times greater than the available thermonuclear energy is released from the star's core and transferred to the star's mantle in a supernova explosion. The energy released corresponds to the change in gravitational potential of the unstable imploding core; the transfer of energy takes place by the emission and deposition of neutrinos.

Colgate and White, (1966), *ApJ*, **143,** 626

see also
Arnett, (1966), *Canadian J Phys*, **44,** 2553
Wilson, (1971), *ApJ*, **163**, 209

Preliminary: *The neutrino emission of a young neutron star?*

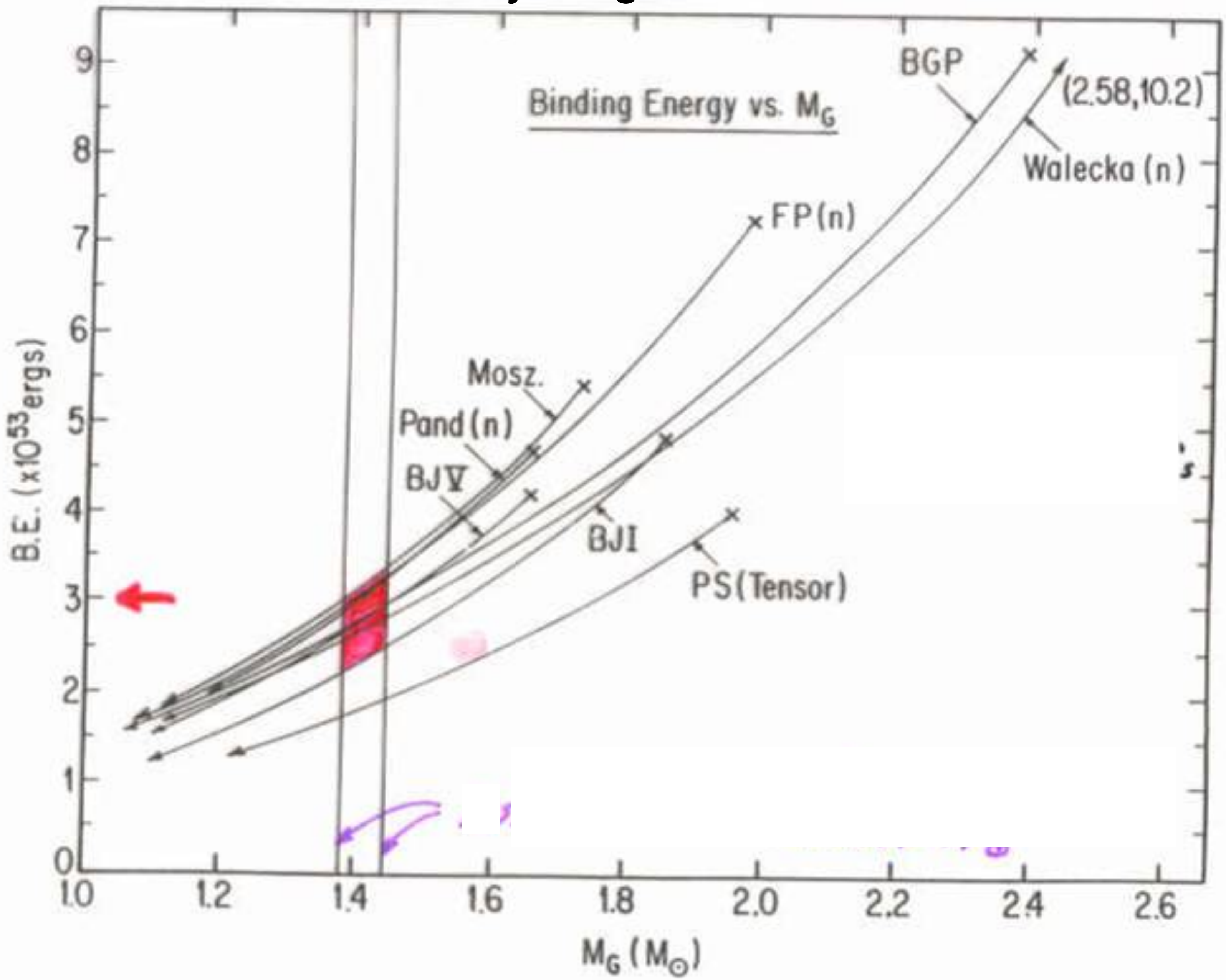


Table 1 Most important neutrino processes in supernova and proto-neutron star matter.

Process	Reaction[a]
Beta-processes (direct URCA processes)	
electron and ν_e absorption by nuclei	$e^- + (A,Z) \longleftrightarrow (A,Z-1) + \nu_e$
electron and ν_e captures by nucleons	$e^- + p \longleftrightarrow n + \nu_e$
positron and $\bar{\nu}_e$ captures by nucleons	$e^+ + n \longleftrightarrow p + \bar{\nu}_e$
"Thermal" pair production and annihilation processes	
Nucleon-nucleon bremsstrahlung	$N + N \longleftrightarrow N + N + \nu + \bar{\nu}$
Electron-position pair process	$e^- + e^+ \longleftrightarrow \nu + \bar{\nu}$
Plasmon pair-neutrino process	$\widetilde{\gamma} \longleftrightarrow \nu + \bar{\nu}$
Reactions between neutrinos	
Neutrino-pair annihilation	$\nu_e + \bar{\nu}_e \longleftrightarrow \nu_x + \bar{\nu}_x$
Neutrino scattering	$\nu_x + \{\nu_e, \bar{\nu}_e\} \longleftrightarrow \nu_x + \{\nu_e, \bar{\nu}_e\}$
Scattering processes with medium particles	
Neutrino scattering with nuclei	$\nu + (A,Z) \longleftrightarrow \nu + (A,Z)$
Neutrino scattering with nucleons	$\nu + N \longleftrightarrow \nu + N$
Neutrino scattering with electrons and positrons	$\nu + e^{\pm} \longleftrightarrow \nu + e^{\pm}$

[a] N means nucleons, i.e., either n or p, $\nu \in \{\nu_e, \bar{\nu}_e, \nu_\mu, \bar{\nu}_\mu, \nu_\tau, \bar{\nu}_\tau\}$, $\nu_x \in \{\nu_\mu, \bar{\nu}_\mu, \nu_\tau, \bar{\nu}_\tau\}$

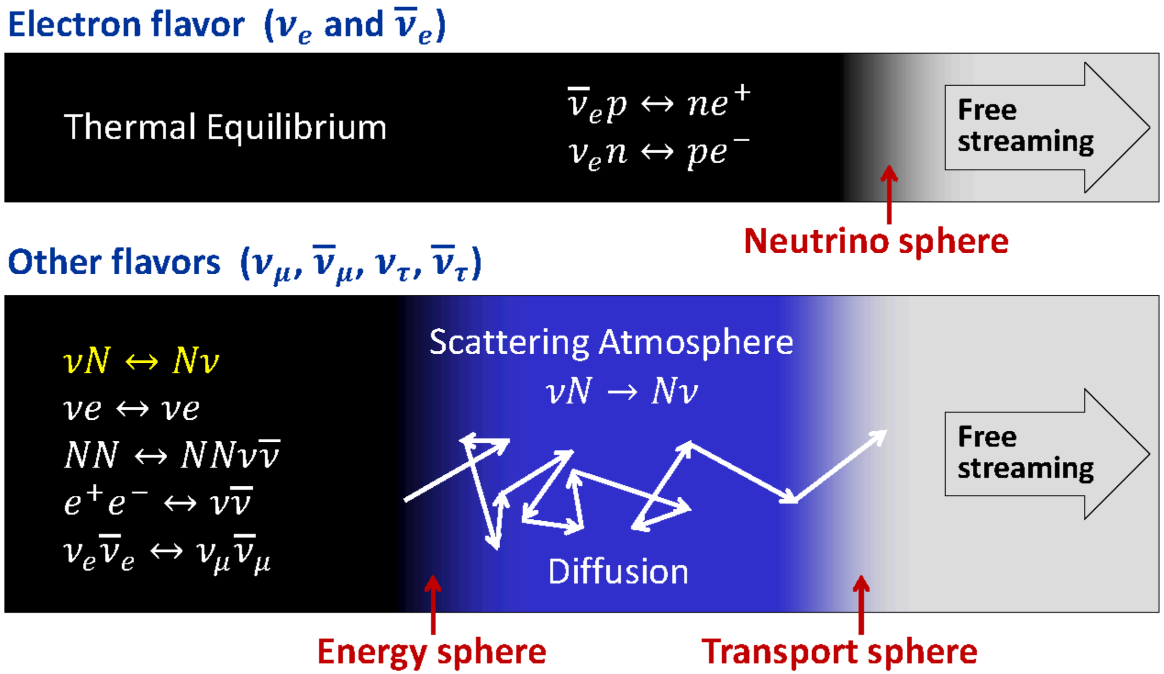
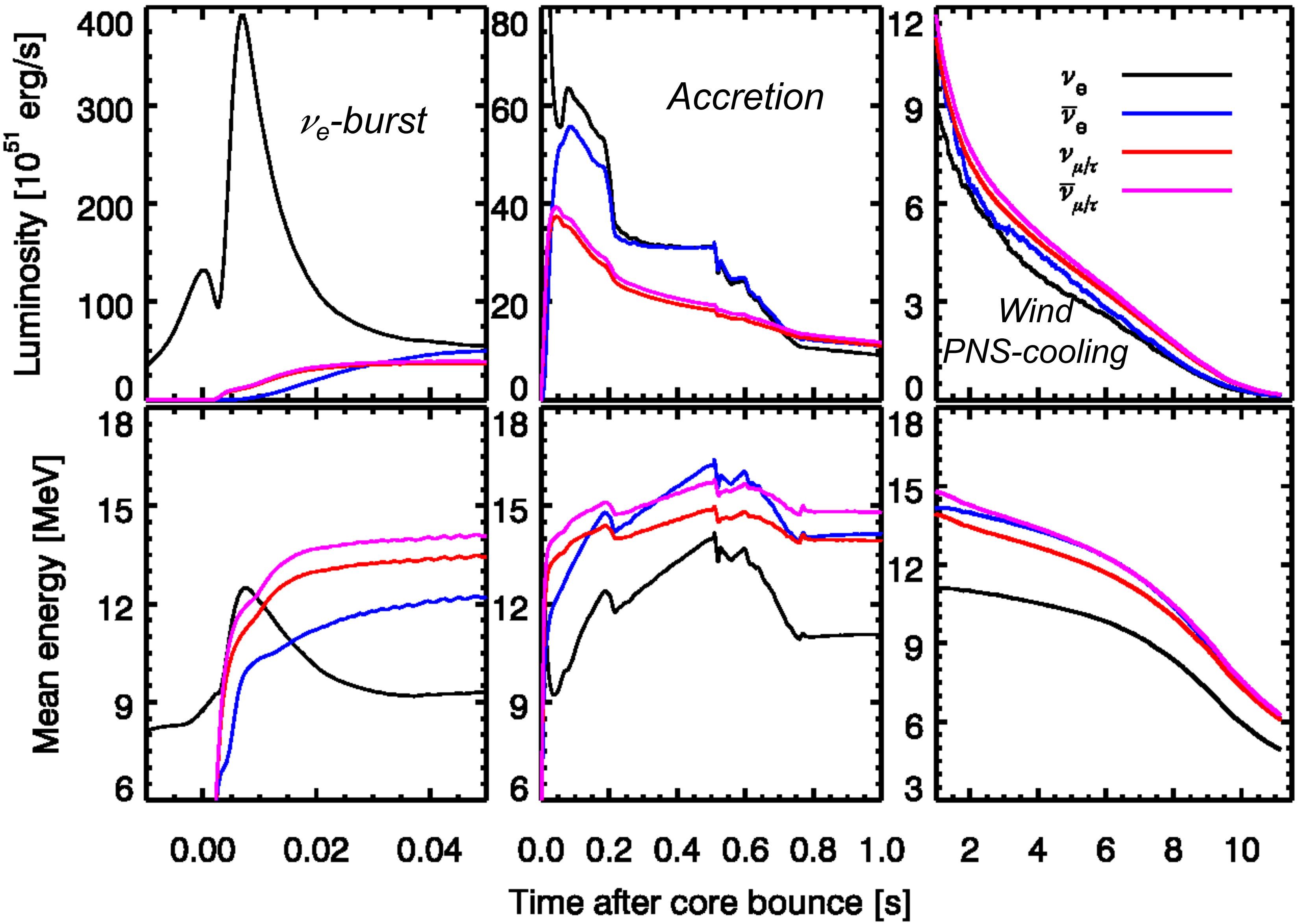
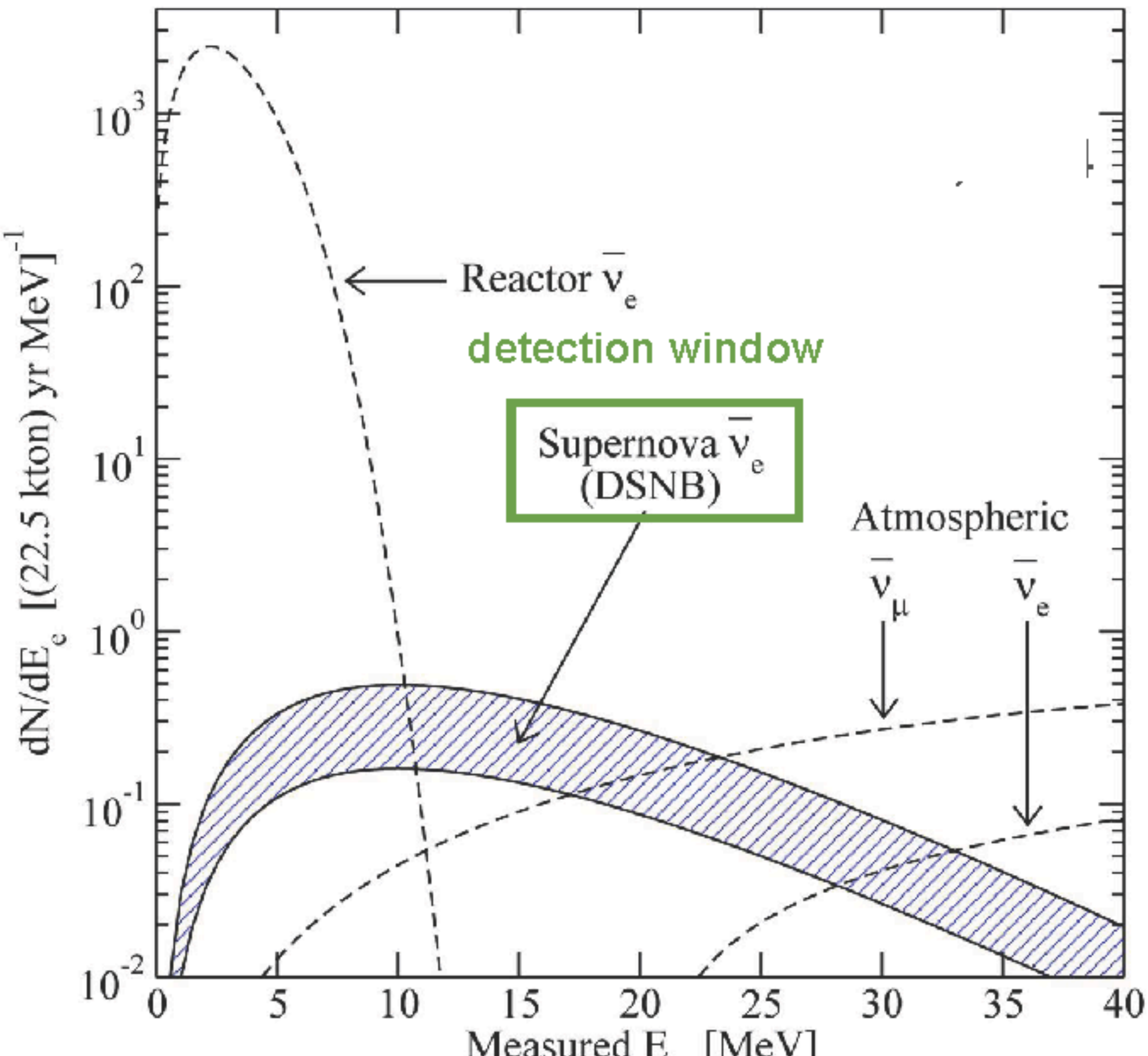


Fig. 4 Sketch of the transport properties of electron-flavor neutrinos and antineutrinos (*upper part*) compared to heavy-lepton neutrinos (*lower part*). In the supernova core ν_e and $\bar{\nu}_e$ interact with the stellar medium by charged-current absorption and emission reactions, which provide a major contribution to their opacities and lead to a strong energetic coupling up to the location of their neutrinospheres, outside of which both chemical equilibrium between neutrinos and stellar matter (indicated by the black region) and diffusion cannot be maintained. In contrast, heavy-lepton neutrinos are energetically less tightly coupled to the stellar plasma, mainly by pair creation reactions like nucleon bremsstrahlung, electron-position annihilation and $\nu_e\bar{\nu}_e$ annihilation. The total opacity, however, is determined mostly by neutrino-nucleon scatterings, whose small energy exchange per scattering does not allow for an efficient energetic coupling. Therefore heavy-lepton neutrinos fall out of thermal equilibrium at an energy sphere that is considerably deeper inside the nascent neutron star than the transport sphere, where the transition from diffusion to free streaming sets in. The blue band indicates the scattering atmosphere where the heavy-lepton neutrinos still collide frequently with neutron and protons and lose some of their energy, but cannot reach equilibrium with the background medium any longer. (Figure adapted from Raffelt, 2012, courtesy of Georg Raffelt)



Luminosity [10^{51} erg/s]
Mean energy [MeV]
Time after core bounce [s]
ν_e-burst
Accretion
Wind
PNS-cooling
ν_e
$\bar{\nu}_e$
$\nu_{\mu/\tau}$
$\bar{\nu}_{\mu/\tau}$
0.00
0.02
0.04
0.0
0.2
0.4
0.6
0.8
1.0
2
4
6
8
10



From talk by Irene Tamborra, MPI, Munich, April 10, 2013

Cosmological Anti-Neutrino Flux

Ando, 2004, *ApJ*, **607**, 20

TABLE 2

FLUX AND EVENT RATE OF SUPERNOVA RELIC NEUTRINOS

super-Kamiokande per year

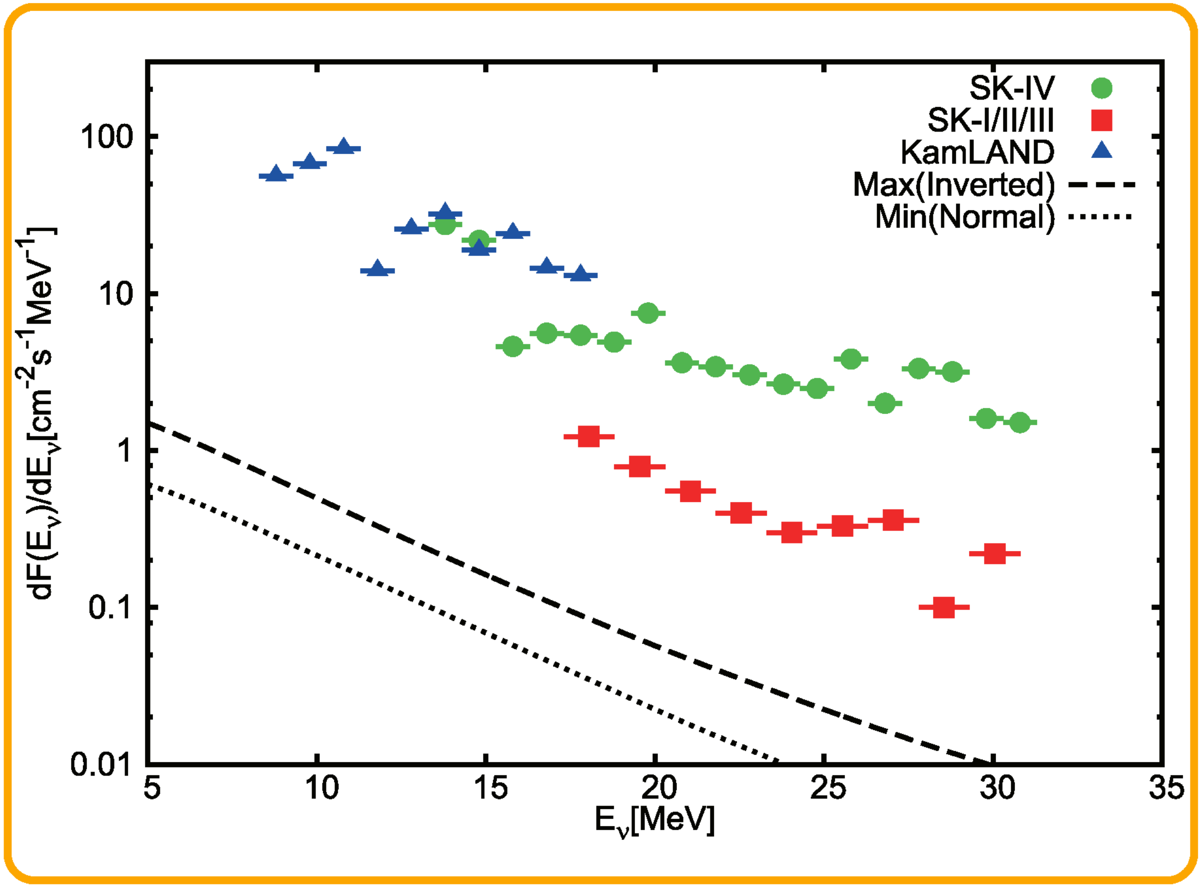
Model	Redshift Range	Flux (cm^{-2} s^{-1}) Total	$E_\nu > 11.3$ MeV	$E_\nu > 19.3$ MeV	Event Rate [(22.5 kton yr)$^{-1}$] $E_e > 10$ MeV	$E_e > 18$ MeV
		Normal Mass Hierarchy				
LL...................	Total	11.7	2.3	0.46	2.3	1.0
	$0 < z < 1$[a]	4.1 (35.3)	1.6 (70.9)	0.39 (85.2)	1.7 (77.5)	0.9 (87.5)
	$1 < z < 2$[a]	4.9 (42.0)	0.6 (26.3)	0.06 (14.0)	0.5 (20.6)	0.1 (11.9)
	$2 < z < 3$[a]	1.8 (15.1)	0.1 (2.5)	0.0 (0.7)	0.0 (1.7)	0.0 (0.5)
	$3 < z < 4$[a]	0.6 (5.3)	0.0 (0.2)	0.0 (0.0)	0.0 (0.1)	0.0 (0.0)
	$4 < z < 5$[a]	0.2 (2.1)	0.0 (0.0)	0.0 (0.0)	0.0 (0.0)	0.0 (0.0)
TBP	Total	16.1	1.3	0.14	0.97	0.25
KRJ	Total	12.7	2.0	0.28	1.7	0.53

LL = Livermore group (1998); TBP = Thompson, Burrows and Pinto (2003); KRJ = Keil, Raffelt, and Janka (2003)

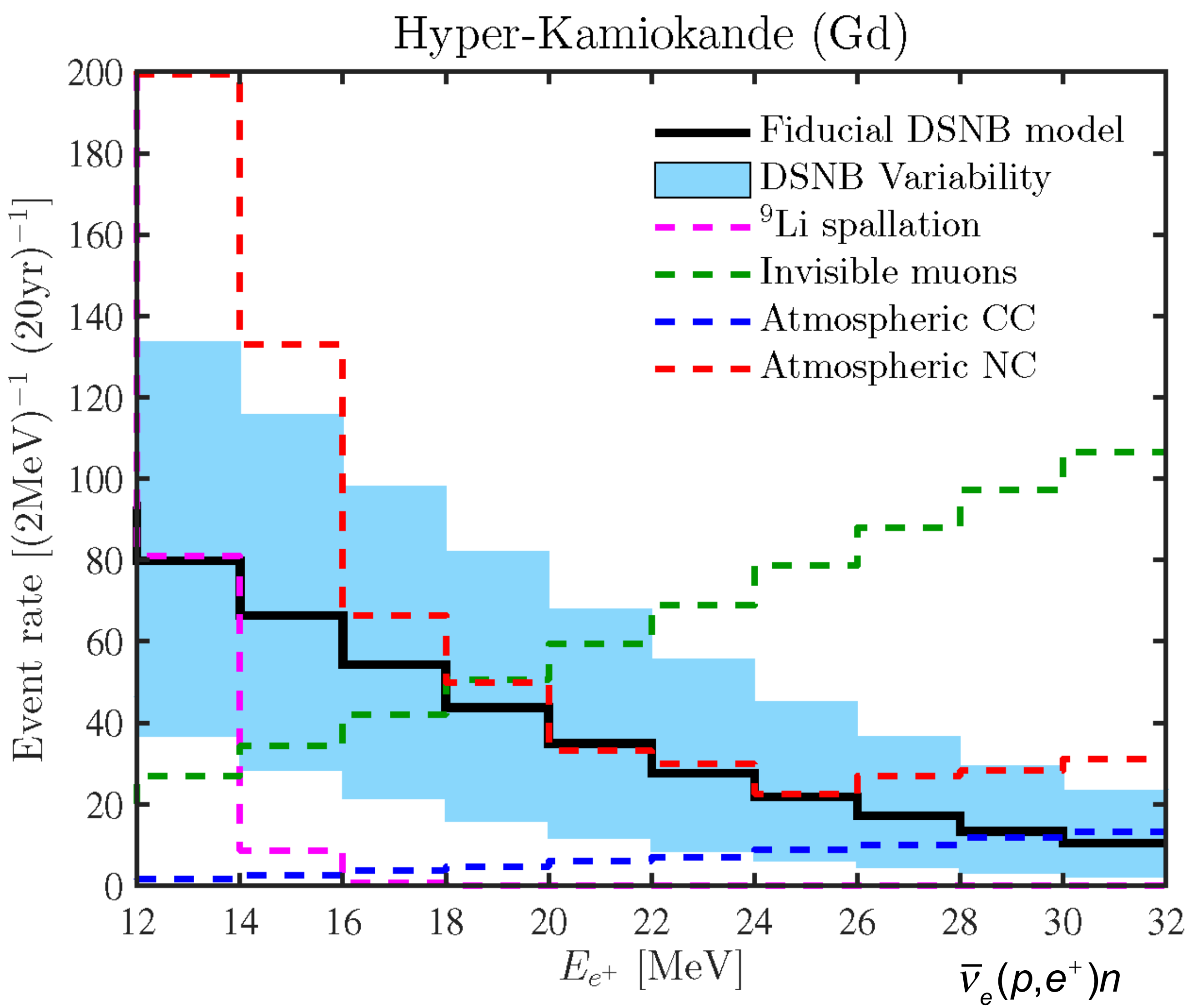
DSNB Detection Perspectives

The DSNB has not been observed yet. Most stringent limit is from Super-Kamiokande (SK):

$$\phi_{\bar{\nu}_e} \leq 2.8 - 3.0 \ \mathrm{cm}^{-2}\mathrm{s}^{-1}$$ for E > 17.3 MeV



From talk Oct 26, 2017 by Irene Tamborro.
See Minzzi, Tamborra et al (2016)



Hyper-K (187 kilotons of water Gadolnium doped) – successor to Super-K (22.5 kilotons) in Japan. "Under development". Gd helps reduce the background from muons.

Closer by …

NEUTRINO BURST
OBSERVED FEBRUARY 23, 1987

- Originated from SN 1987A in the Large Magellanic Cloud, 55 kpc distant. First signal from supernova (supernova detected optically, neutrino data then searched).

- Detected in 3 locations - IMB - Cleveland; Kamiokande - Japan; and Baksan - USSR.

- Observed at Kamiokande and IMB* - 19 neutrino events, energies 8 to 40 MeV. Inferred neutrino temperature - 5 MeV. Total neutrino energy inferred at LMC - 2 to 5 x 10^{53} erg. Duration about 10 s with most emission during first 3 s.

- Neutrino flux at Earth about 5×10^{10} cm^{-2} s^{-1}.

- Observed coming through the Earth!

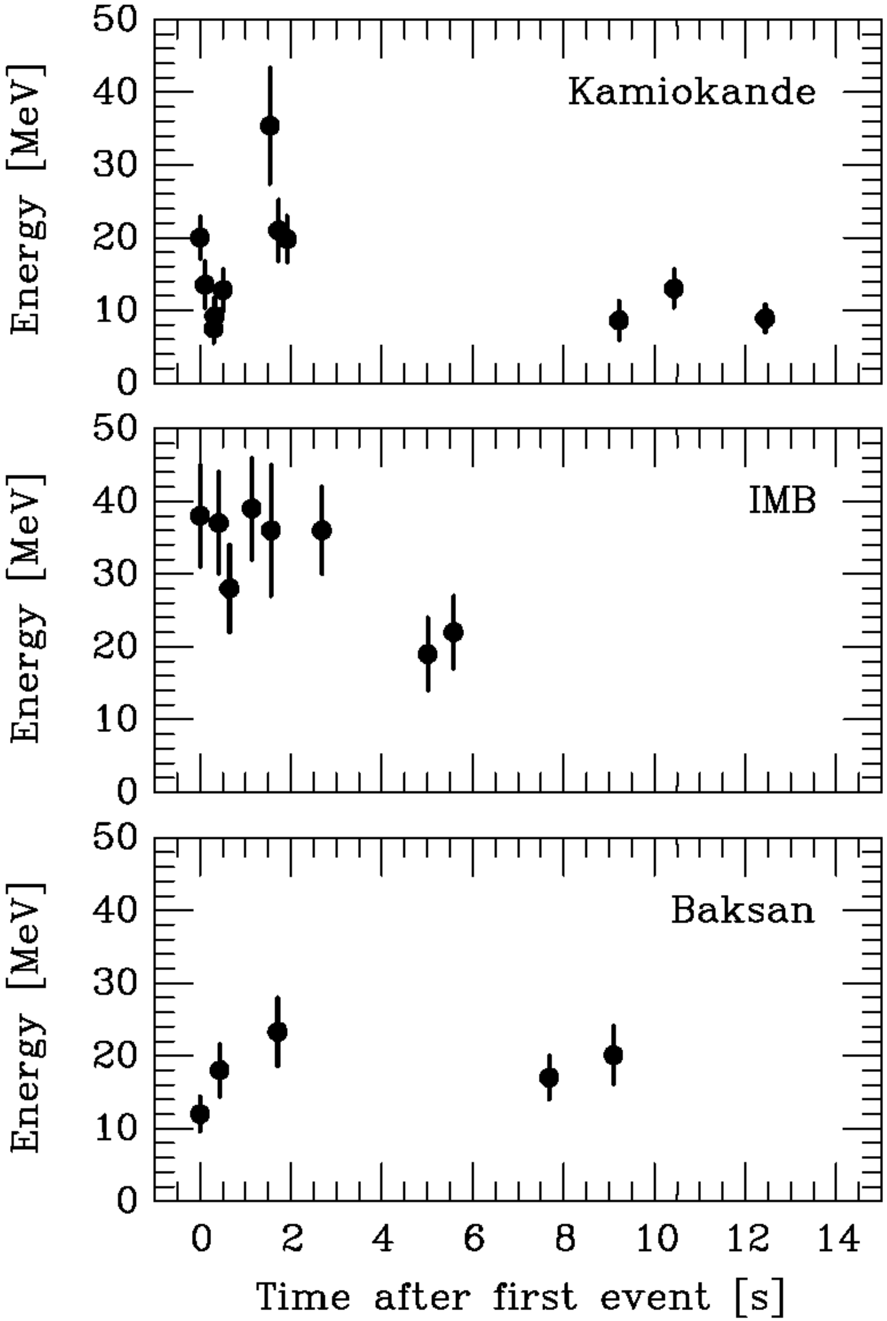
- Arrival at same time as light puts limits on neutrino mass (very small)

K II 2140 tons H_2O
IMB 6400 tons "

Cerenkov radiation from

$\bar{\nu}$ (p,n)e^+ - dominates
$\nu(e^-,e^-)\nu$ - relativistic e
all flavors ν

less than solar neutrino flux but neutrinos more energetic individually.



Japan

Ohio

Russia

Neutrino Burst Properties:

$$E_{tot} \sim \frac{3}{5}\frac{GM^2}{R}$$

$$\boxed{\sim 3\times10^{53} \ \text{erg}}$$

$M = 1.5\ M_\odot$

$R = 10$ km

emitted roughly equally in $\nu_e, \bar{\nu}_e, \nu_\mu, \bar{\nu}_\mu, \nu_\tau$, and $\bar{\nu}_\tau$

Time scale

$$\tau_{Diff} \sim \left(\frac{R^2}{l\ c}\right) \qquad l = \frac{1}{\kappa_\nu \rho}$$

$\kappa_\nu \sim 10^{-16}\ \text{cm}^2\ \text{gm}^{-1}$ for $\varepsilon_\nu =$ 50 MeV (next page)

$\rho \sim 3\times10^{14}\ \text{gm cm}^{-3} \quad \Rightarrow \quad l \sim$ 30 cm $\qquad R \sim 20$ km

$$\tau_{Diff} \sim \left(\frac{(2\times10^6)^2}{30\bullet3\times10^{10}}\right) \sim \boxed{5 \text{ sec}}$$

Very approximate

At densities above nuclear, the coherent scattering cross section (see last lecture) is no longer appropriate. One instead has scattering and absorption on individual neutrons and protons.

$$\text{Scattering:} \quad \kappa_{\nu s} \approx 1.0\times 10^{-20} \left(\frac{E_\nu}{\text{MeV}}\right)^2 \text{cm}^2\,\text{gm}^{-1}$$

$$\text{Absorption:}\ \kappa_{\nu \text{a}} \approx 4\kappa_{\nu s}$$

The actual neutrino energy needs to be obtained from a simulation but is at least tens of MeV. Take 50 MeV for the example here. Then $\kappa_\nu \sim 10^{-16}\,\text{cm}^2\ \text{g}^{-1}$. Gives $l_{\text{mfp}} \sim$ 1 m and $\tau_{\text{diff}} \sim$ few seconds.

Temperature:

$$L_\nu \approx \frac{E_{tot}}{6\tau_{Diff}} \approx 10^{52} \text{ erg s}^{-1} \text{ per flavor}$$

$$\approx \frac{7}{16}\left(4\pi\sigma R_\nu^2 T_\nu^4\right) \Rightarrow \boxed{T_\nu \approx 4.5 \text{ MeV}}$$

for $R_\upsilon \approx 20$ km and $\tau_\nu = 3$ sec

Actually $\bar{R}_\nu$ is a little bit smaller and τ_{Diff} is a little bit longer but 4.5 MeV is about right.

A victory for theory

Back to supernovae:

There were fundamental problems in the late 1960's and early 1970's that precluded a physically complete description:

- Lack of realistic progenitor models (addressed in the 80s)
- Primitive radiation transport or none
- Neglect of weak neutral currents – discovered 1974
- Uncertainty in the equation of state at super-nuclear densities (started to be addressed in the 80s)
- Inability to do realistic multi-dimensional models - the current frontier
- Missing fundamental physics (still discussed – flavor mixing?)

1.C

Nuclear Physics **A324** (1979) 487–533 © *North-Holland Publishing Co., Amsterdam*

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EQUATION OF STATE IN THE GRAVITATIONAL COLLAPSE OF STARS

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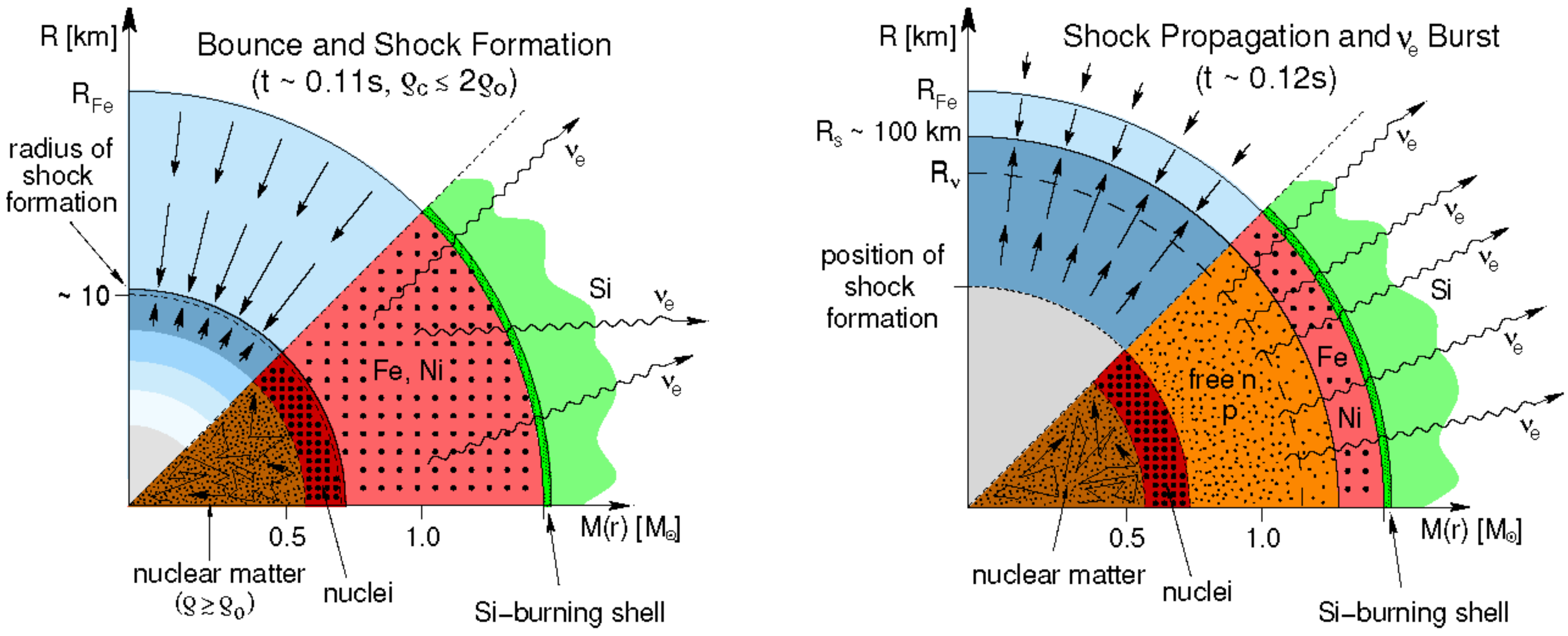
Received 12 February 1979

Abstract: The equation of state in stellar collapse is derived from simple considerations, the crucial ingredient being that the entropy per nucleon remains small, of the order of unity (in units of k), during the entire collapse. In the early regime, $\rho \sim 10^{10}–10^{13}$ g/cm^3, nuclei partially dissolve into α-particles and neutrons; the α-particles go back into the nuclei at higher densities. At the higher densities, nuclei are preserved right up to nuclear matter densities, at which point the nucleons are squeezed out of the nuclei. The low entropy per nucleon prevents the appearance of drip nucleons, which would add greatly to the net entropy.

We find that electrons are captured by nuclei, the capture on free protons being negligible in

BBAL 1979

- The explosion was low entropy
- Heat capacity of excited states kept temperature low
- Collapse continues to nuclear density and beyond
- Bounce on the nuclear repulsive force
- Possible strong hydrodynamic explosion - no longer believed



Radial distances R are indicated on the vertical axes, the corresponding enclosed masses M(r) are given on the horizontal axes. R_{Fe} , R_s, R_ν , R_g, and R_{ns} denote the iron-core radius, shock radius, neutrinospheric radius, gain radius (which separates neutrino cooling and heating layers), and proto-neutron star (PNS) radius, respectively. MCh defines the effective Chandrasekhar mass, M_{hc} the mass of the homologously collapsing inner core (where velocity u propto r), ρ_χ the central density, and r_O = 2.7 x 10^{14} g cm^{-3} the nuclear saturation density. (Figure taken from Janka et al, 2007)

THE ASTROPHYSICAL JOURNAL, **295**: 14–23, 1985 August 1
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REVIVAL OF A STALLED SUPERNOVA SHOCK BY NEUTRINO HEATING*

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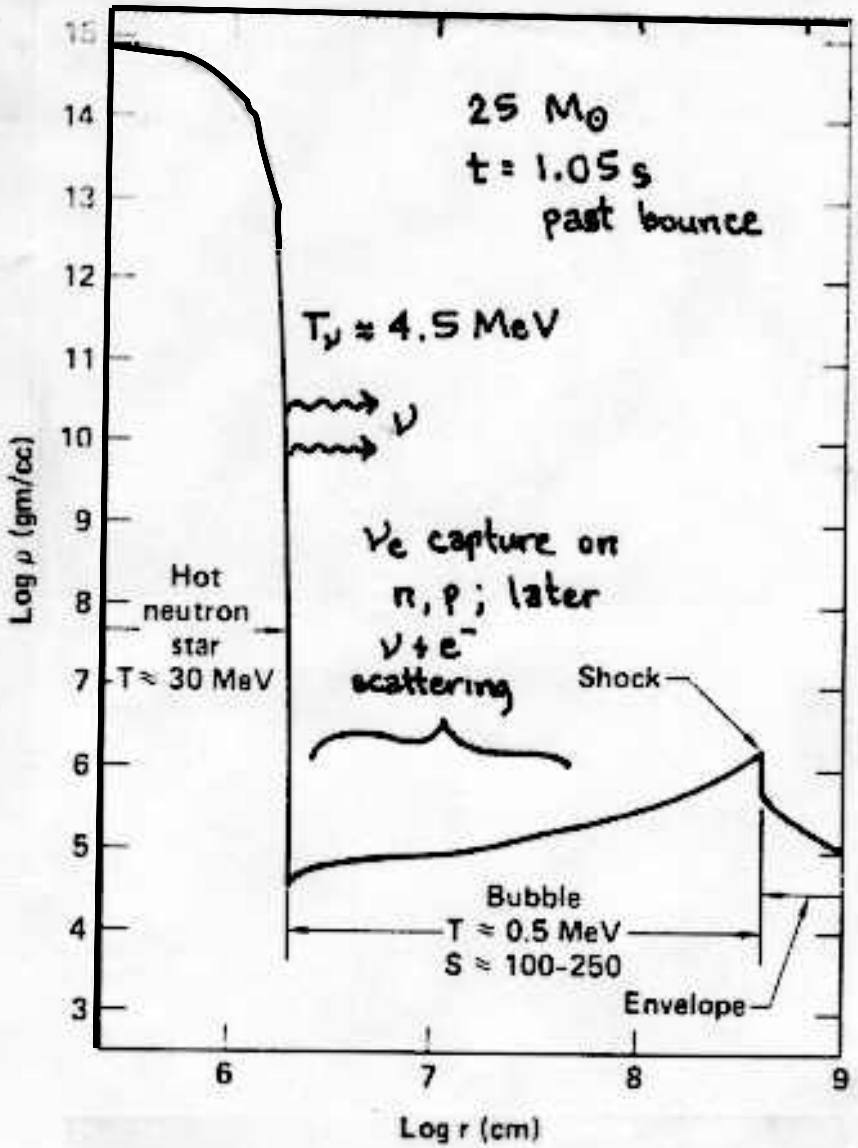
Received 1984 March 23; accepted 1985 February 5

ABSTRACT

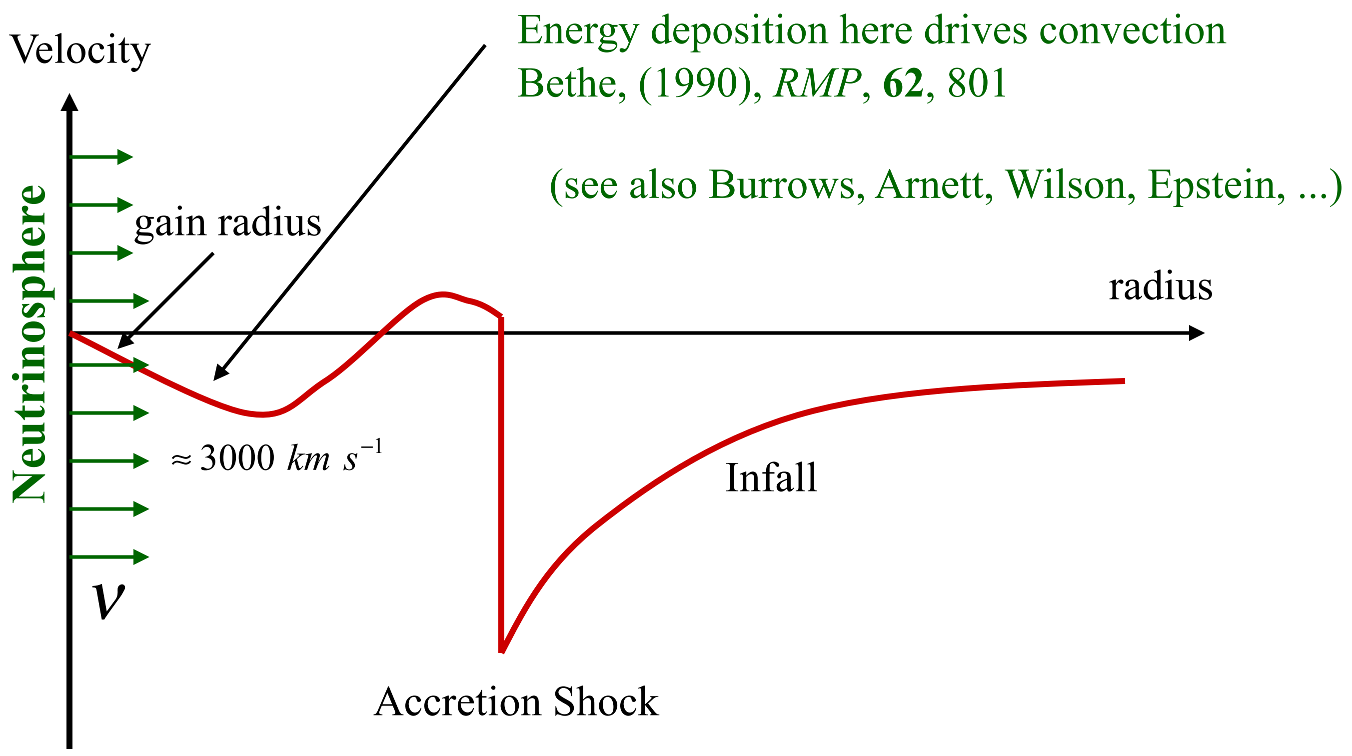
We analyze the mechanism for revival of a stalled supernova shock found by one of us (J. R. W.) in a computation. Neutrinos from the hot, inner core of the supernova are absorbed in the outer layers, and although only about 0.1% of their energy is so absorbed, this is enough to eject the outer part of the star and leave only enough mass to form a neutron star. The neutrino absorption is independent of the density of material. After the shock recedes to some extent, neutrino heating establishes a sufficient pressure gradient to push the material beyond about 150 km outward, while the material further in falls rapidly toward the core. This makes the density near 150 km decrease spectacularly, creating a quasi-vacuum in which the pressure is mainly carried by radiation. This is a perfect condition to make the internal energy of the matter sufficient to escape from the gravitational attraction of the star. The net energy of the outgoing shock is about 4×10^{50} ergs.

Subject headings: neutrinos — shock waves — stars: supernovae

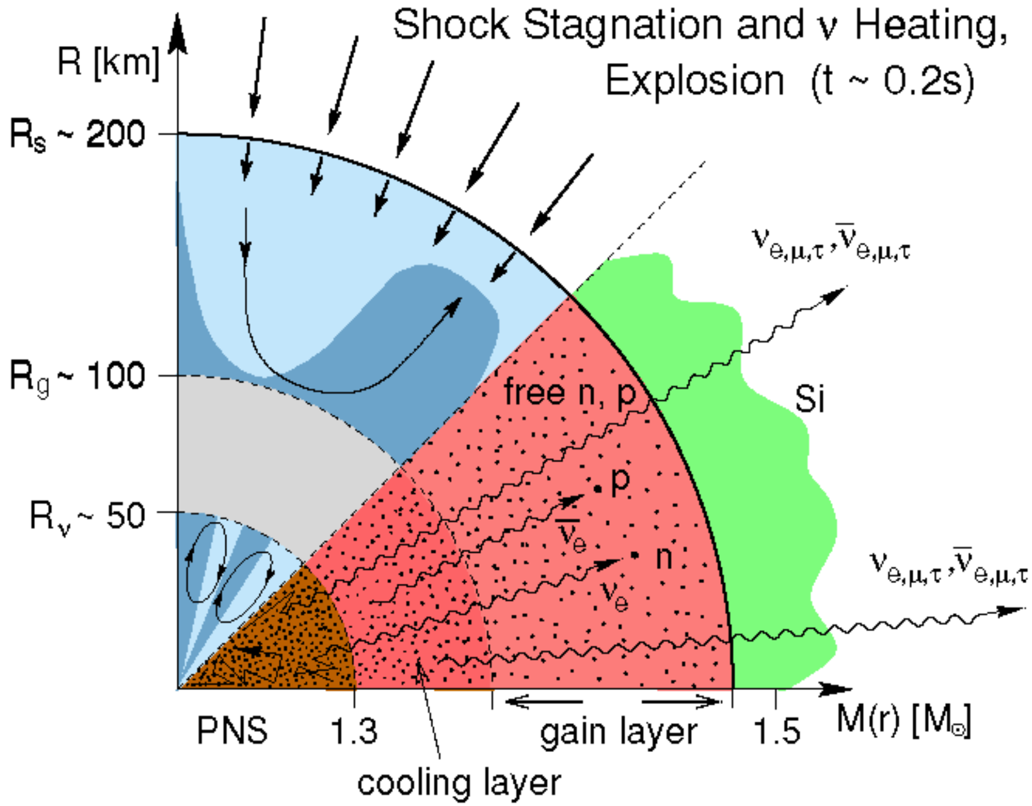
* See also conference proceedings by Wilson (1982)



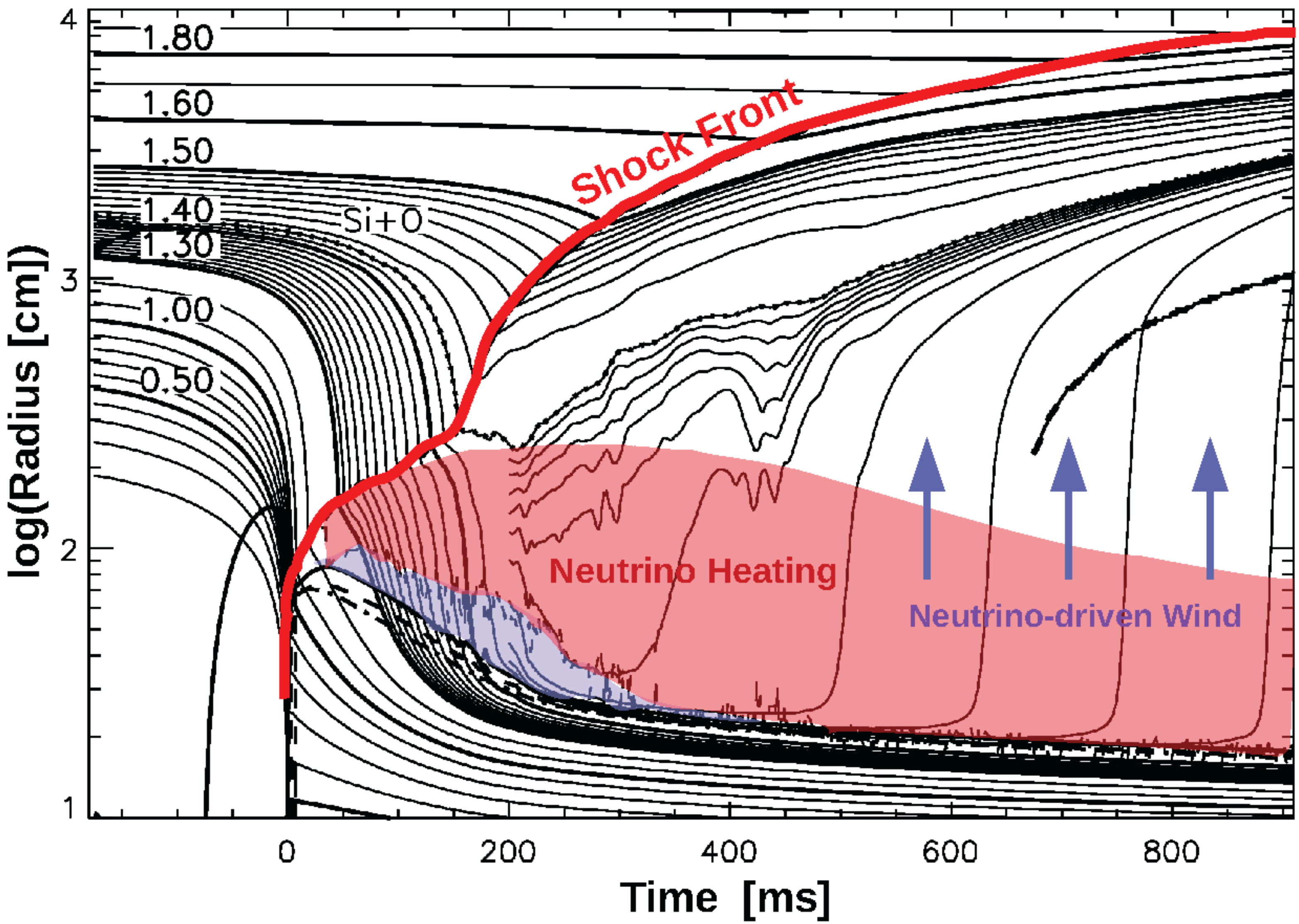
25 $M_\odot$
t = 1.05 s
past bounce
$T_\nu \approx 4.5$ MeV
ν
ν_e capture on
n, p; later
$\nu + e^-$
scattering
Hot
neutron
star
T ≈ 30 MeV
Shock
Bubble
T ≈ 0.5 MeV
S ≈ 100-250
Envelope
Log ρ (gm/cc)
Log r (cm)



Inside the shock, matter is in approximate hydrostatic equilibrium. Inside the gain radius there is net energy loss to neutrinos. Outside there is net energy gain from neutrino deposition. At any one time there is about 0.1 solar masses in the gain region absorbing a few percent of the neutrino luminosity.



Radial distances R are indicated on the vertical axes, the corresponding enclosed masses M(r) are given on the horizontal axes. R_{Fe} , R_s, R_ν , R_g, and R_{ns} denote the iron-core radius, shock radius, neutrinospheric radius, gain radius (which separates neutrino cooling and heating layers), and proto-neutron star (PNS) radius, respectively. MCh defines the effective Chandrasekhar mass, M_{hc} the mass of the homologously collapsing inner core (where velocity u propto r), ρ_χ the central density, and r_O = 2.7 x 10^{14} g cm^{-3} the nuclear saturation density. (Figure taken from Janka et al, 2007)


1.80
1.60
1.50
1.40
1.30
Si+O
1.00
0.50
Shock Front
Neutrino Heating
Neutrino-driven Wind
log(Radius [cm])
Time [ms]
4
3
2
1
0
200
400
600
800

The Neutrino Wind

An unavoidable consequence of neutron star formation is the "neutrino wind". As the ~ 3 x 10^{53} erg of neutrinos flow through the atmosphere of the cooling contracting protoneutron star mass loss is driven. The power for the wind is deposited chiefly by electron-neutrinos and antineutrinos interacting with neutrons and protons. We will discuss this more in the context of the r-process.

For now, note that it sets a lower bound to the kinetic energy of a supernova from a low mass star or accretion induced collapse of a white dwarf.

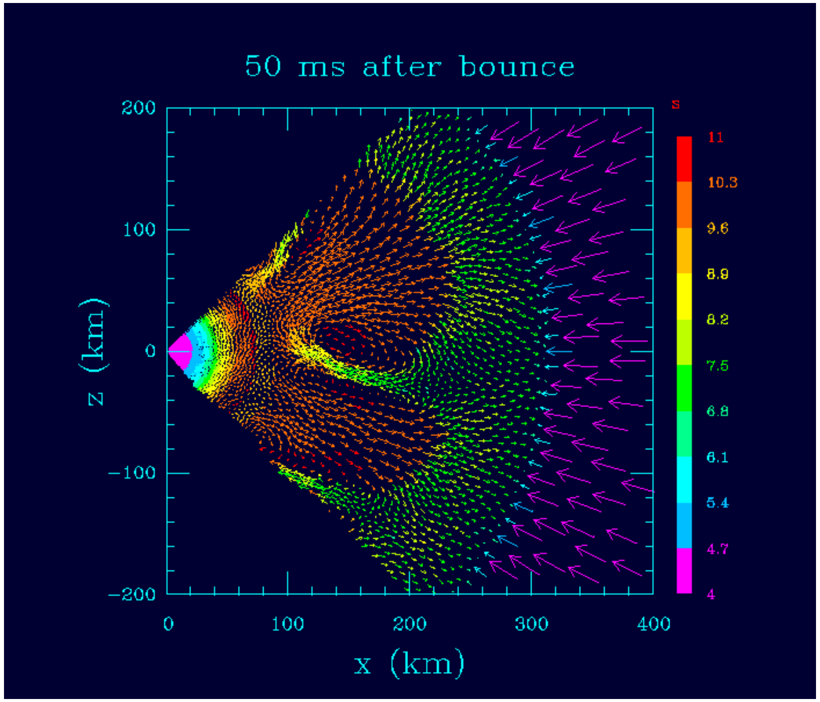
Power $= \dot{M}\, q$ where q is the energy of the wind per gram.

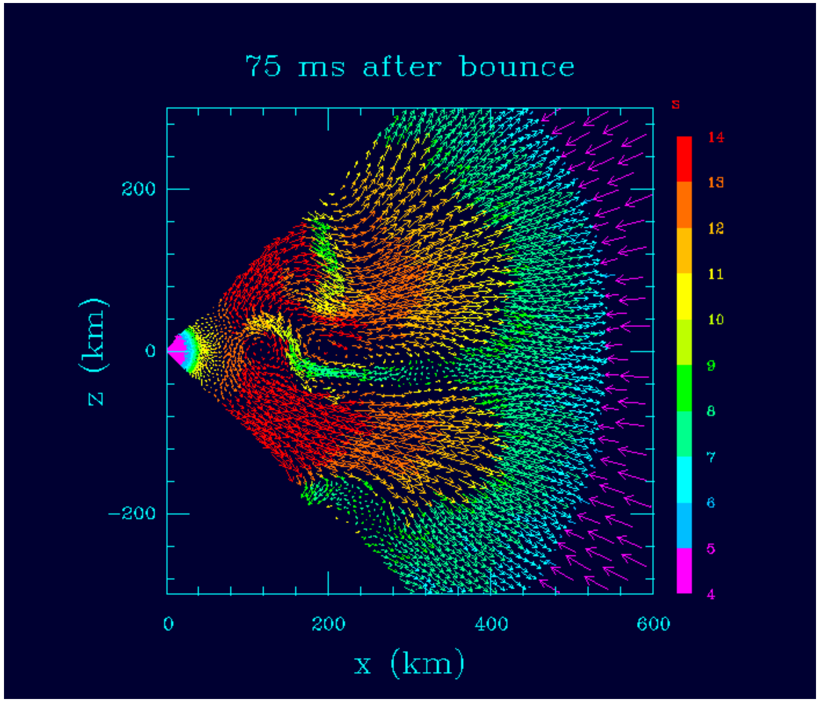
This turns out to be rather model independent, ~5 MeV.

Typical mass lost is 0.001 - 0.01 $M_\odot$, so an energy of

$\sim 10^{49} - 10^{50}$ erg is typical.

"Normal" Explosions





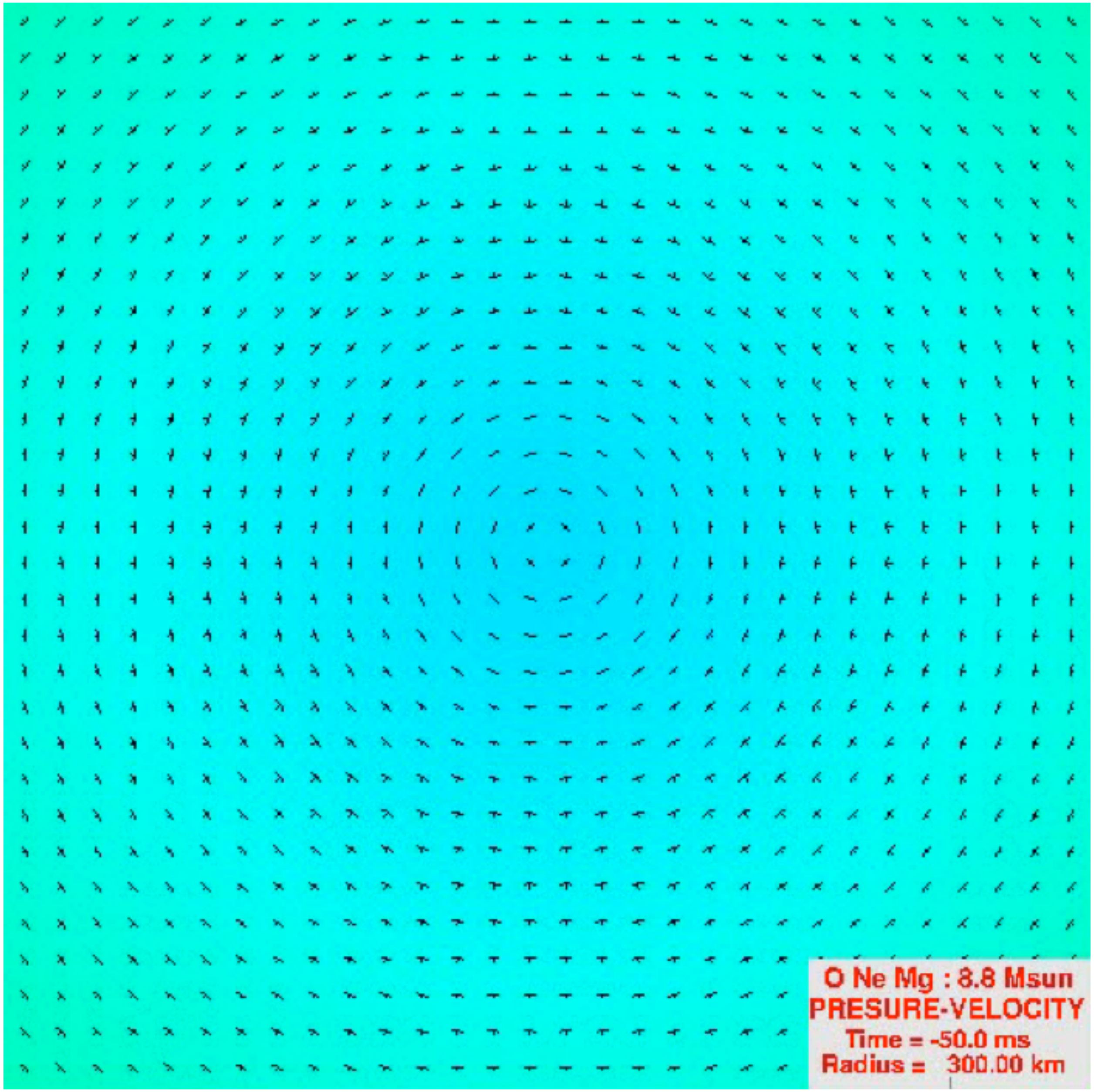
Herant and Woosley, 1995. 15 solar mass star.
successful explosion.
(see also Herant, Benz, & Colgate (1992), *ApJ*, **395**, 642)

8.8-Solar mass Progenitor of Nomoto: Neutrino-driven Wind Explosion

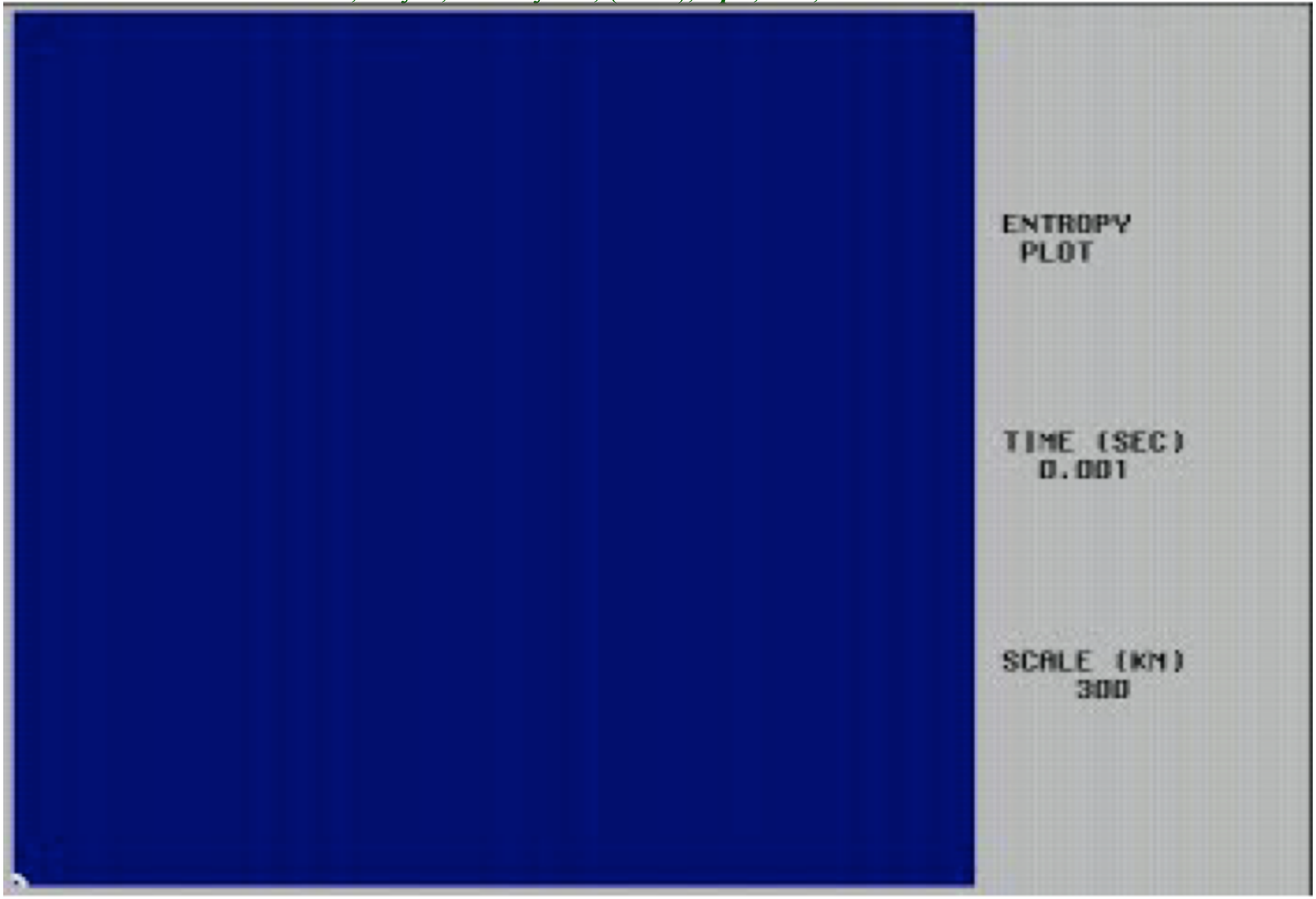
Burrows et al , 2007, AIPC, 937, 370

Explosion energy

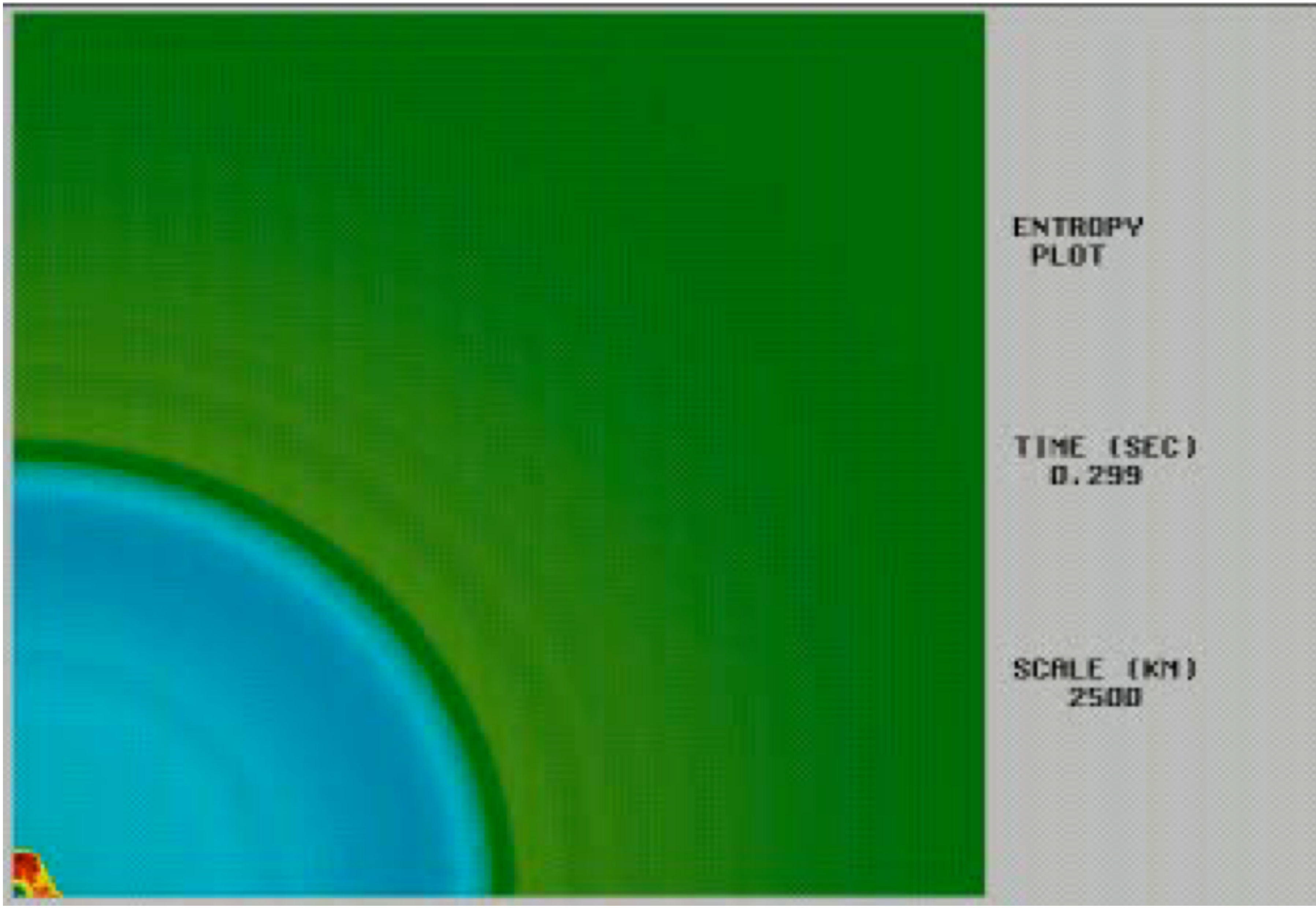
$\leq 10^{50}$ erg



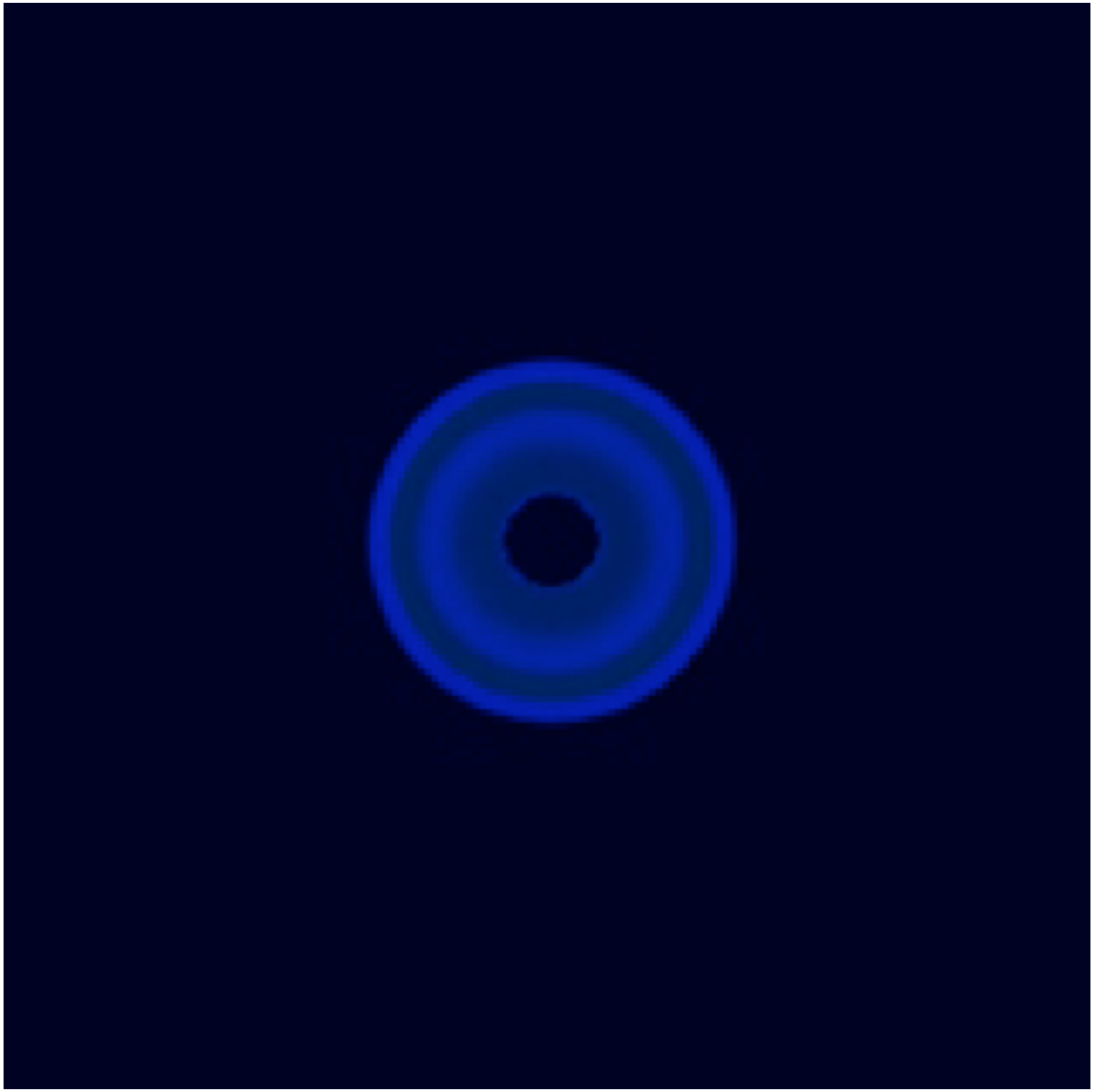
Burrows, Hayes, and Fryxell, (1995), ***ApJ***, **450**, 830



15 Solar masses – exploded with an energy of order 10^{51} erg.
see also Janka and Mueller, (1996), *A&A*, **306**, 167



At 408 ms, KE = 0.42 foe, stored dissociation energy is 0.38 foe, and the total explosion energy is still growing at 4.4 foe/s



Mezzacappa et al. (1998), *ApJ*, **495**, 911.

Using 15 solar mass progenitor WW95. Run for 500 ms.
1D flux limited multi-group neutrino transport coupled to 2D hydro.

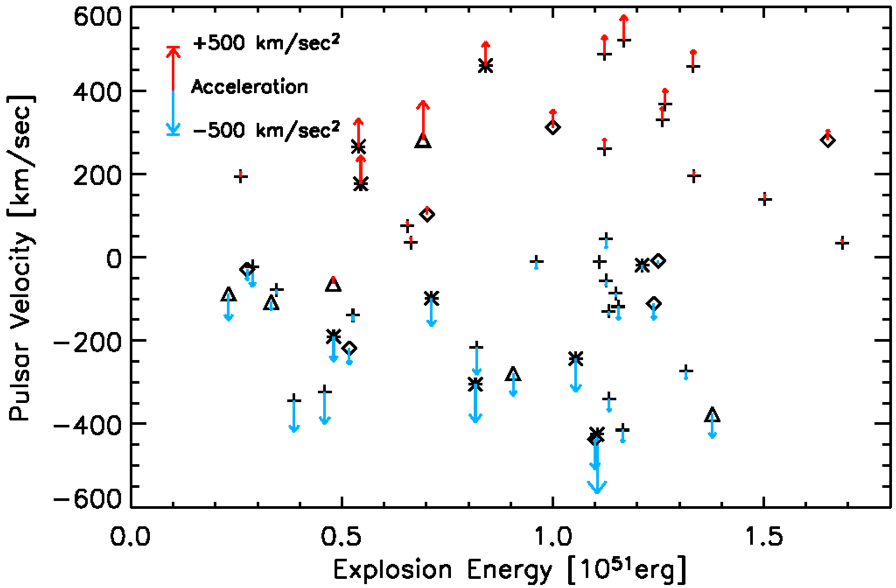
No explosion.

Beneficial Aspects of Convection

- Increased luminosity from beneath the neutrinosphere
- Turbulent motion is an extra source of pressure
- Transport of energy to regions far from the neutrinosphere (i.e., to where the shock is)

Also Helpful

- Decline in the accretion rate and accompanying ram pressure as time passes
- A shock that stalls at a large radius
- Accretion sustaining a high neutrino luminosity as time passes (able to continue at some angles in multi-D calculations even as the explosion develops).



Scheck et al. (2004)

Figure 1: Neutron star velocities and accelerations at one second after core bounce for a sample of simulations [4]. Different symbols denote different progenitor stars.

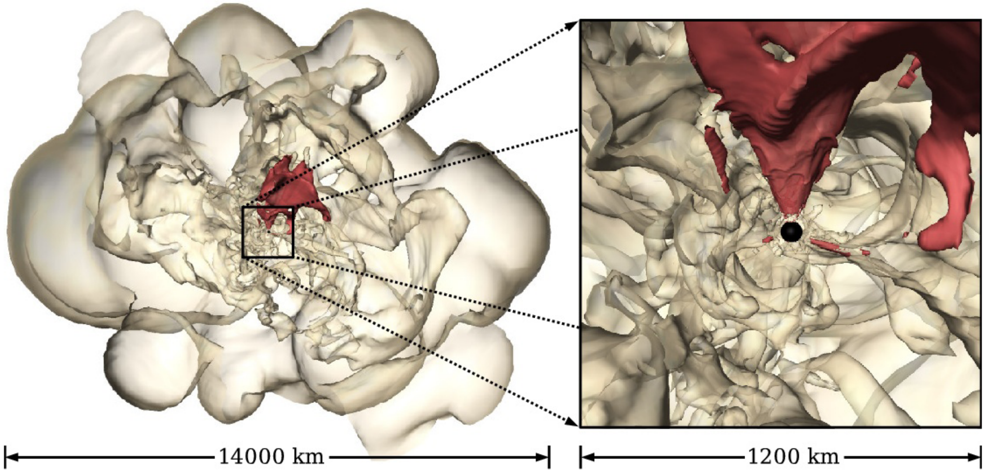
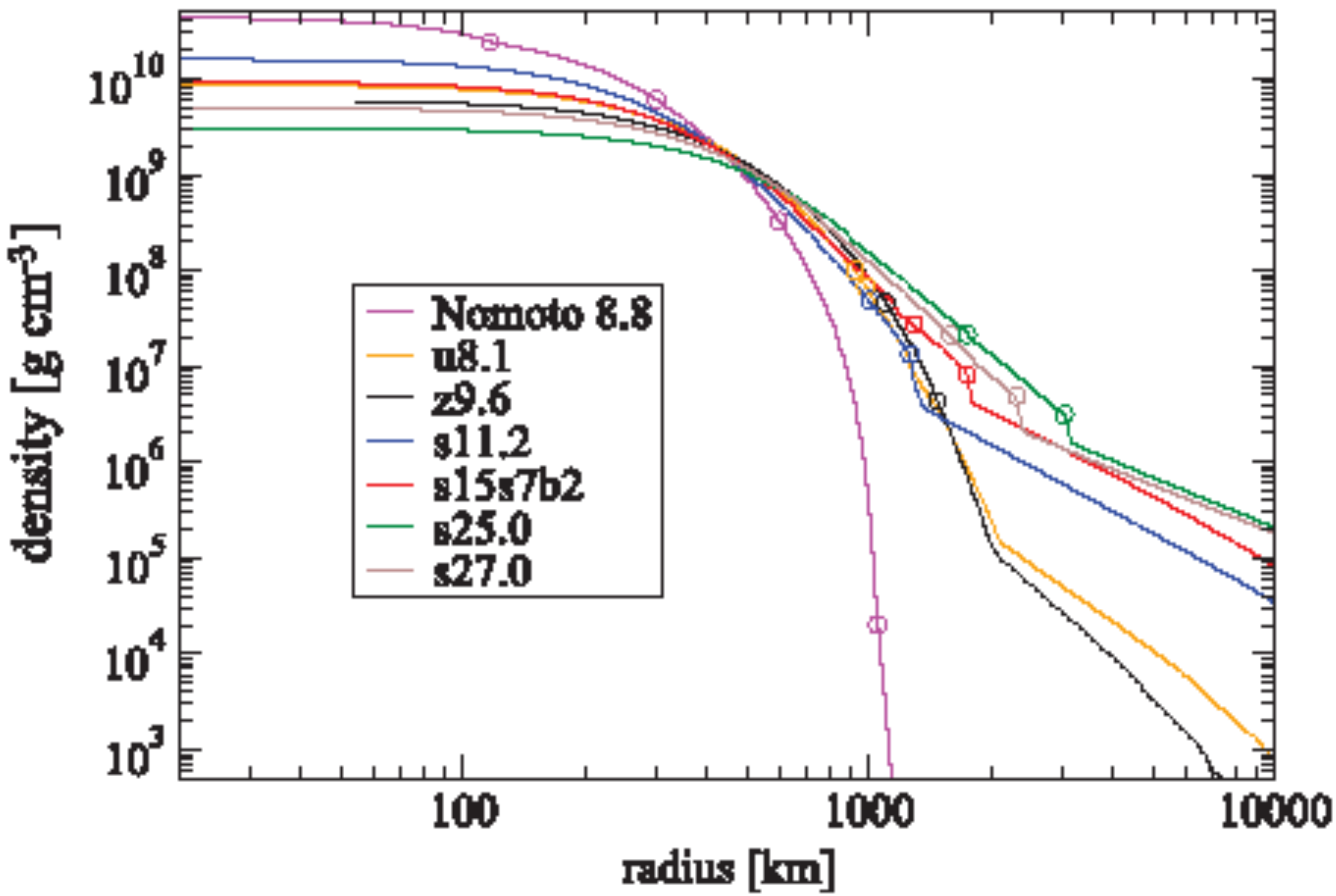
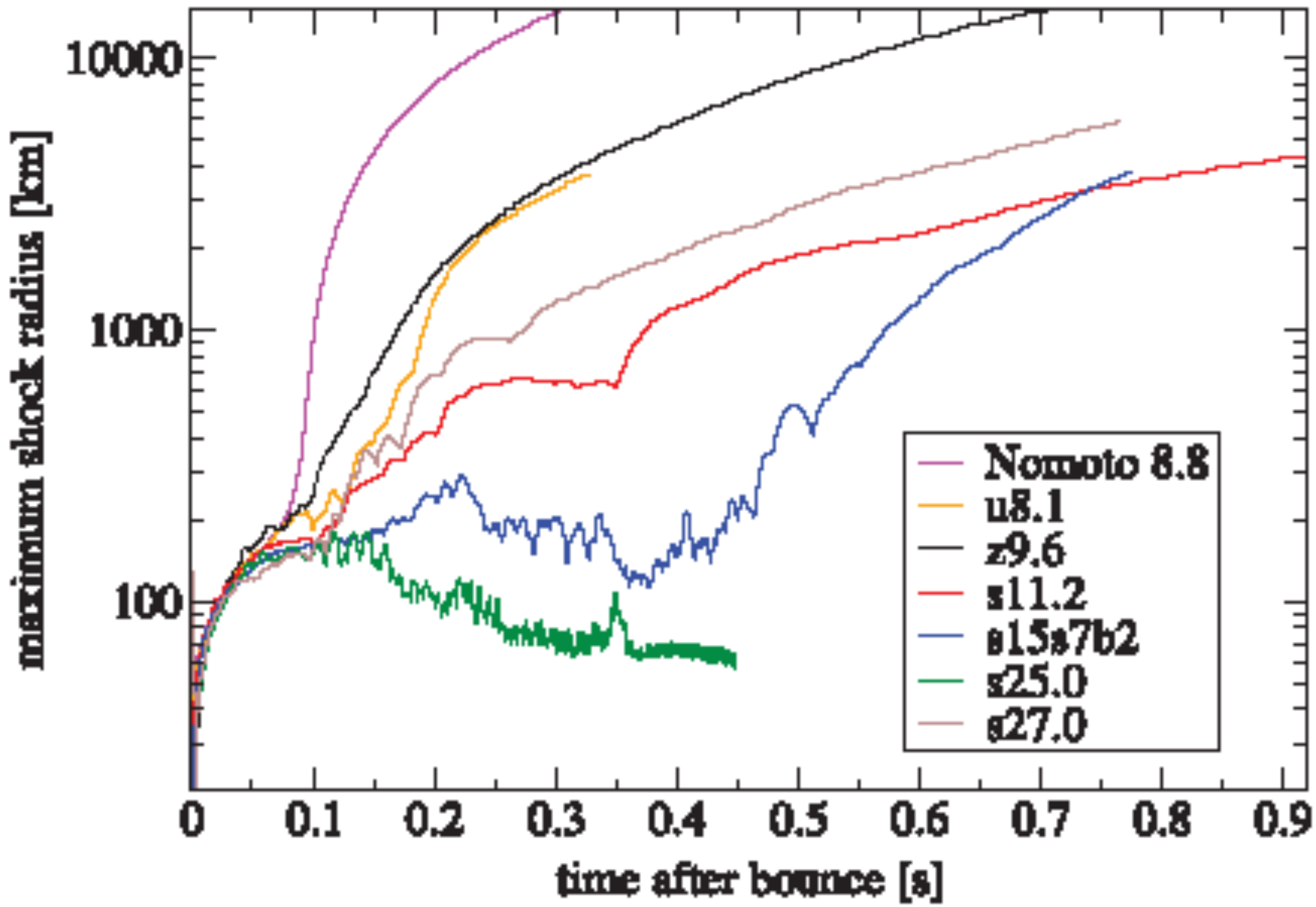


Figure 2: Three-dimensional simulation [7] one second after core bounce. The bright structure is a surface of constant proton-to-neutron ratio which roughly marks the outer boundaries of the neutrino-heated high-entropy bubbles. The dark surface, blown up in the right figure, is defined by a constant value for the mass flux per unit area and defines a downflow of matter towards the neutron star, the surface of which is indicated by the black sphere (corresponding to a density of 10^{11} g/cm^3).



density [g cm^{-3}]
radius [km]
10^{10}
10^9
10^8
10^7
10^6
10^5
10^4
10^3
100
1000
10000
Nomoto 8.8
u8.1
z9.6
s11.2
s15s7b2
s25.0
s27.0

Janka et al. 2012, *Prog. Theor. Exp. Phys.*, 01A309



Weak explosions for all 6 models in 2D except for 25 solar masses

Outcome sensitive to resolution and initial perturbations – Couch and Ott (2015)

The Astrophysical Journal, 799:5 (12pp), 2015 January 20 Couch & Ott

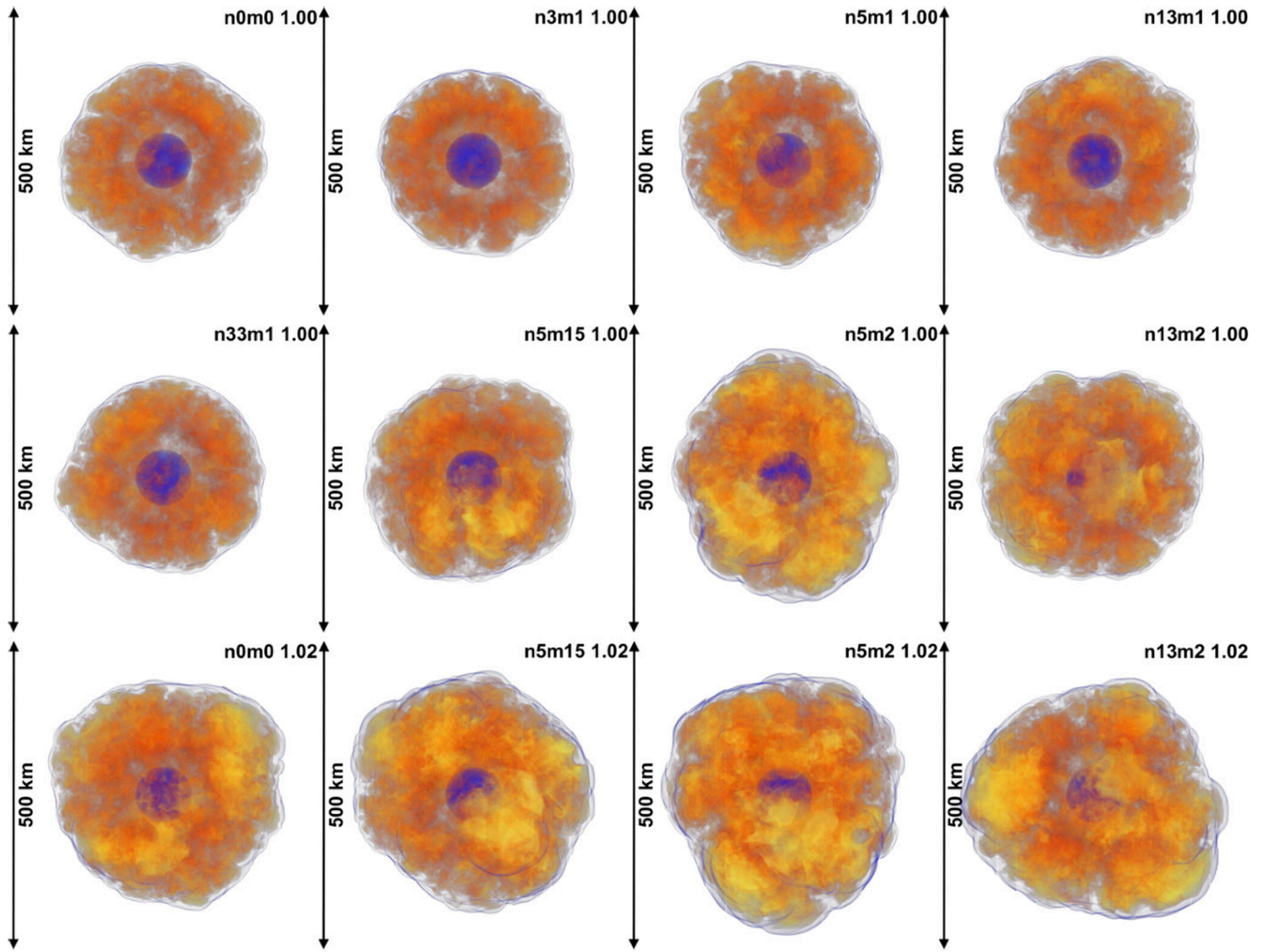


Figure 1. Volume renderings of specific entropy for several of the 3D simulations at 150 ms after bounce. Darker, red colors correspond to specific entropies of $\sim 14\,k_B$ baryon^{-1} while lighter, yellow colors correspond to entropies of $\sim 18\,k_B$ baryon^{-1}. The blue colors, which highlight the shock surface and the lower-entropy cooling region near the protoneutron star, correspond to specific entropies of $\sim 5\,k_B$ baryon^{-1}. Models with stronger perturbations show higher specific entropies in the gain layer and a greater shock extension. This is a result of the stronger turbulence and concomitant higher neutrino heating efficiency and turbulent pressure in these models.

Challenges

- Tough physics – nuclear EOS, neutrino opacities
- Tough problem computationally – must be 3D (convection is important). 6 flavors of neutrinos out of thermal equilibrium (thick to thin region crucial). Must be follwoed with multi-energy group and multi-angles
- Magnetic fields and rotation may be important
- If a black hole forms, problem must be done using relativistic (magnto-)hydrodynamics (general relativity, special relativity, magnetohydrodynamics)

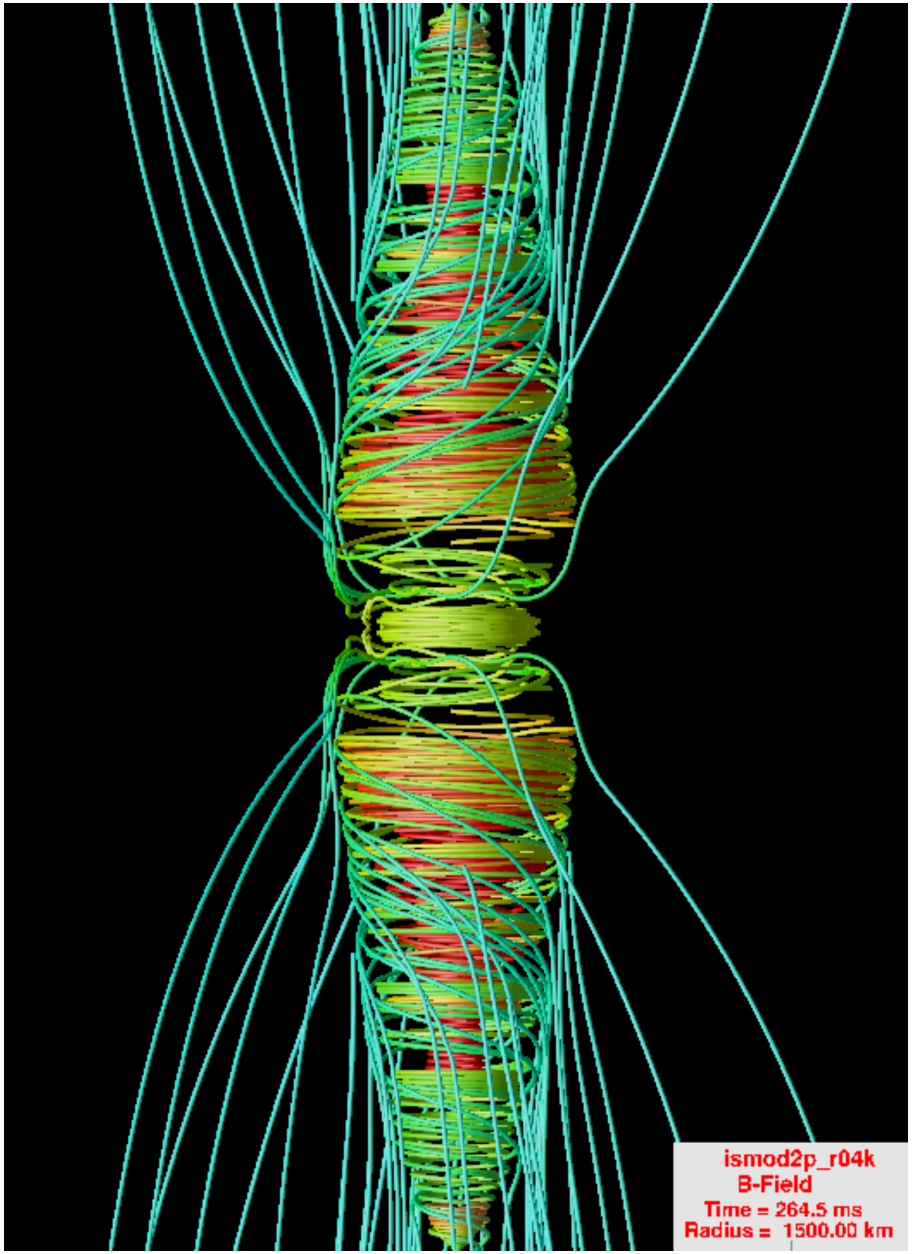
Rotationally Powered Models

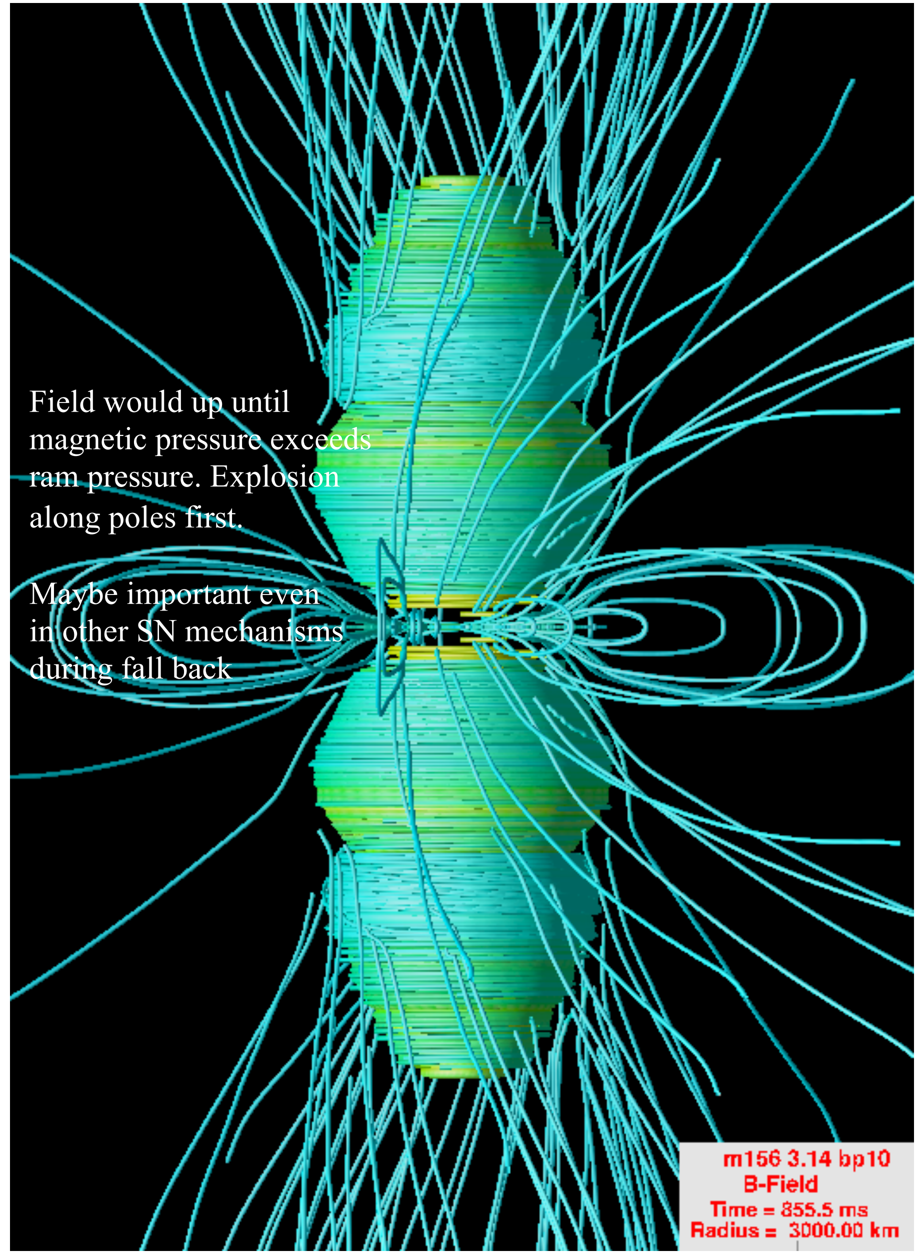
Common theme:

Need iron core rotation at death to correspond to a pulsar of < 4 ms period if rotation and B-fields are to matter. This is much faster than observed in common pulsars.

A concern:

If calculate the presupernova evolution with the same efficient magnetic field generating algorithms as used in some core collapse simulations, will it be rotating at all?

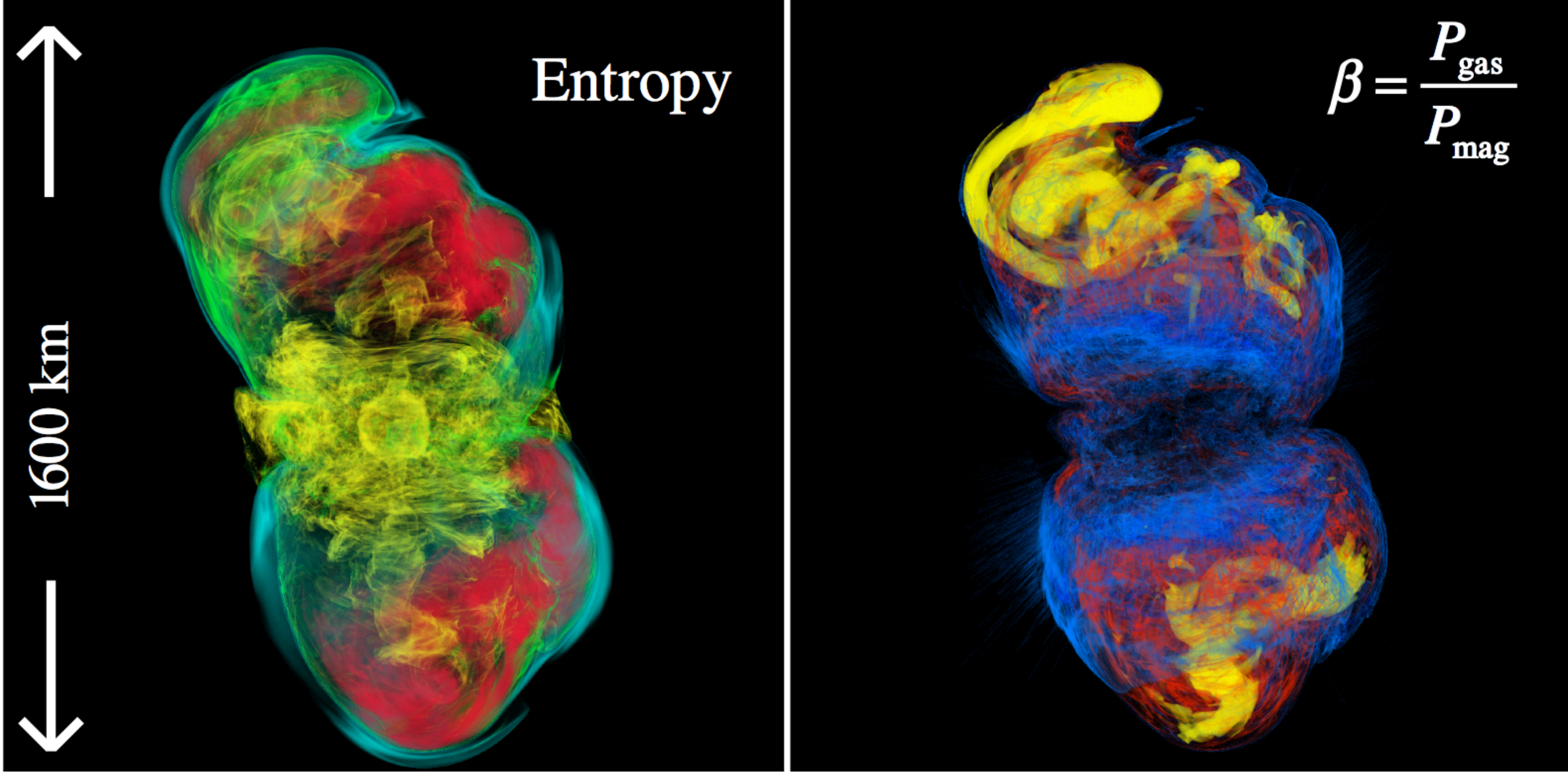




Field would up until magnetic pressure exceeds ram pressure. Explosion along poles first.

Maybe important even in other SN mechanisms during fall back

Burrows et al 2007, *ApJ*, **664**, 416



3D, GR-MHD
"Leakage scheme" for neutrinos
Mosta, Ott, et al (2014)
Does not produce explosion or jets during time followed

Assuming the emission of high amplitude ultra-relativistic MHD waves, one has a radiated power

$$P \sim 6 \text{ x } 10^{49} \; (1 \text{ ms}/P)^4 \, (B/10^{15} \text{ gauss})^2 \text{ erg s}^{-1}$$

and a total rotational kinetic energy

$$E_{rot} \sim 4 \; x \; 10^{52} \, (1 \text{ ms}/P)^2 \; (10 \text{ km}/R)^2 \text{ erg}$$

For magnetic fields to matter one thus needs magnetar-like magnetic fields and rotation periods (for the cold neutron star) of < 4 ms. This is inconsistent with what is seen in common pulsars. Where did the energy go?

Aside: Note an interesting trend. Bigger stars are harder to explode using neutrinos because they are more tightly bound and have big iron cores.

But they also rotate faster when they die.

Table 4: Pulsar Rotation Rate With Variable Remnant Mass[a]

Mass	Baryon[b] ($M_\odot$)	Gravitational[c] ($M_\odot$)	$J(M_{\rm bary})$ (10^{47} erg s)	BE (10^{53} erg)	Period[d] (ms)
$12\,M_\odot$	1.38	1.26	5.2	2.3	15
$15\,M_\odot$	1.47	1.33	7.5	2.5	11
$20\,M_\odot$	1.71	1.52	14	3.4	7.0
$25\,M_\odot$	1.88	1.66	17	4.1	6.3
$35\,M_\odot$ [e]	2.30	1.97	41	6.0	3.0

[a] Assuming a constant radius of 12 km and a moment of inertia $0.35MR^2$ (Lattimer & Prakash 2001)

[b] Mass before collapse where specific entropy is $4\,k_{\rm B}$/baryon

[c] Mass corrected for neutrino losses

[d] Not corrected for angular momentum carried away by neutrinos

[e] Became a Wolf-Rayet star during helium burning

Magnetic torques as described by Spruit, *A&A*, **381**, 923, (2002)

Table 5: Periods and Angular Momentum Estimates for Observed Young Pulsars

pulsar	current (ms)	initial (ms)	J_o (erg s)
PSR J0537-6910 (N157B, LMC)	16	∼10	8.8×10^{47}
PSR B0531+21 (crab)	33	21	4.2×10^{47}
PSR B0540-69 (LMC)	50	39	2.3×10^{47}
PSR B1509-58	150	20	4.4×10^{47}

Summary – Reasonable Expectations For Most Core-Collapse Supernovae

- Whether a given star will blow up by neutrino transport depends sensitively on the presupernova structure – on its mass. Even more so than the details of the collapse calculation

- The masses of stars that explode may not be a simply connected set

- Stars around 10 solar masses (+- 1 say) will be very easy to explode

- Typical supernovae (SN IIp) are the result of neutrino energy transport in stars with main sequence masses 8 to ~19 solar masses.

- Rotation may boost the explosion and mixing of supernovae coming from (rapidly rotating) stars above 20 solar masses, but many/most stars above ~20 solar masses become black holes.

- There is an island of “compact” pre-supernova stars at around 30 solar masses that might be exploded by unboosted neutrino transport

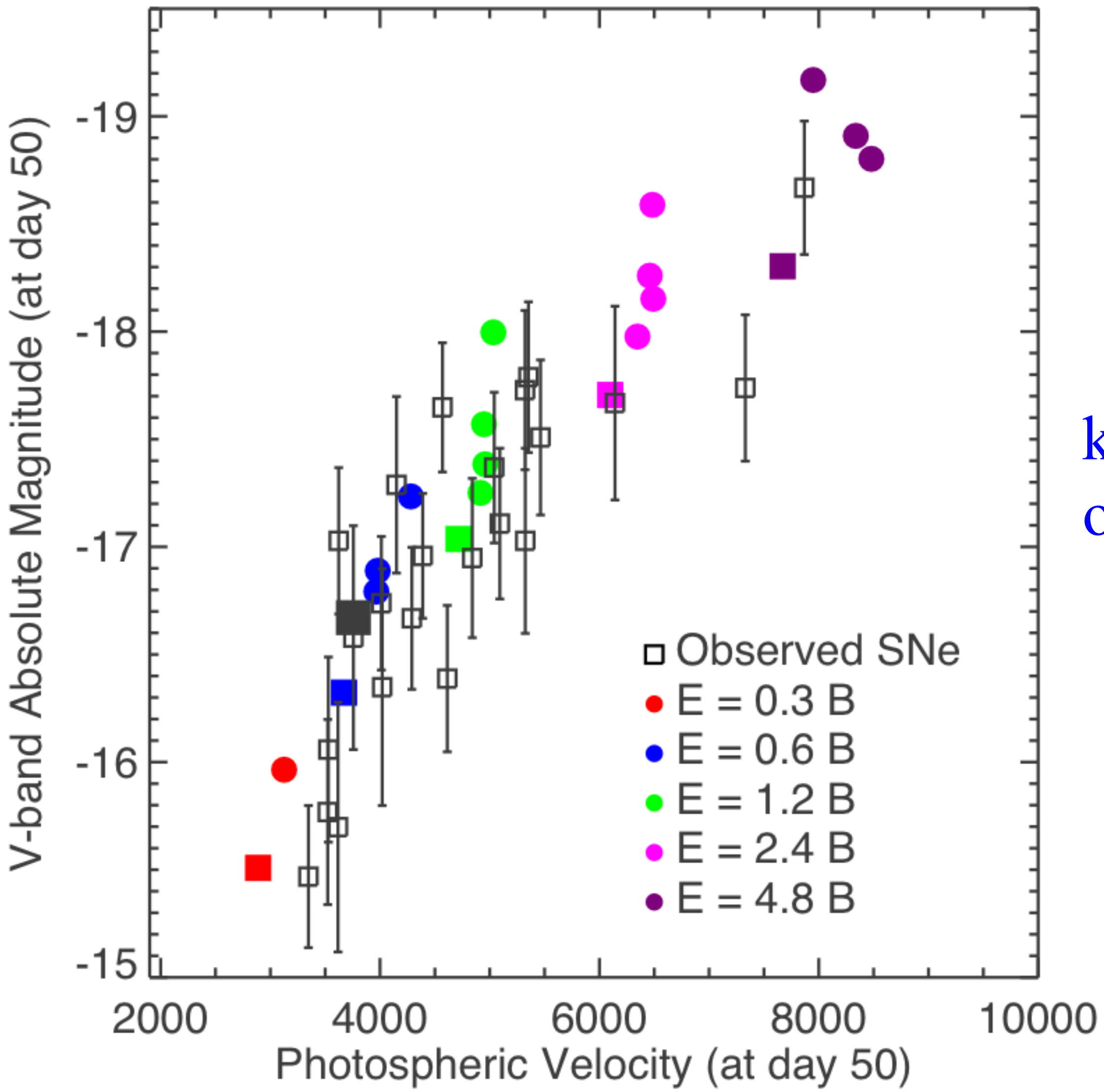
Continued

- Supernovae with explosion energies over 3 x 10^{51} probably do not come from unboosted neutrino transport.

$$\text{Explosion E} \sim \text{BE}_{n*} \times (\text{fraction in } \nu_e \bar{\nu}_e) \times (\frac{\tau_{\exp}}{\tau_{KH}}) \times (\text{Deposition efficiency})$$

$$\sim 3 \times 10^{53}\ \text{erg} \left(\frac{1}{3}\right)\left(\frac{1}{10}\right)(0.1) \sim 10^{51}\ \text{erg}$$

Kasen and Woosley (2009)



Observationally
The typical SN IIp has kinetic energy at infinity of 6 x 10^{50} erg, but with a wide spread.

Mixing During the Explosion

The Reverse Shock and Rayleigh-Taylor Instability:

The Sedov solution (adiabatic blast wave)

For $\rho = Ar^{-\omega}$

$$\boxed{v_{shock} = A^{\frac{1}{\omega-5}} \; E^{\frac{1}{5-\omega}} \; t^{(\omega-3)/(5-\omega)}}$$

$\omega < 5$

$\omega = 3 \quad \rightarrow \quad v_{shock} = \text{constant}$

$\omega < 3 \quad \rightarrow \quad v_{shock}$ slows down

$\omega > 3 \quad \rightarrow \quad v_{shock}$ speeds up

$$R_S = \left(\frac{Et^2}{\alpha A} \right)^{\frac{1}{\nu+2-\omega}}$$

$\nu =$ dimension of space

1, 2, or 3

$\alpha =$ const = f(ν)

Korobeinikov (1961)

If ρr^3 increases with radius, the shock will slow down.
The information that slowing is occuring will propagate inwards as a decelerating force directed towards the center. This force is in the opposite direction to the density gradient, since the density, even after the explosion, generally decreases for the material farther out.

$\Rightarrow$ Rayleigh-Taylor instability and mixing

Example:

For constant density and an adiabatic blast wave.
The constants of the problem are $E_{initial}$ and ρ. We seek a solution $r(t, E_{initial}, \rho)$. Assume that these are the only variables to which r is sensitive.

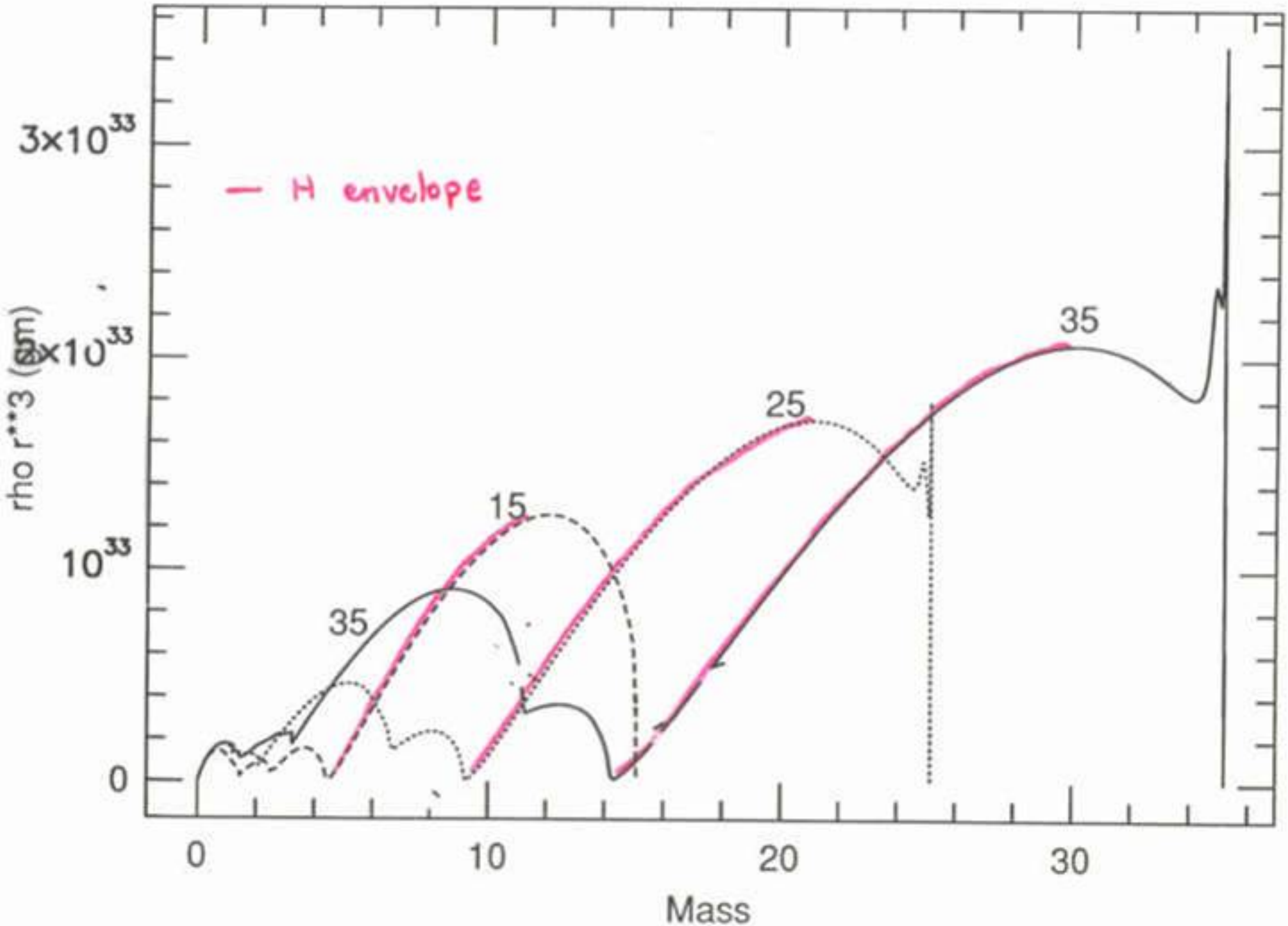
$$\text{Units} \quad E \quad \frac{\text{gm cm}^2}{\text{sec}^2}$$

$$\rho \quad \frac{\text{gm}}{\text{cm}^3}$$

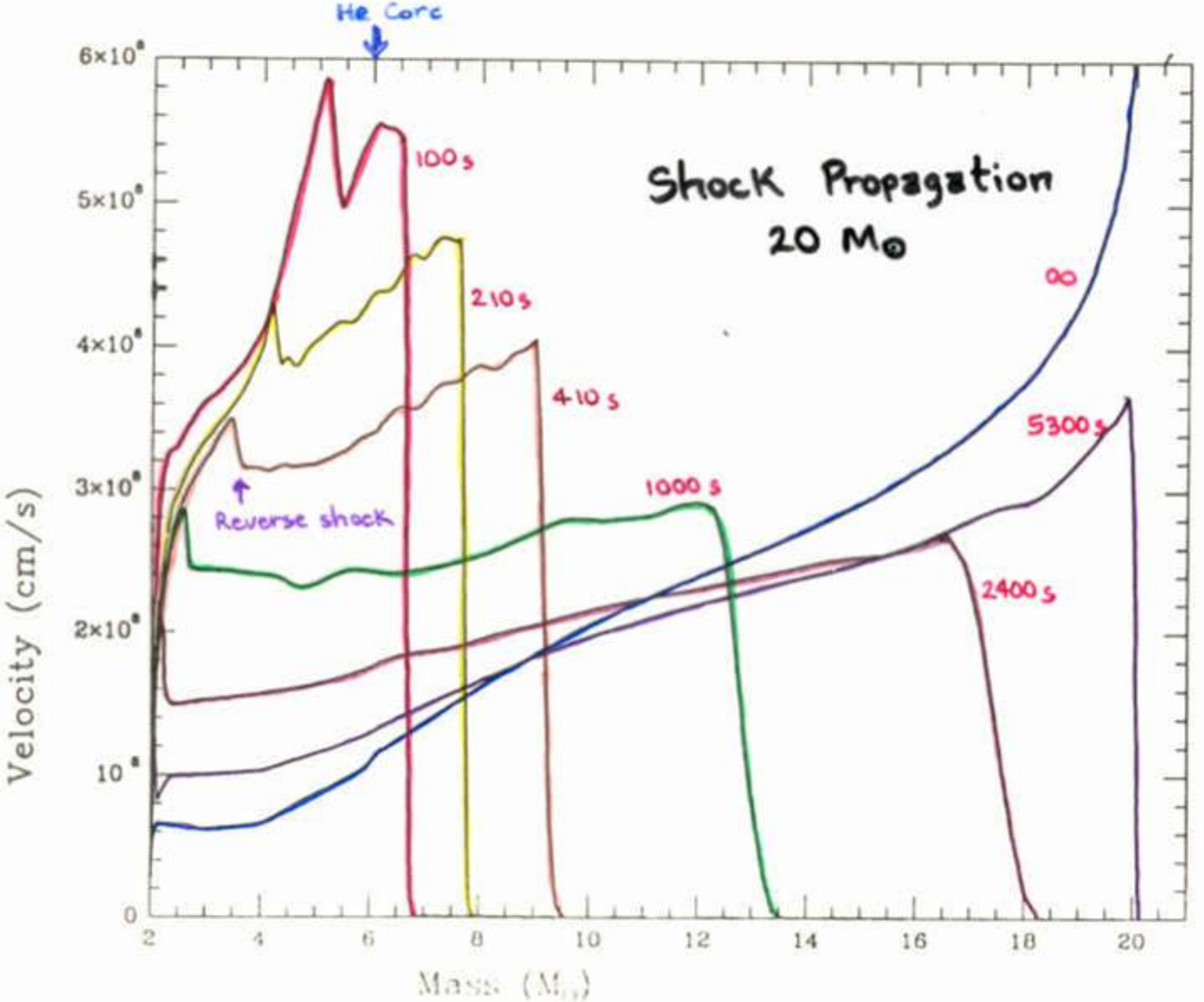
$$\frac{E}{\rho} \quad \frac{\text{cm}^5}{\text{sec}^2} \Rightarrow r^5 \propto \frac{E}{\rho} t^2$$

$$\boxed{r = K\left(\frac{E_{initial}}{\rho}\right)^{1/5} t^{2/5}}$$

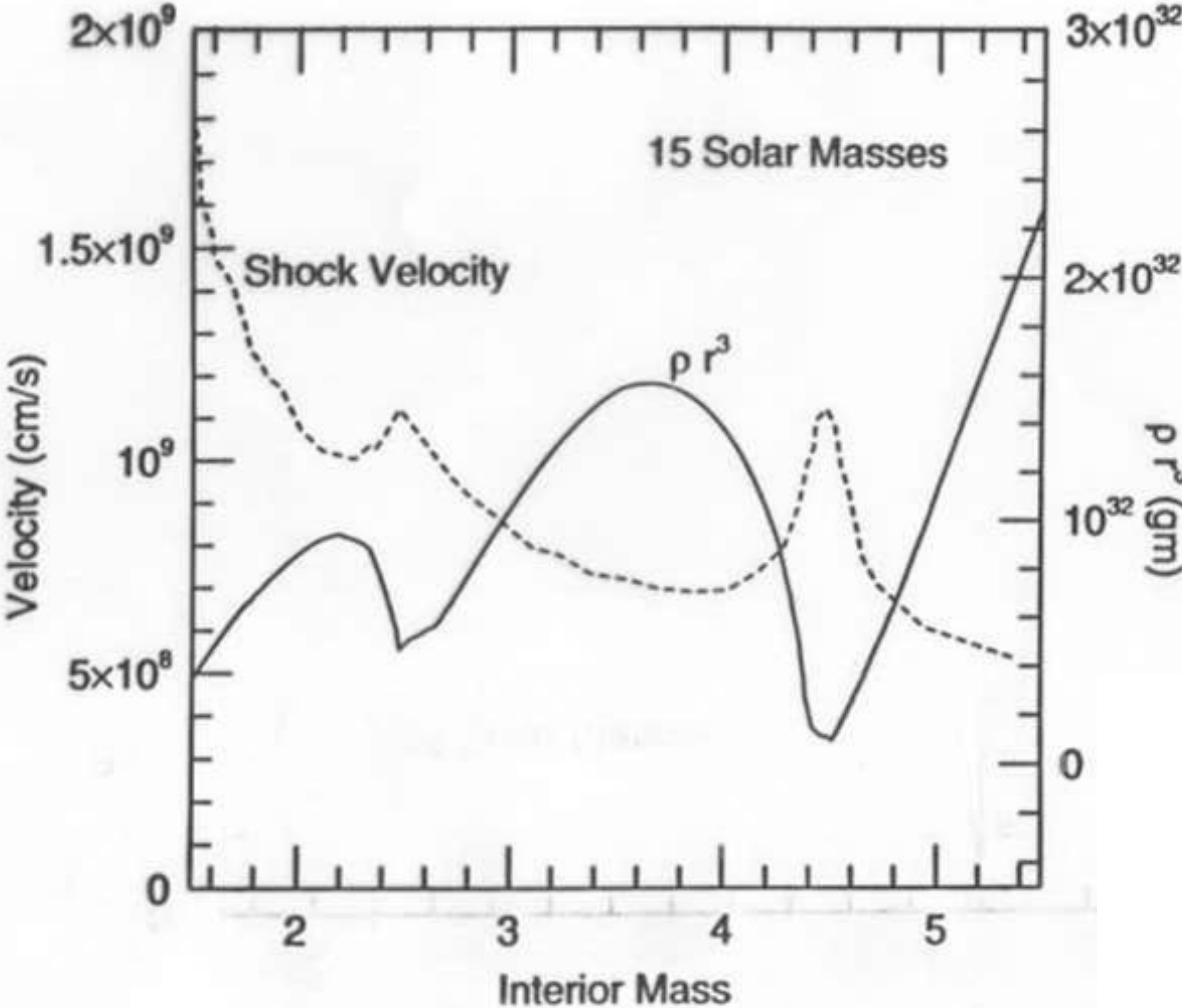
$$v = \frac{dr}{dt} = \frac{2}{5} K \left(\frac{E_{initial}}{\rho}\right)^{1/5} t^{-3/5}$$ which is our ω=0 case



3×10^33
10^33
0
— H envelope
rho r**3 (gm)
35
25
15
35
0
10
20
30
Mass

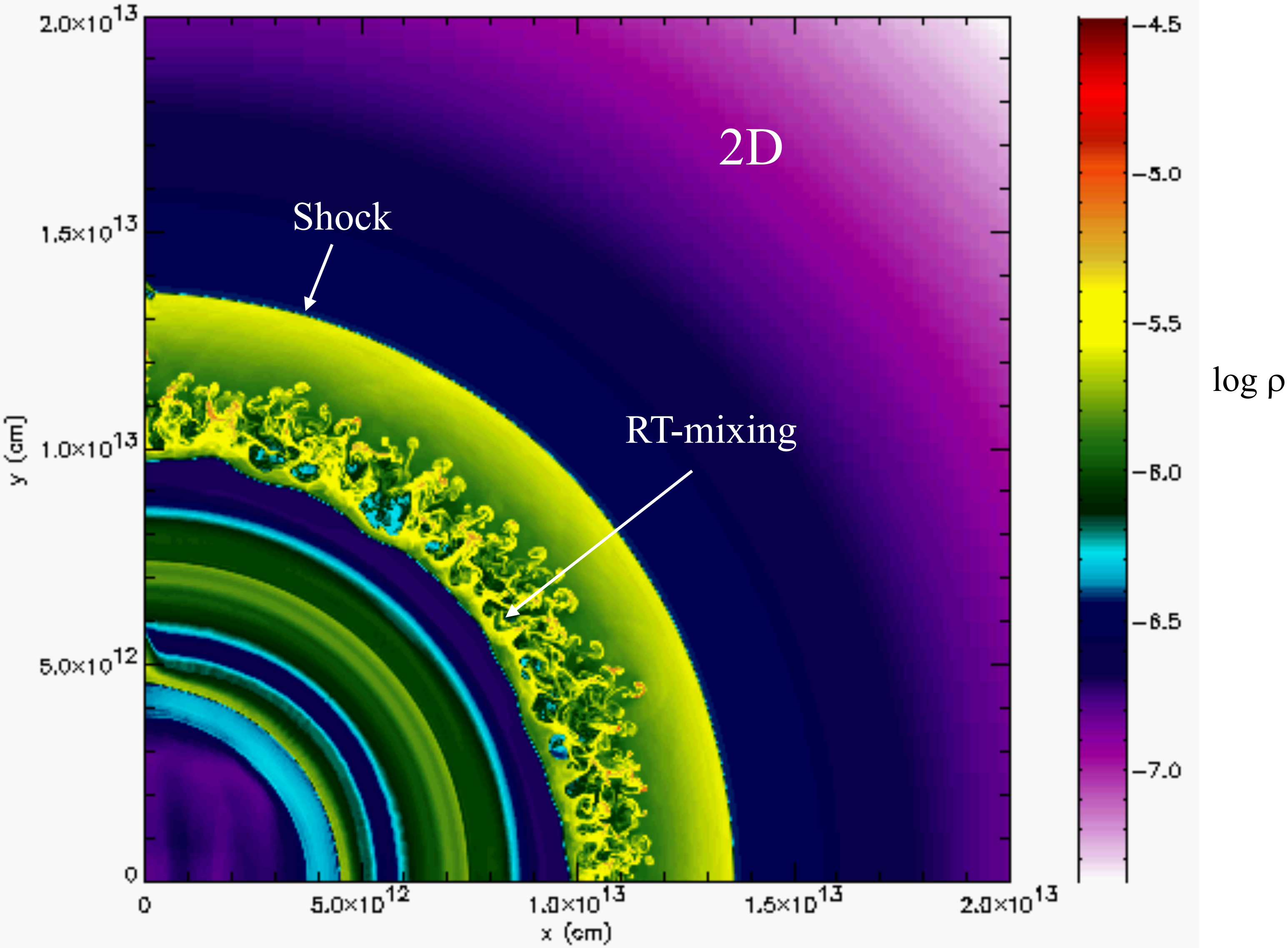


He Core
Shock Propagation
20 M⊙
100 s
210 s
410 s
1000 s
2400 s
5300 s
∞
Reverse shock
6×10⁸
5×10⁸
4×10⁸
3×10⁸
2×10⁸
10⁸
0
2
4
6
8
10
12
14
16
18
20
Velocity (cm/s)
Mass (M⊙)

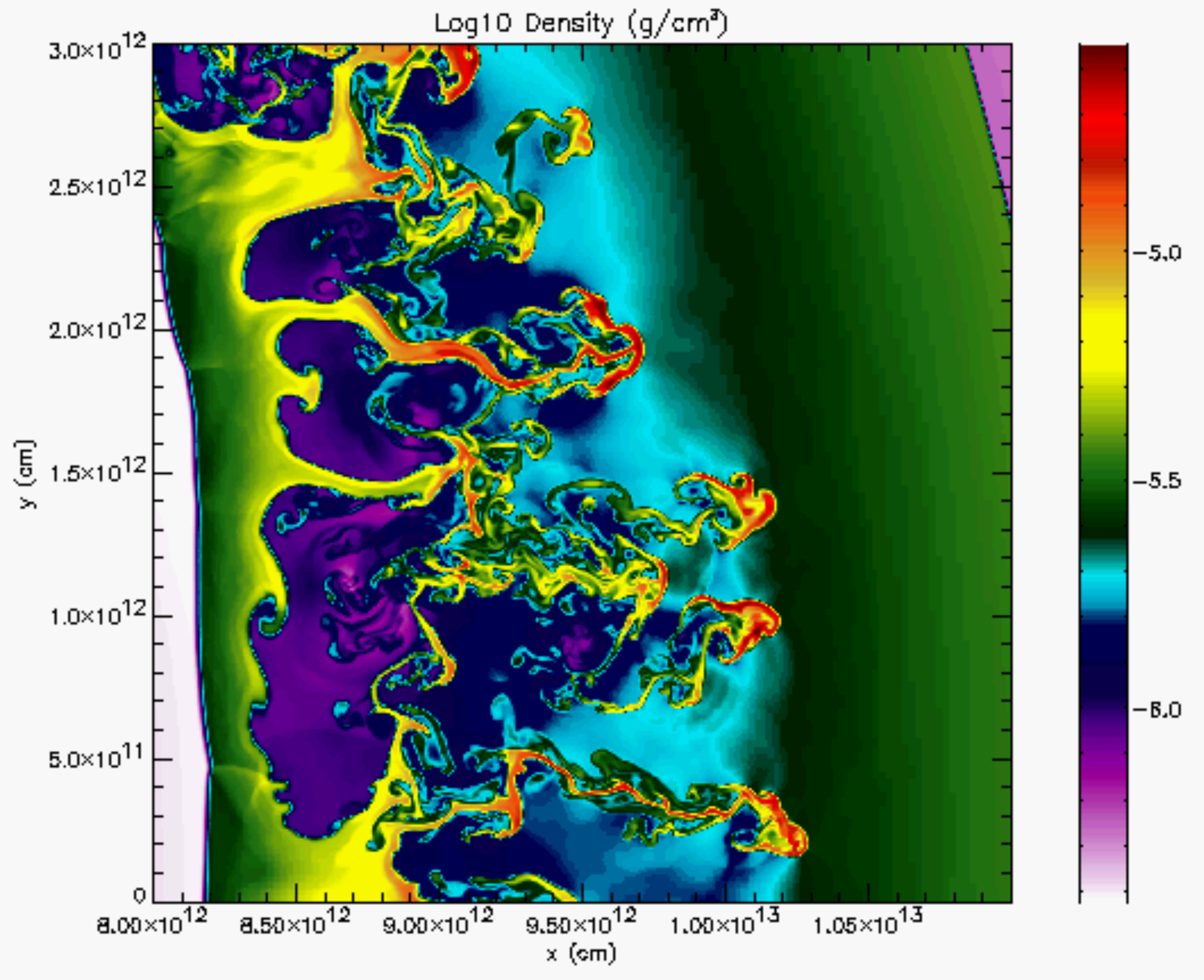


2×10⁹
1.5×10⁹
10⁹
5×10⁸
0
Velocity (cm/s)
3×10³²
2×10³²
10³²
0
ρ r³ (gm)
15 Solar Masses
Shock Velocity
ρ r³
2
3
4
5
Interior Mass

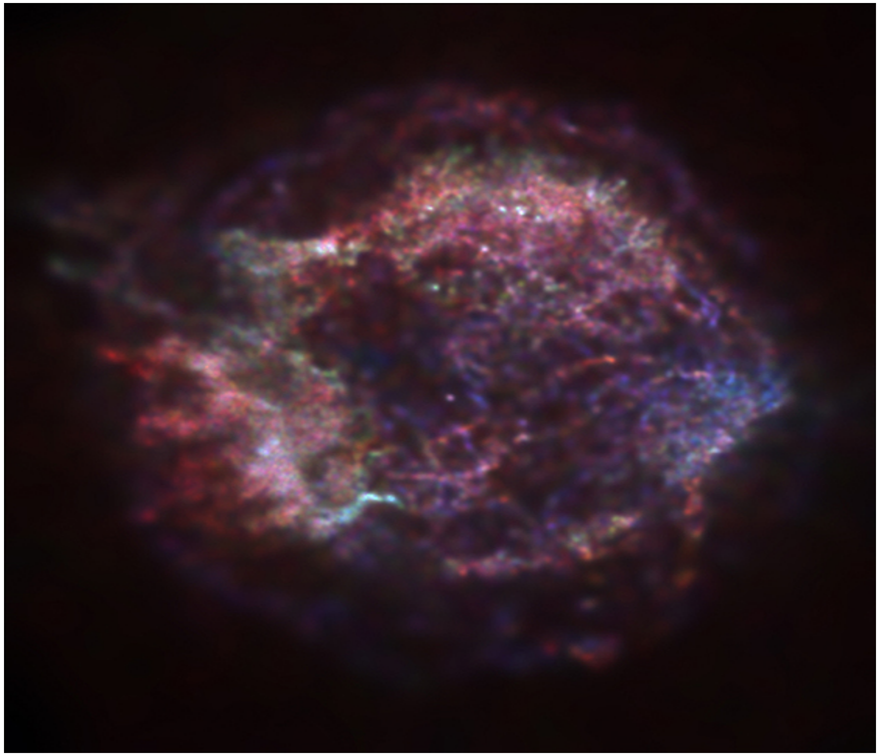
25 solar mass supernova, 1.2×10^{51} erg explosion



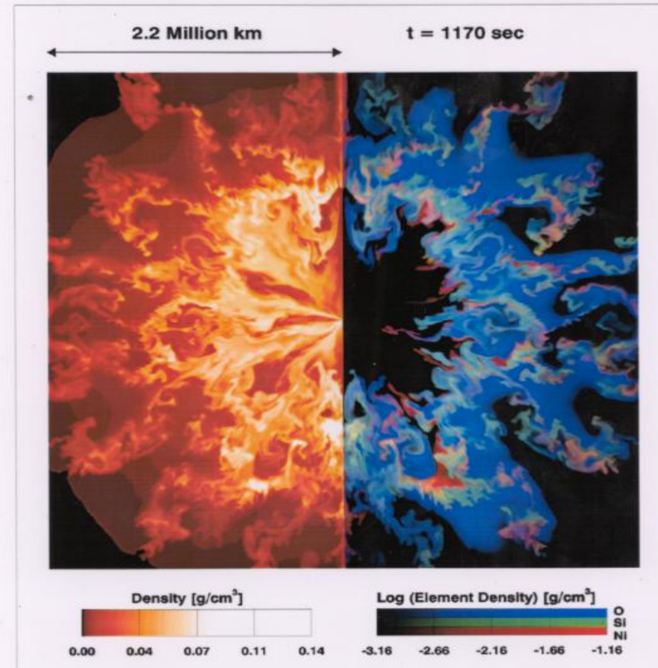
Calculation using modified FLASH code – Zingale & Woosley


Log10 Density (g/cm³)
3.0×10^12
2.5×10^12
2.0×10^12
1.5×10^12
1.0×10^12
5.0×10^11
0
y (cm)
8.00×10^12
8.50×10^12
9.00×10^12
9.50×10^12
1.00×10^13
1.05×10^13
x (cm)
-5.0
-5.5
-6.0

Diagnosing an explosion



Kifonidis et al. (2001), *ApJL*, *531*, 123

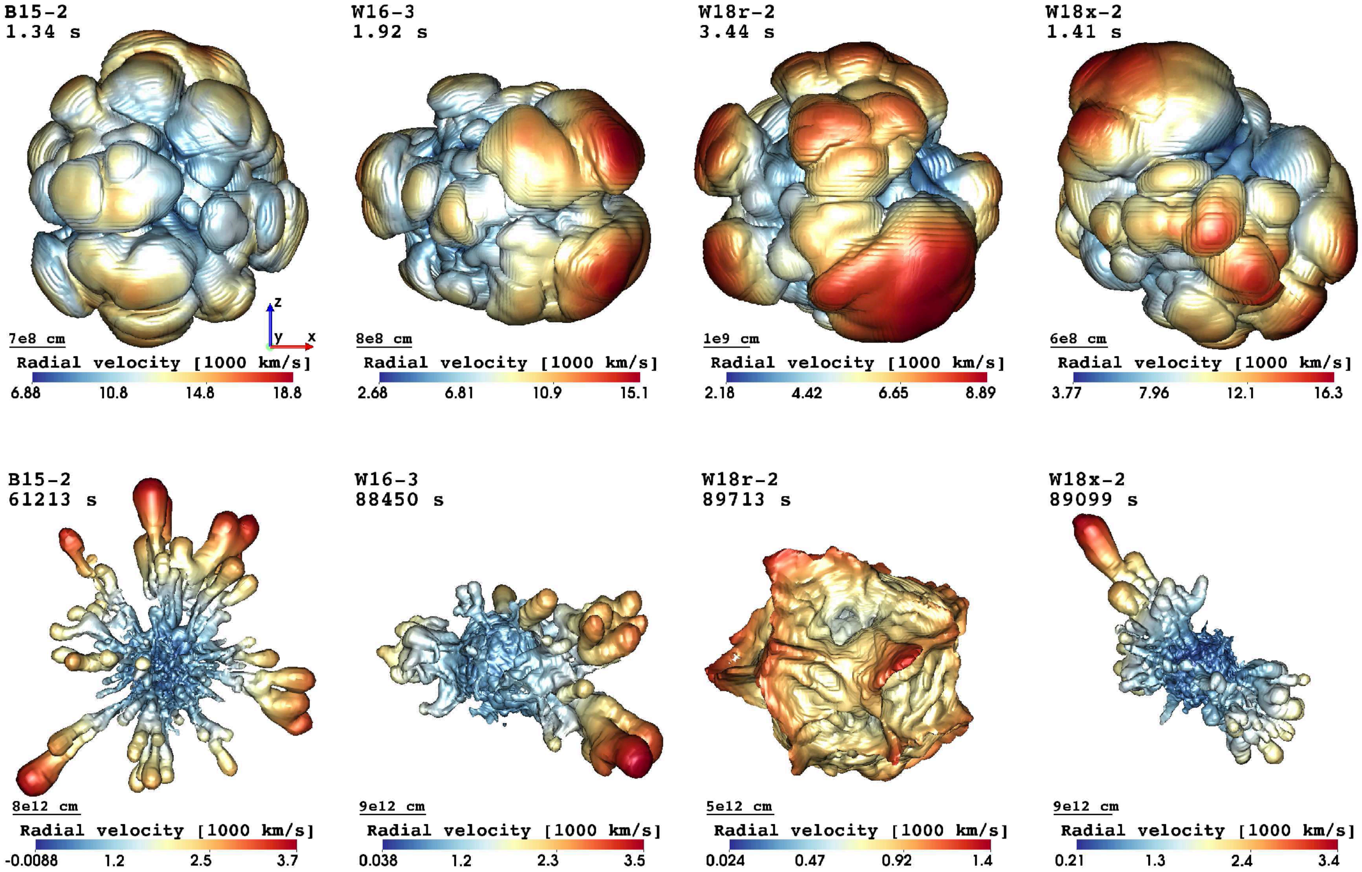


Left - Cas-A SNR as seen by the Chandra Observatory Aug. 19, 1999

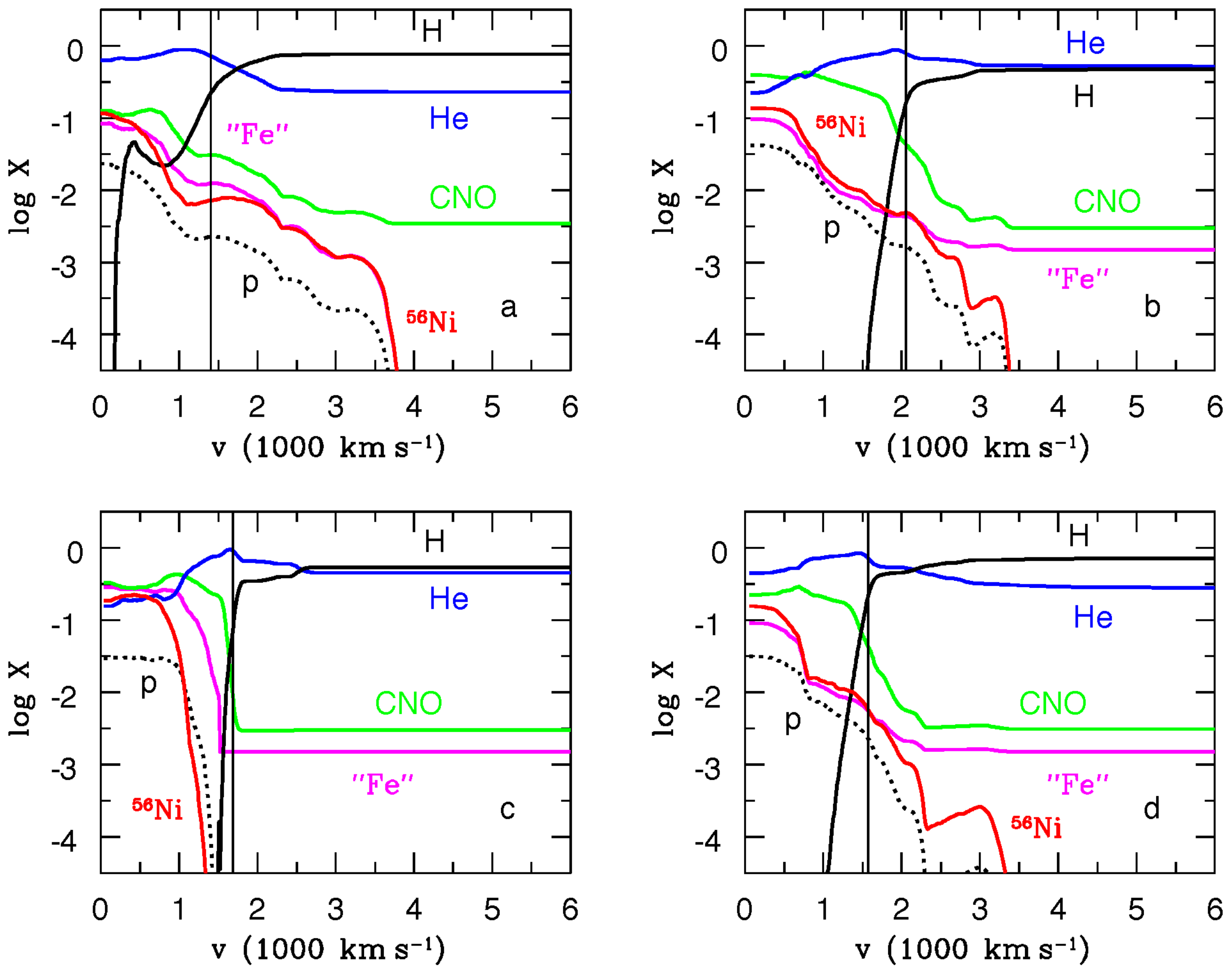
The red material on the left outer edge is enriched in iron. The greenish-white region is enriched in silicon. Why are elements made in the middle on the outside?

Right - 2D simulation of explosion and mixing in a massive star - Kifonidis et al, Max Planck Institut fuer Astrophysik

Mixing in SN 1987A – Utrobin et al (2019)



^{56}Ni-rich material at two times in 4 3D models


H
He
"Fe"
CNO
p
^{56}Ni
a
b
c
d
log X
v (1000 km s^{-1})