Nuclear Astrophysics
Lecture 1

Gail McLaughlin

North Carolina State University

Neutrino Astrophysics

- What are the fundamental properties of neutrinos?
- What do they do in astrophysical environments?
- What do neutrinos in a core collapse supernova do?
- What do neutrinos in a black hole accretion disk do?

Nuclear Astrophysics

- What is the origin of the elements?
- Where are the heaviest elements made?
- What elements are made in compact object mergers/ stellar explosions?

Outline

- Lecture 1: Introduction
- Lecture 2: Nucleosynthesis
- Lecture 3: The r-process

Gas clouds that make stars

As a gas cloud is made primarily of hydrogen and helium, with trace amounts of other elements.

The small amounts of other elements were made in the Big Bang (Li, Be, B) or else were ejecta in prior stellar explosions (carbon and heavier).

Stars are formed when a gas cloud collapses. Very old stars have mostly Big Bang elements, whereas newer stars have more heavier elements (although hydrogen and helium still dominate).

Stars

As the cloud collapses, the density and temperature increase. If the cloud is large enough, then the nuclear reactions begin. Hydrogen starts a series of reactions, the pp chain, the burn hydrogen to helium. Hydrogen burning occurs at around 10^7K.

$$p + p \rightarrow D + e^+ + \nu_e$$

$$D + p \rightarrow {}^3He + \gamma \qquad \text{(pp I, 86\%)}$$

$$^3He +^3 He \rightarrow {}^4He + p + p$$

Stars: what do we see?

Light is the primarily and traditional tool of astronomy, and we use it to look at stars.

Light has many excellent features

- it doesn't change identity
- it travels in a straight line (w/o GR)
- it is easy to detect
- it contains spectral information

Light has one less excellent feature:

- it is not very penetrating

Even very dilute matter is opaque to light, and this prevents us from seeing inside stars. It takes 10,000 years for a photon to random walk its way out of the center of the sun.

Stars and Neutrinos

In order to use light to test for nuclear reactions in the sun, we need a model e.g. John Bahcall's standard solar model. There is a lot of physics in between the nuclear reactions and the surface emission of the sun.

Another tool is the neutrino:

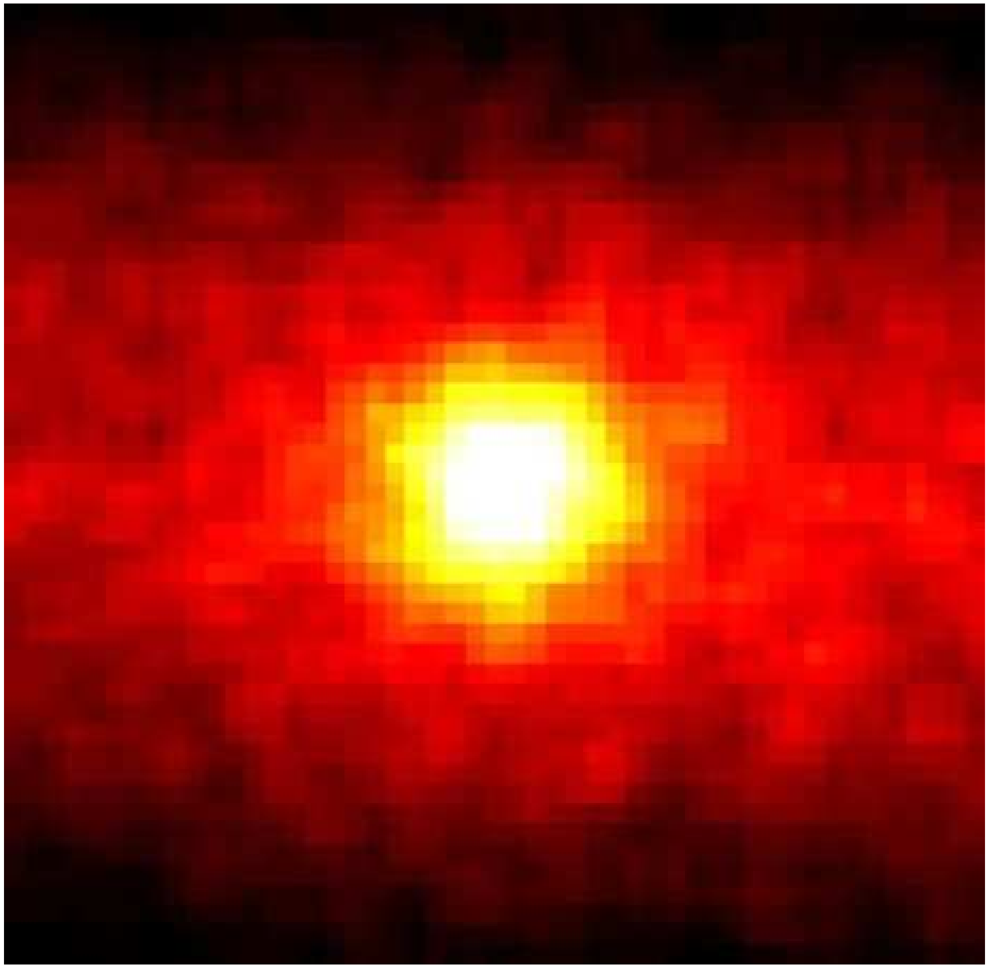
The neutrino has many excellent features:

- it is very penetrating
- it travels in a straight line (w/o GR)

Neutrinos have a few less excellent features:

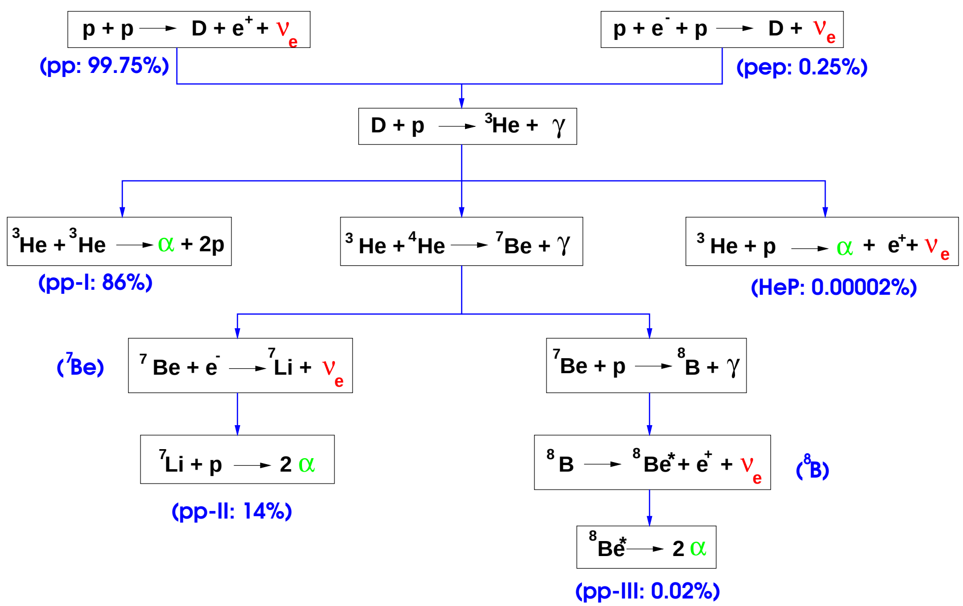
- they change identity
- they are hard to detect

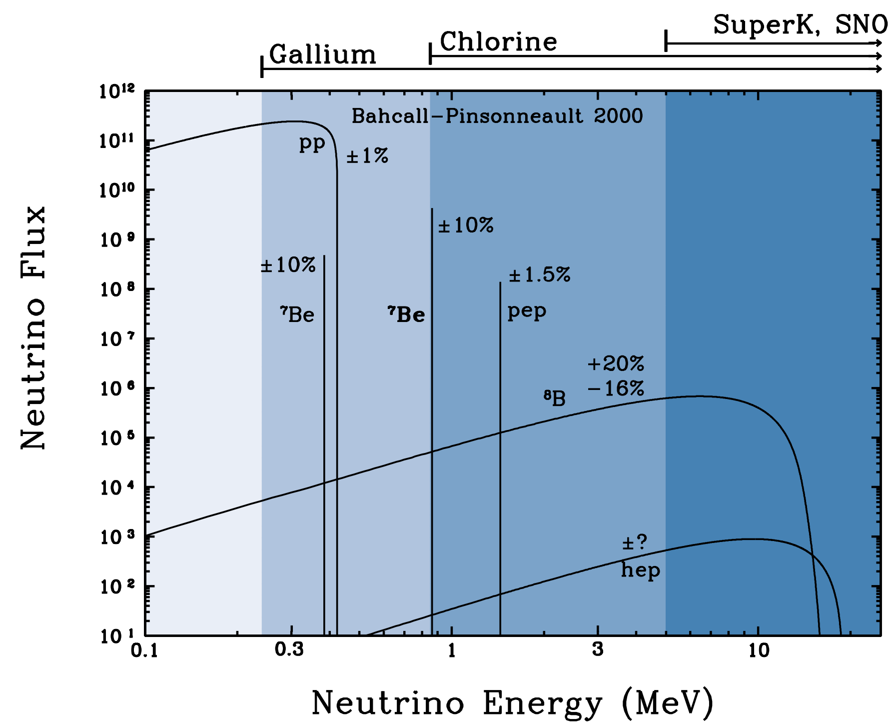
Neutrinos from the sun that are detected here on earth have traveled straight from the nuclear reaction in which they were produced and were emitted just a few minutes before they were detected.



Reconstruction of the sun from neutrinos, SuperK collaboration

Detecting neutrinos from the sun

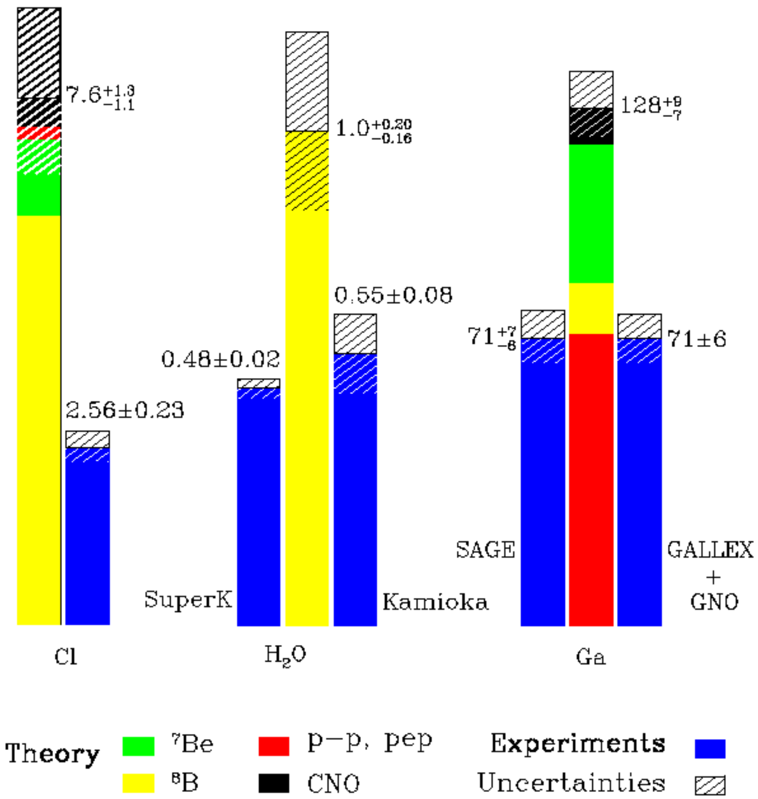




Detectors like SuperKamiokande and SNO have/had higher thresholds and see primarily Boron-8 neutrinos. To see other neutrinos, a detection reaction *and* a detector with a low threshold is required, e.g. Borexino.

Experiments measure deficit of ν_e

(Homestake, Gallex, Sage, Kamiokande, SuperK, SNO)



Homestake

$$\nu_e + Cl \rightarrow e^- + Ar$$

SAGE, GALLEX

$$\nu_e + Ga \rightarrow e^- + Ge$$

SuperK

$$\nu_e + e^- \rightarrow \nu_e + e^-$$

Neutrino Oscillations

Neutrinos come in different flavors $\nu_e, \nu_\mu, \nu_\tau, \ldots$

Flavor eigenstates $\neq$ mass eigenstates   Pontecorvo (1957)

The rest is just quantum mechanics ...

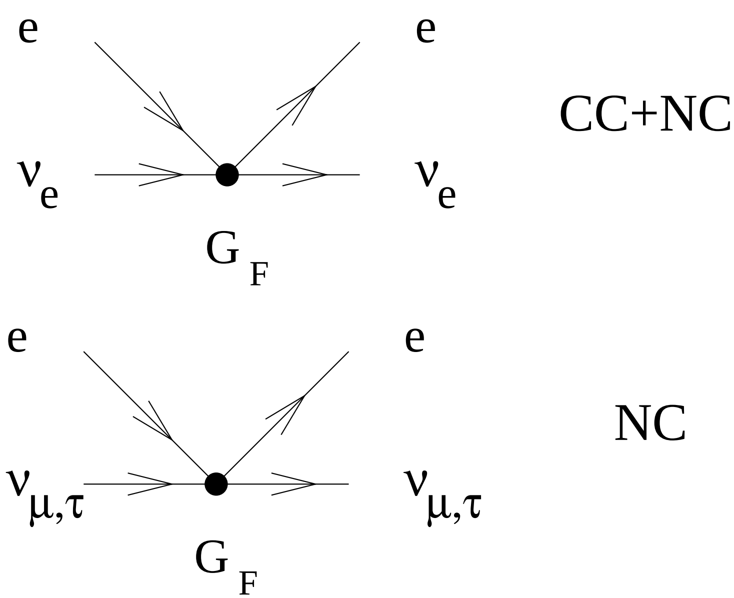
Consider two flavors ν_e, ν_μ

$$\begin{aligned} |\nu_e\rangle &= \cos\theta|\nu_1\rangle + \sin\theta|\nu_2\rangle \\ |\nu_\mu\rangle &= -\sin\theta|\nu_1\rangle + \cos\theta|\nu_2\rangle \end{aligned}$$

Flavor eigenstates $|\nu_{e,\mu}\rangle \quad \leftrightarrow \quad$ Mass eigenstates $|\nu_{1,2}\rangle$

Matter Enhanced (MSW) Oscillations

Neutrino propagation in matter: forward scattering on electrons leads to effective potential



$$V = \frac{V_e - V_x}{2} = 2\sqrt{2} G_F N_e(r)$$

electron density $N_e(r)$

Wolfenstein (1978)

Mikheyev-Smirnov (1985)

Modified wave equation

$$i\hbar c \frac{d}{dr}\psi_\nu = \begin{pmatrix} V - \frac{\delta m^2}{4E}\cos(2\theta) & \frac{\delta m^2}{4E}\sin(2\theta) \\ \frac{\delta m^2}{4E}\sin(2\theta) & -V + \frac{\delta m^2}{4E}\cos(2\theta) \end{pmatrix} \psi_\nu$$

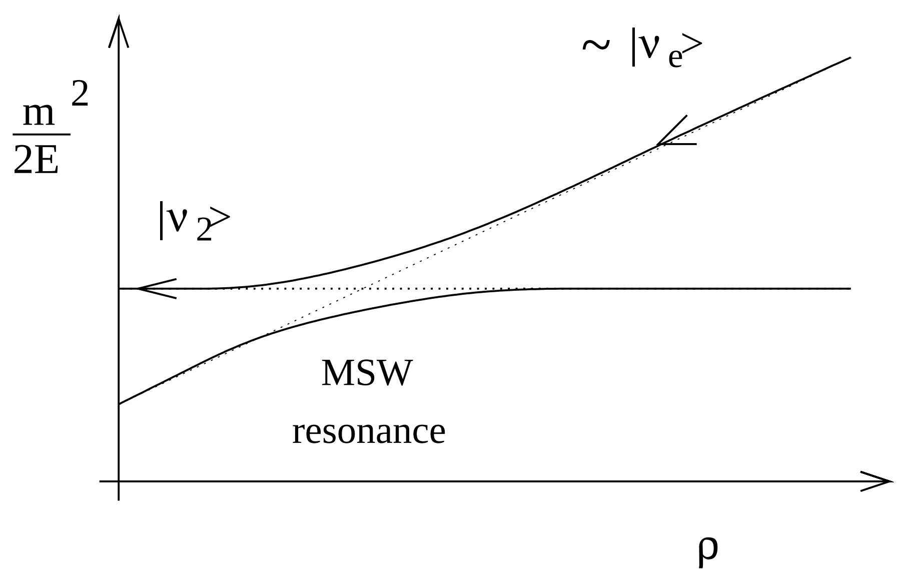
Consider eigenstates of RHS ("matter eigenstates")

Start at high density

matter eigenstates $\simeq$ flavor eigenstates

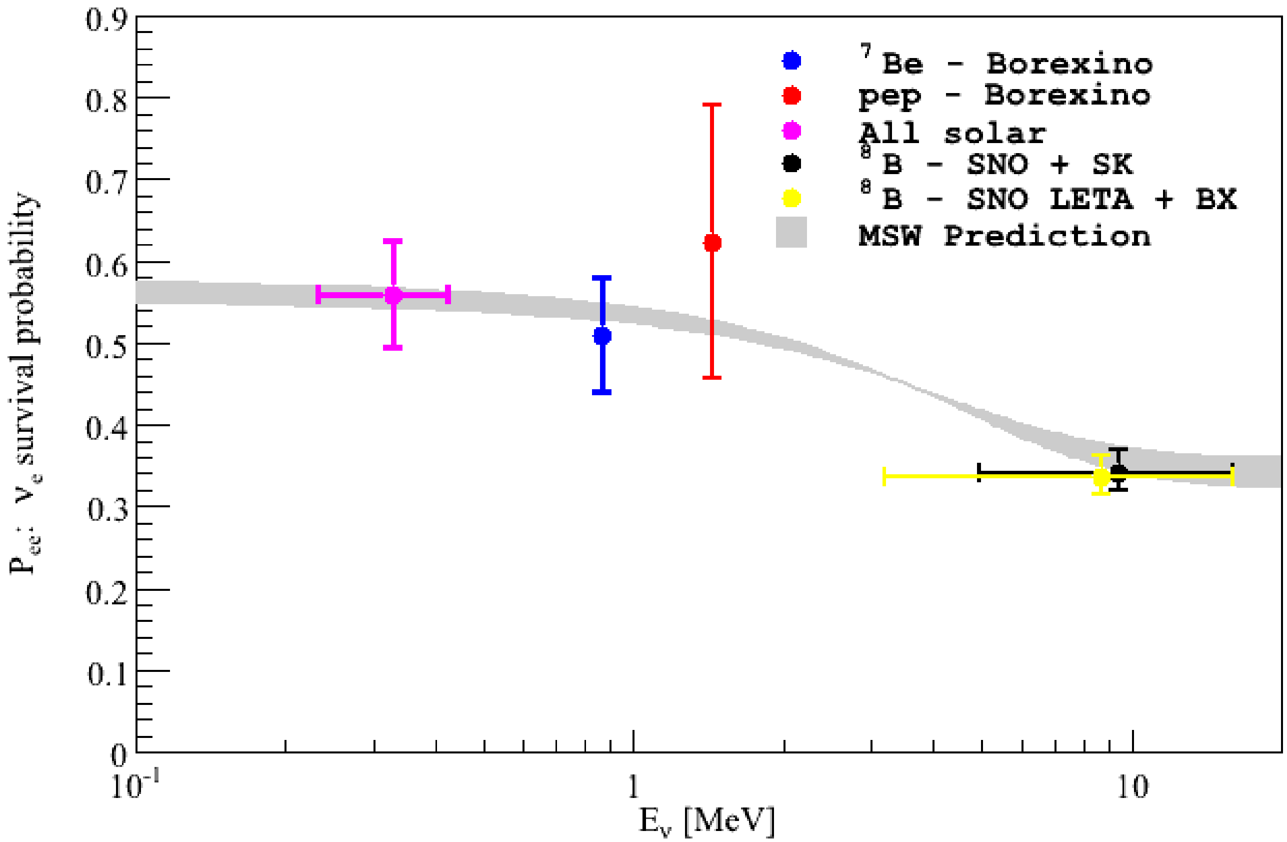
Resonance occurs if diagonal element vanishes. Possibilities

$$
\begin{array}{llll}
1) & \nu_e \rightarrow \nu_1 & \mathrm{non-adiabatic} & P(\nu_e) \sim \cos^2\theta, \quad P(\nu_\mu) \sim \sin^2\theta \\
2) & \nu_e \rightarrow \nu_2 & \mathrm{adiabatic} & P(\nu_e) \sim \sin^2\theta, \quad P(\nu_\mu) \sim \cos^2\theta
\end{array}
$$



Boron-8 neutrinos undergo MSW oscillations

pp neutrinos undergo vacuum oscillations



Helium burning

When the core of our sun runs out of hydrogen, it will start to burn helium. It contracts until the temperature reaches around 10^8K.

$$\alpha + \alpha + \alpha \rightarrow \; ^{12}C$$

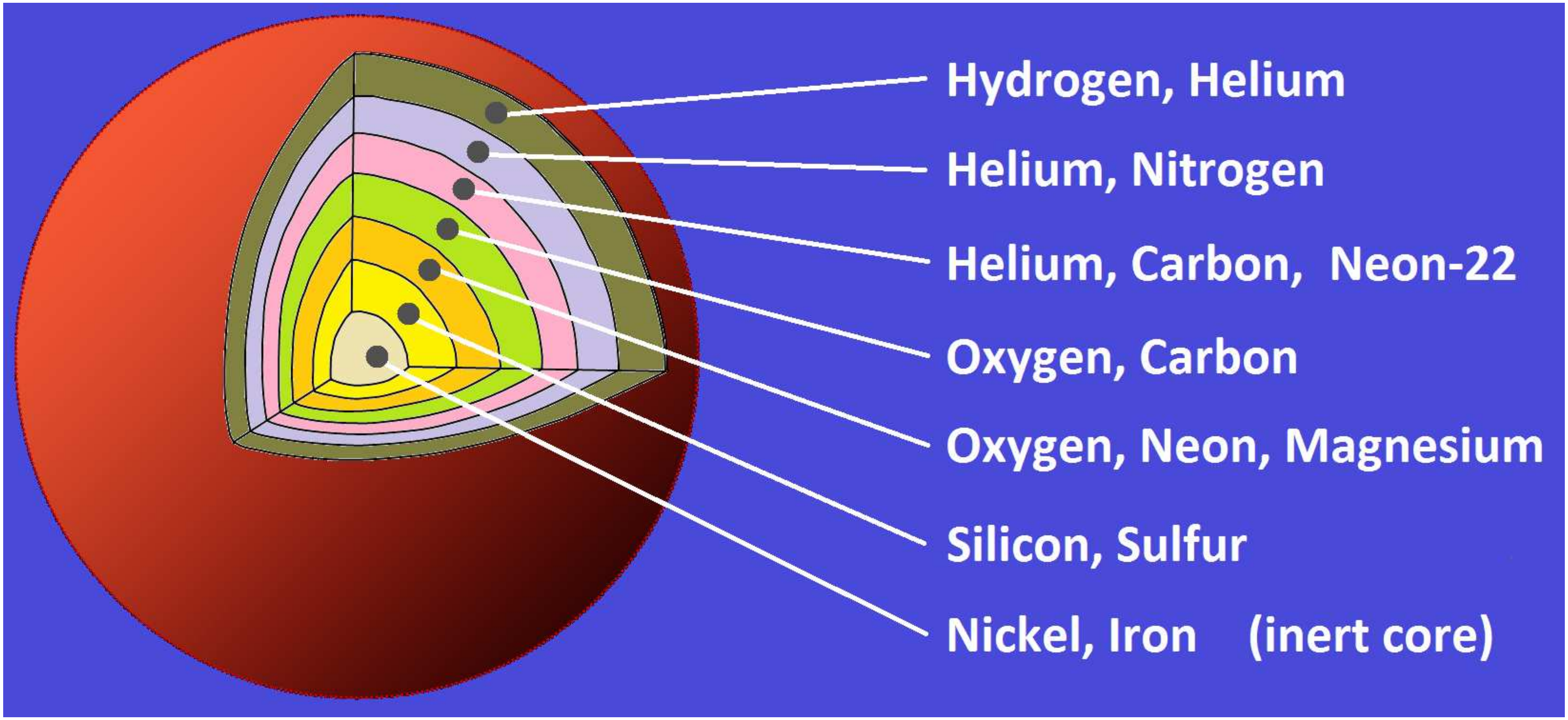
$$\alpha +^{12} C \rightarrow \; ^{16}C$$

Unfortunately this happens when the sun becomes a red giant and incinerates the earth. Luckily that is five billion years away.

Our sun is will not become hot enough to burn past carbon and will end its like as a carbon-oxygen white dwarf.

Burning past carbon and oxygen

A bigger star can get hotter and burn past oxygen. It develops either an oxygen-neon magnesium core (~ 8 solar mass star) or an iron core (10 solar masses). The later burning stages occur at around 10^9K.



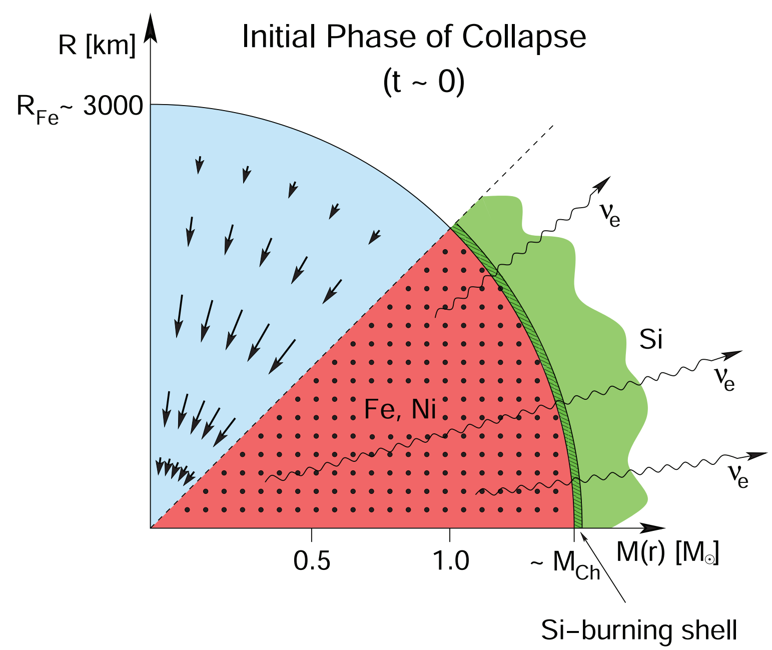
Iron core collapse

Iron is tightly bound and no more energy can be released as it burns past iron. The iron core becomes larger and larger and is supported by electron degeneracy pressure. At some point it can no longer maintain hydrostatic equilibrium and it begins to collapse.

Specifically, the iron starts to dissociate and electrons capture on the protons.

$$e^- + p \rightarrow n + \nu_e$$

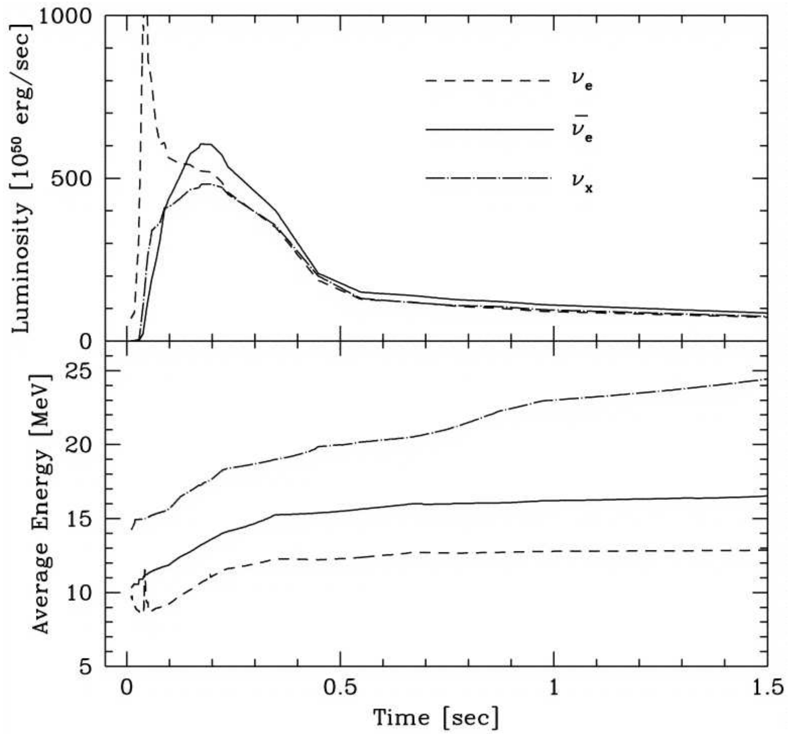
Once initiated, the process runs away due to energy losses.



Neutronization burst

The neutrinos produced in this way during collapse are called the "neutronization burst".

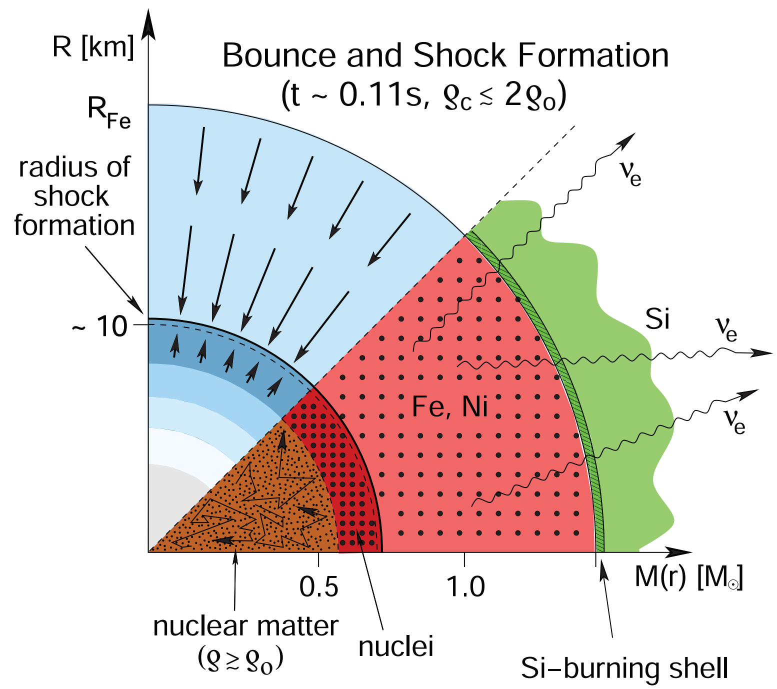
$$e^- + p \to n + \nu_e$$



The core rebounds and creates a shock

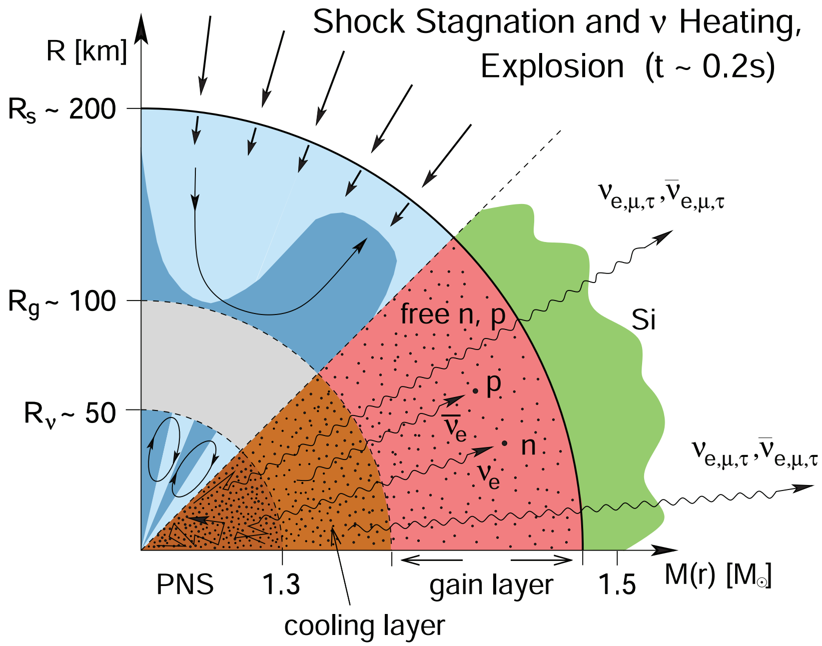
The core collapses to very high density and then rebounds. The details of the bounce depend on the equation of state.

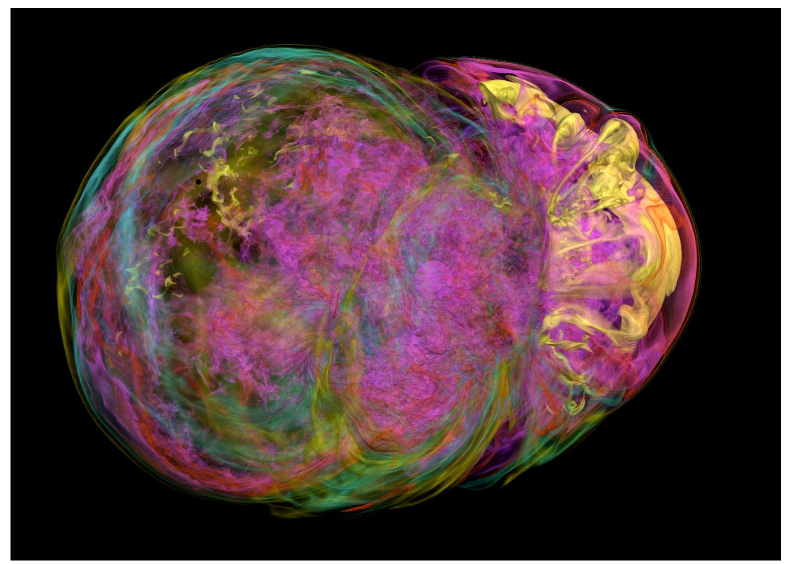
Material moving outward collides with material moving inward creating a shock.



The shock stalls

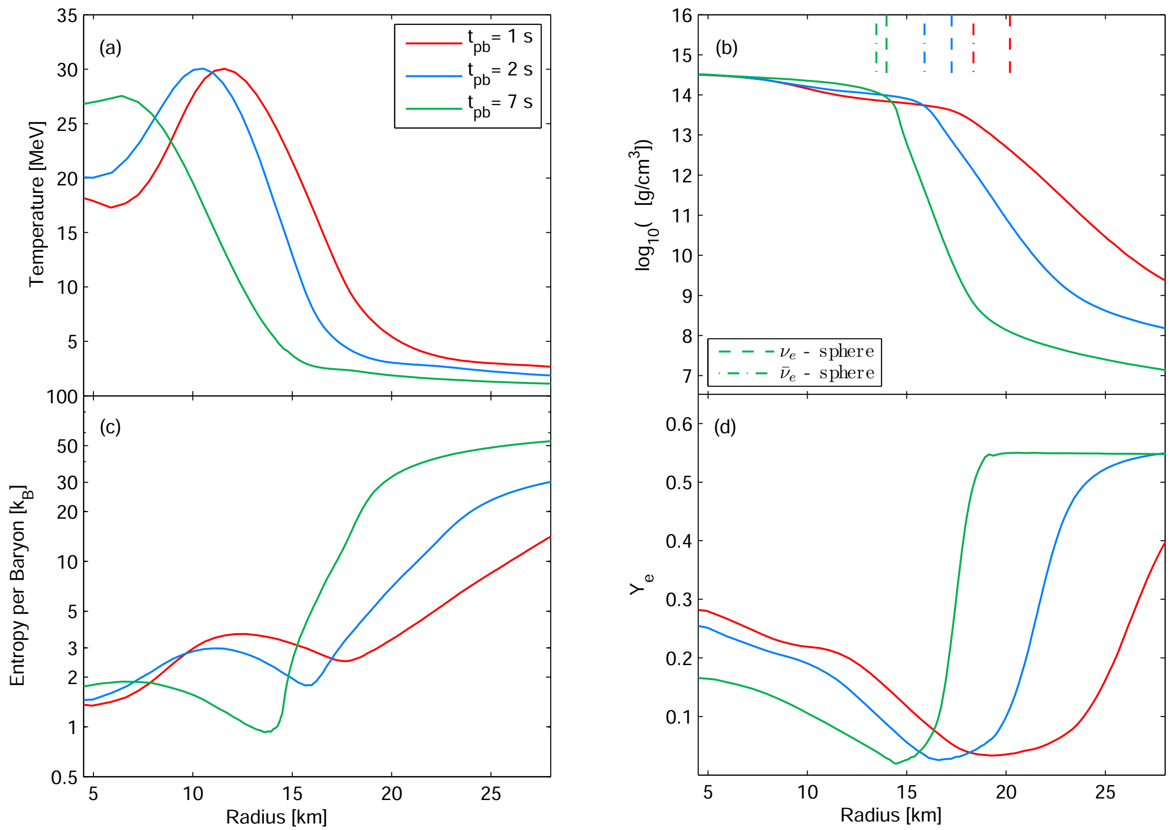
The shock uses up some of its energy photodissociating nuclei along its path. Then it stalls and becomes a standing accretion shock (SAS).





A protoneutron star is at the center

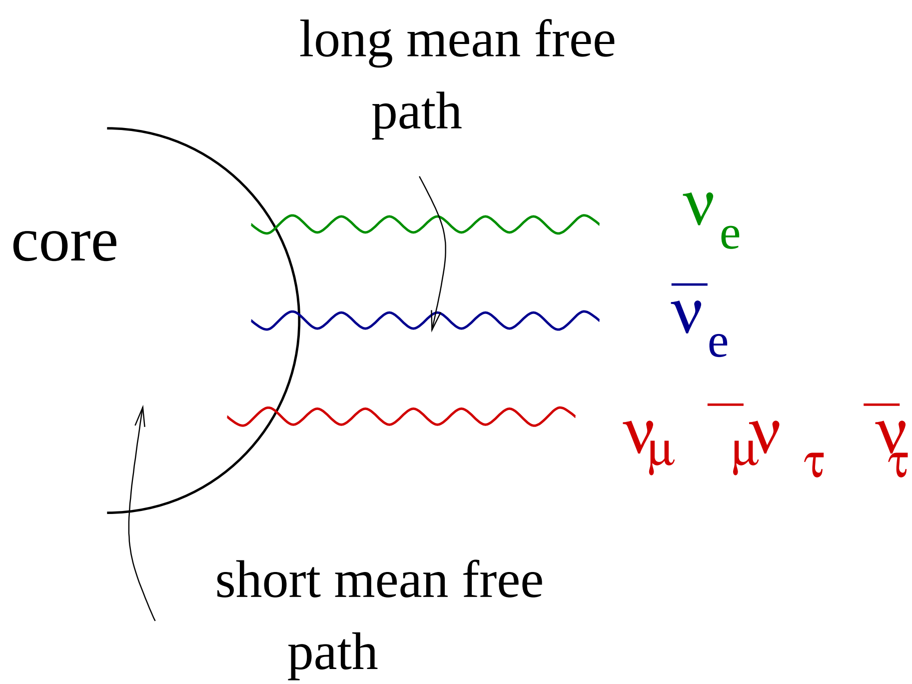
The core is now has a radius of about 30-40km, and is called a proto-neutron star. The very center of the protoneutron star has temperatures of about a several tens of MeV, $\sim 10^{12}$K, and densities of about 10^{14} g/cm^2. The exact composition is not known.



The proto-neutron star releases thermal neutrinos

The proto-neutron star is so hot that everything is in thermal equilibrium, which means that thermal processes, e.g. neutrino bremsstahlung, produce neutrinos of all flavors. Neutrinos come in three flavors ν_e, ν_μ and ν_τ.

These neutrinos scatter out of the proto-neutron star and then free stream.



Neutrino diffusion: order of magnitude estimate

How long does it take a 30 MeV neutrino to scatter out of a proto neutron star core?

1. 10 milliseconds

2. 10 seconds

3. 10 minutes

4. 10 days

5. 10 years

some useful quantities:
core size, R = 40 km; density, $\rho = 10^{14} \mathrm{g\,cm}^{-3}$
weak cross section $\sigma = \sigma_0 \left(\frac{E_\nu}{m_e c^2}\right)^2$, $\sigma_0 = \frac{4G_f^2 m_e^2}{\hbar^4} = 1.76 \times 10^{-44}\,\mathrm{cm}^2$
diffusion time $t_{\mathrm{diff}} = \frac{R^2}{l_\nu c}$, where l_ν is the mean free path

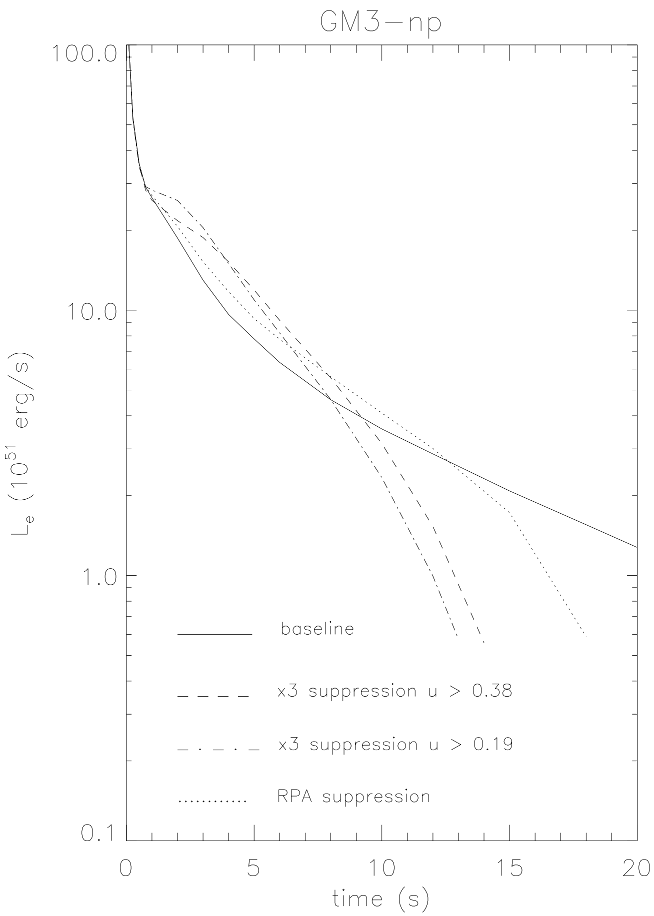
Neutrino Scattering

Neutrinos scatter from nucleons, nuclei:

- Scattering on neutrons, protons $\nu(\bar{\nu}) + n \to n + \nu(\bar{\nu})$
- electron (neutrino) capture $e^- + p \leftrightarrow n + \nu_e$
- positron (antineutrino) capture $e^+ + n \leftrightarrow p + \bar{\nu}_e$
- ν scattering on nuclei $\nu + A(Z,N) \leftrightarrow A(Z,N) + \nu$
- ν_e capture on nuclei $\nu_e + A(Z,N) \leftrightarrow A(Z+1,N-1) + e^-$
- $\bar{\nu}_e$ capture on nuclei $\bar{\nu}_e + A(Z,N) \leftrightarrow A(Z-1,N+1) + e^+$

Neutrino diffusion timescale

The neutrino diffusion timescale depends on composition of the core. Here are a few predictions made from different equations of state.



The timescale on which neutrinos are emitted provides some information about the composition of the core.

Leptons above the protoneutron star

Above the protoneutron star the neutrinos are free streaming - almost - a small amount of scattering still goes on. It is a reasonably good approximation to assume that the temperatures are frozen in at the decoupling surface, and the flux declines roughly as $1/r^2$.

Above the proto-neutron star electrons and baryons are still thermally coupled and the temperature declines as you move away from the neutrino surface. The flux is determined by thermal distributions.

Neutrino Decoupling

Inside the core neutrinos scatter, and outside they free stream. The transition is called the decoupling surface.

Which neutrinos have the lowest temperature at decoupling?

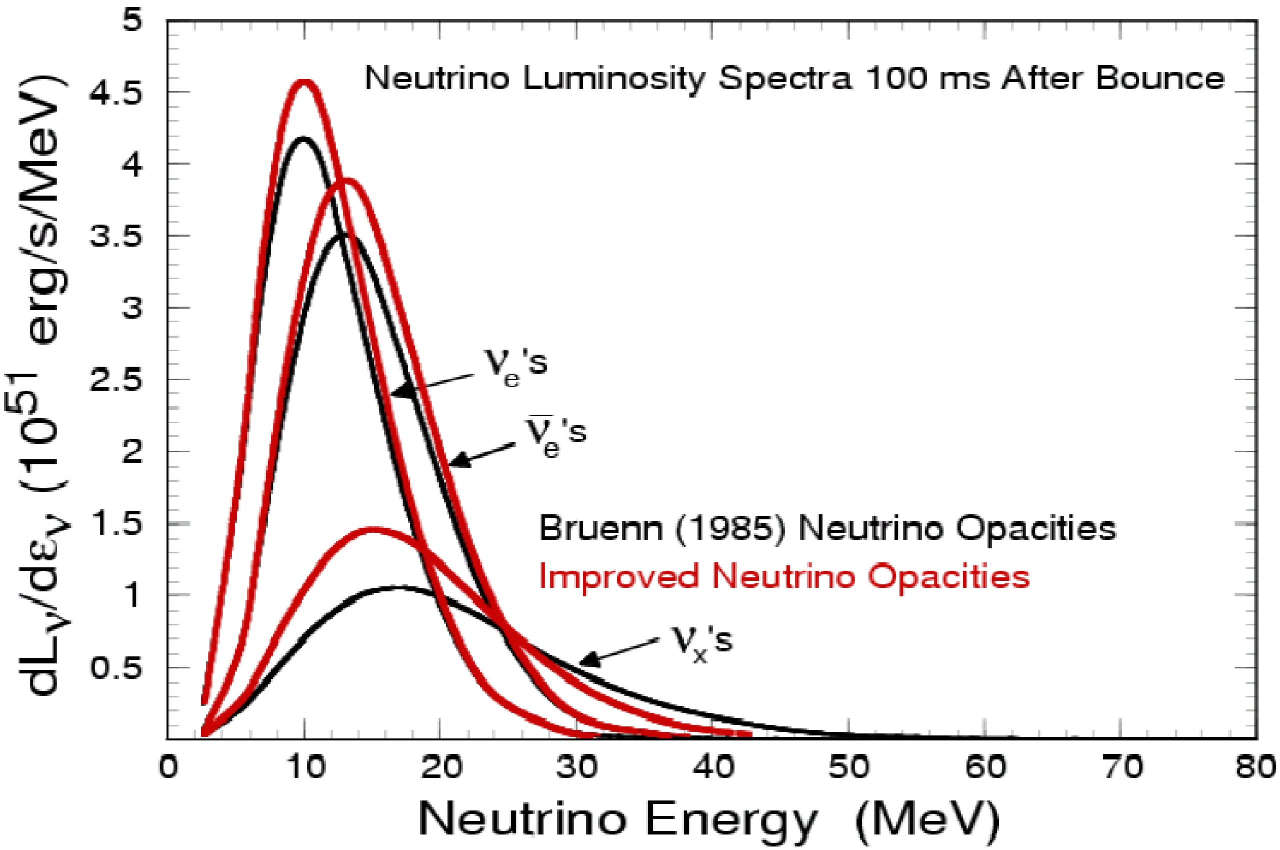
1. ν_e
2. $\bar{\nu}_e$
3. ν_μ
4. $\bar{\nu}_\mu$

some useful information: a proto-neutron star has more neutrons than protons

the mass of an electron is 0.511 MeV
the mass of a muon is 105.7 MeV

Neutrino spectra

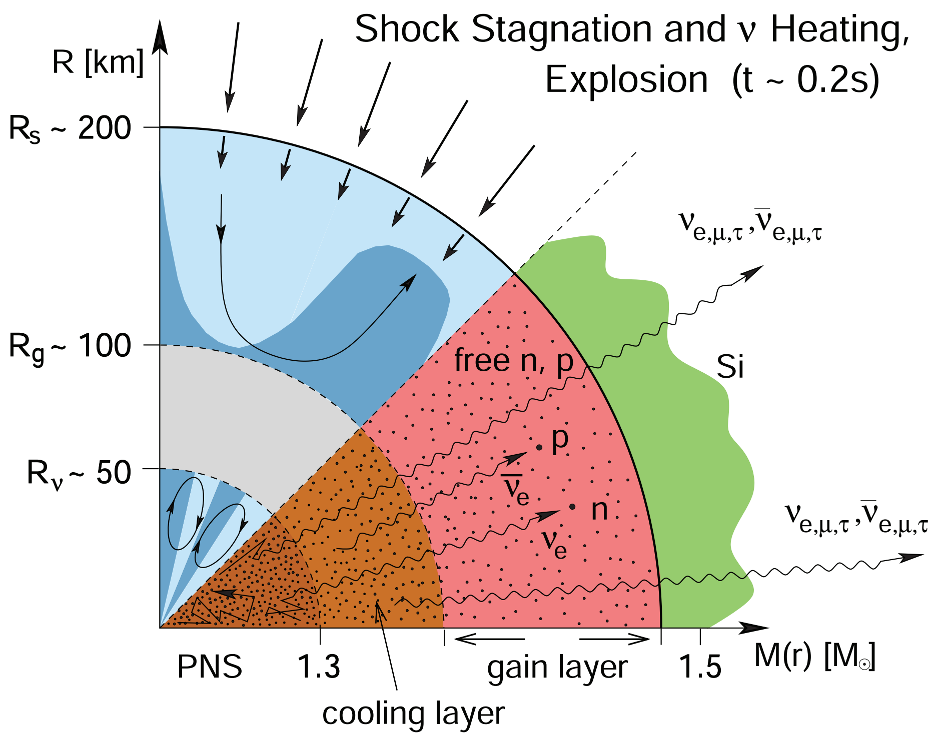
Spectra are determined by the physics at the surface of the proto-neutron star.



There is some uncertainty in the spectra. Different opacities predict different spectra.

Neutrino diffusion

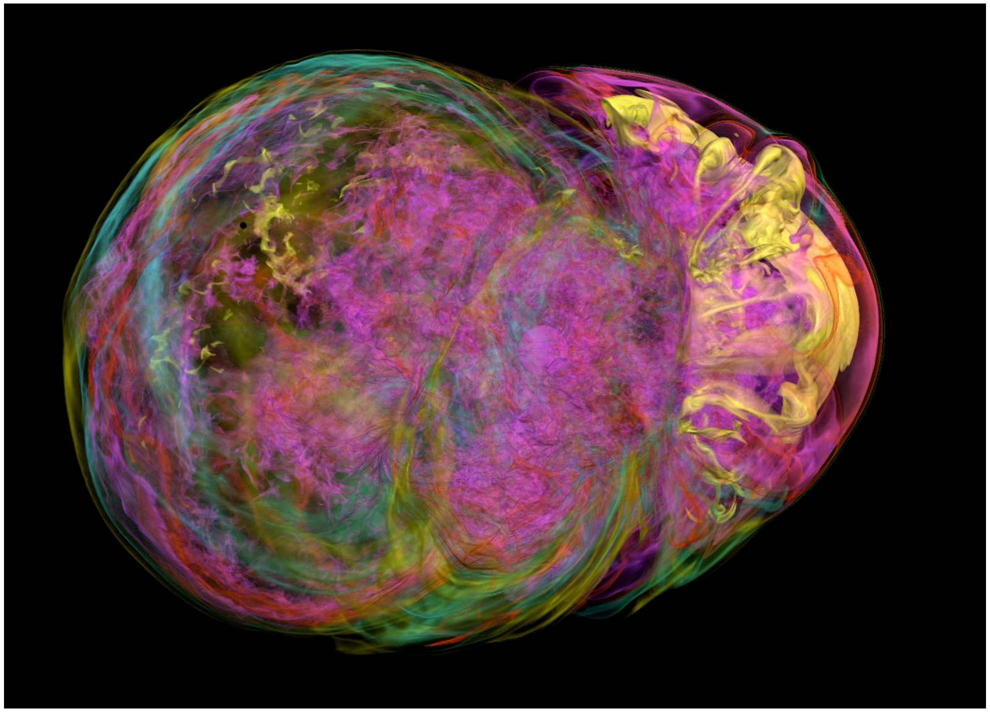
Neutrinos first cool and then heat the material above the material above the protoneutron star.



Heating occurs when small numbers of neutrinos deposit energy in the region above the gain radius at a higher rate than the material produces neutrinos.

Re-energizing the stalled shock

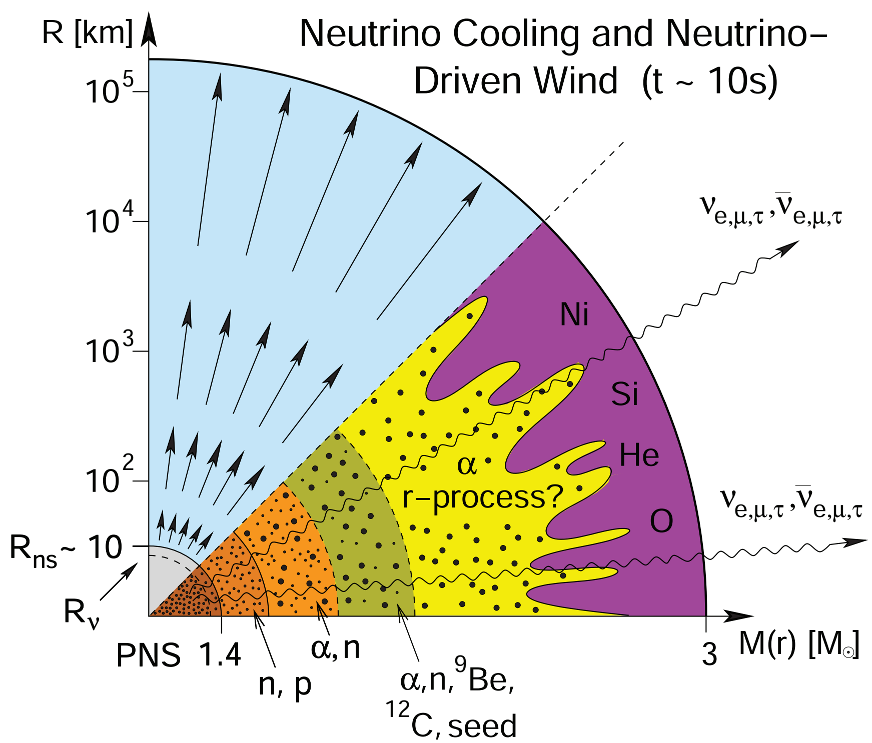
Neutrinos heat the material below the stalled shock, helping along the standing accretion shock instability.



Explosion!

Details of the explosion are still under investigation. But when the explosion occurs, explosive nucleosynthesis occurs as well.

The energy of the shocked material allows nuclear burning, and a number of elements are produced. Lots of nickel, but smaller amounts of other elements as well, like titanium.



Why nickel?

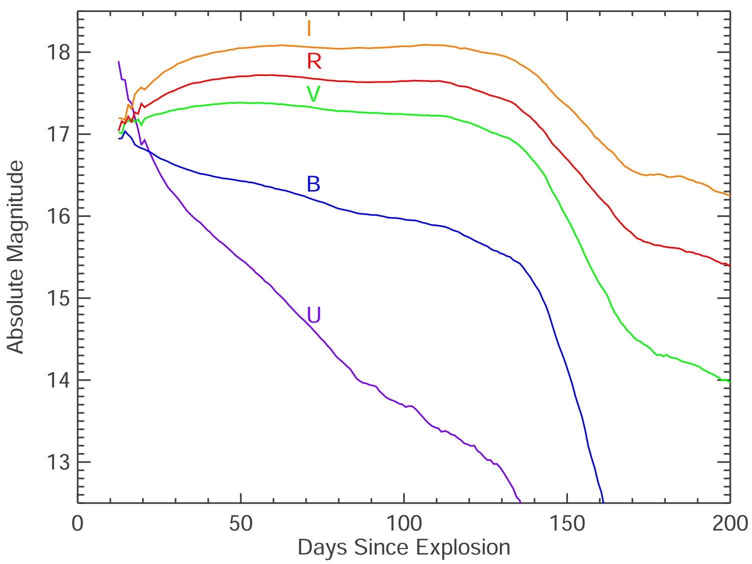
Nickel is a tightly bound iron group element that has equal numbers of neutrons and protons. Thus, nuclear statistical equilibrium favors nickel.

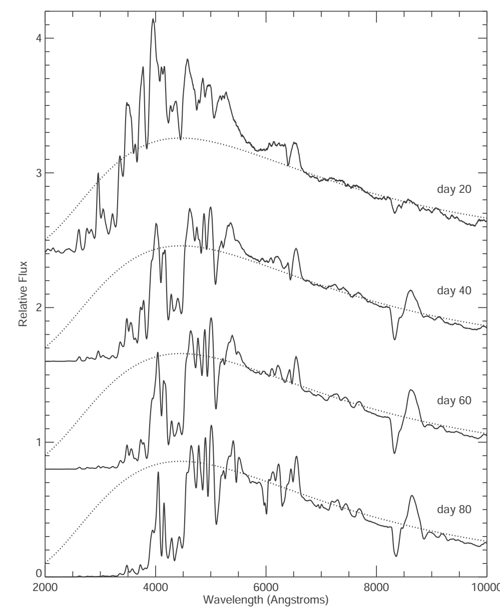
At high temperatures, e.g. T $=$ a few $\times 10^{10}$ K, the distribution of elements is determined by their binding energies, not nuclear reactions. i.e. all nuclear reactions are balanced in nuclear statistical equilibrium.

Nickel-56 is important for the explosion

Lots of nickel is produced in explosive burning and this is reflected in the light curve.

The 56Ni beta decays to 56Co which decays to 56Fe. Gamma rays are produced at each step that thermalize. The timescale for these decays is months, which is reflected in the light curve.





Kasem and Woosley 2012

Other nucleosynthetic products

As the shock moves out, it loses energy and the temperature drops. Nuclear burning will still occur, just not in nuclear statistical equilibrium. Nuclear burning also occurs in the neutrino driven wind.

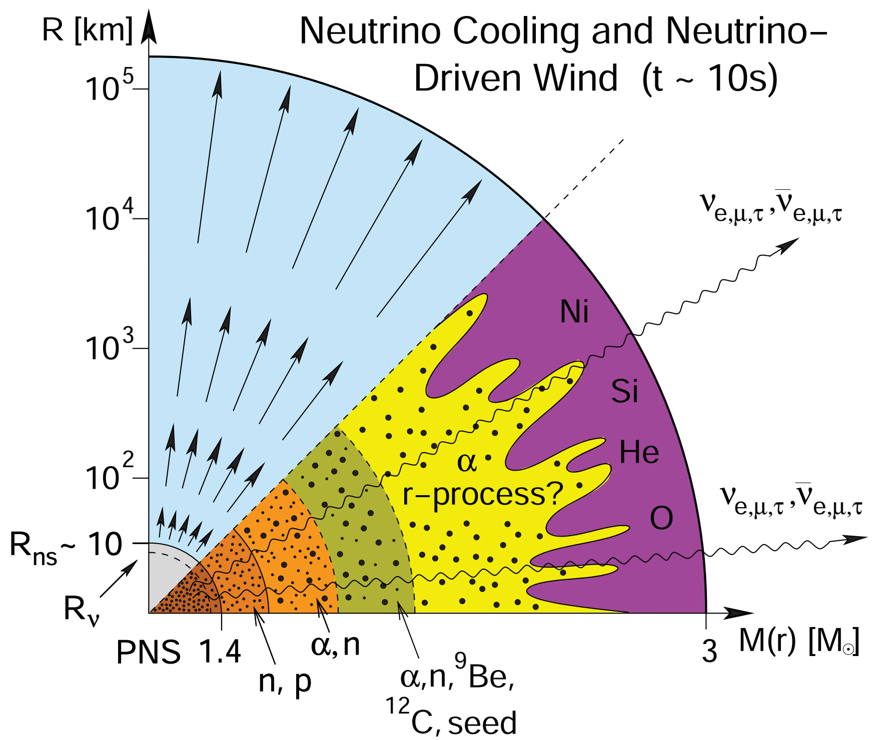
Ways to detect nucleosynthetic products

- supernova remnants
- spectroscopy of gas clouds, stars
- meteorites

Only elements that are made in large quantities can be seen in remnants

Neutrino driven wind

After the shock moves out, neutrinos deposit small amounts of energy near the surface of the proto-neutron star.



Nucleosynthesis in hot outflows

What matters?

- outflow timescale, milliseconds to seconds
- entropy $s \sim 20$ to $s \sim 400$
- electron fraction, $Y_e = \frac{\mathrm{p}}{\mathrm{n+p}}$, $Y_e \sim 0.1$ to 0.6,

Electron fraction is set by the weak interactions:

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

The weak interaction

The only way to convert protons to neutrons and vice-versa

- beta decay $n \rightarrow p + e^- + \bar{\nu}_e$
- electron (neutrino) capture $e^- + p \leftrightarrow n + \nu_e$
- positron (antineutrino) capture $e^+ + n \leftrightarrow p + \bar{\nu}_e$
- ν_e capture on nuclei $\nu_e + A(Z, N) \leftrightarrow A(Z+1, N-1) + e^-$
- $\bar{\nu}_e$ capture on nuclei $\bar{\nu}_e + A(Z, N) \leftrightarrow A(Z-1, N+1) + e^+$

p-process elements

e.g. $^{92}\mathrm{Mo}$ Z $=$ 42, N $=$ 50 $\rightarrow$ proton rich side of the valley of β stability

How do you make light p elements? Three possibilities:

- Just the right Y_e (around 0.485)
- The right flux of neutrinos on protons during element formation
- The right flux of neutrinos on nuclei during element formation

What are the candidate environments?

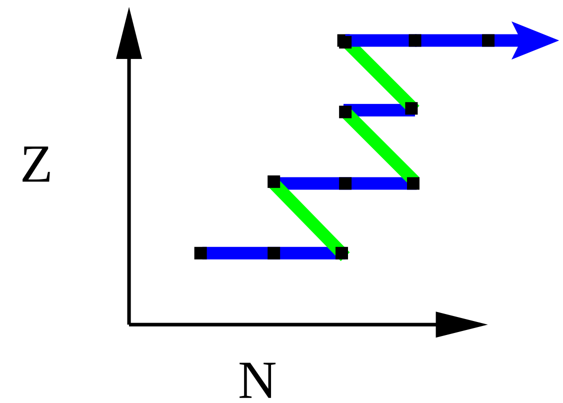
- Supernovae
- Accretion disks

The r-process elements

e. g. Uranium-238 Z=92, N=146 $\rightarrow$ need lots of neutrons

$$A(Z,N) + n \leftrightarrow A+1(Z,N+1) + \gamma$$

$$A(Z,N) \rightarrow A(Z+1,N-1) + e^- + \bar{\nu}_e$$

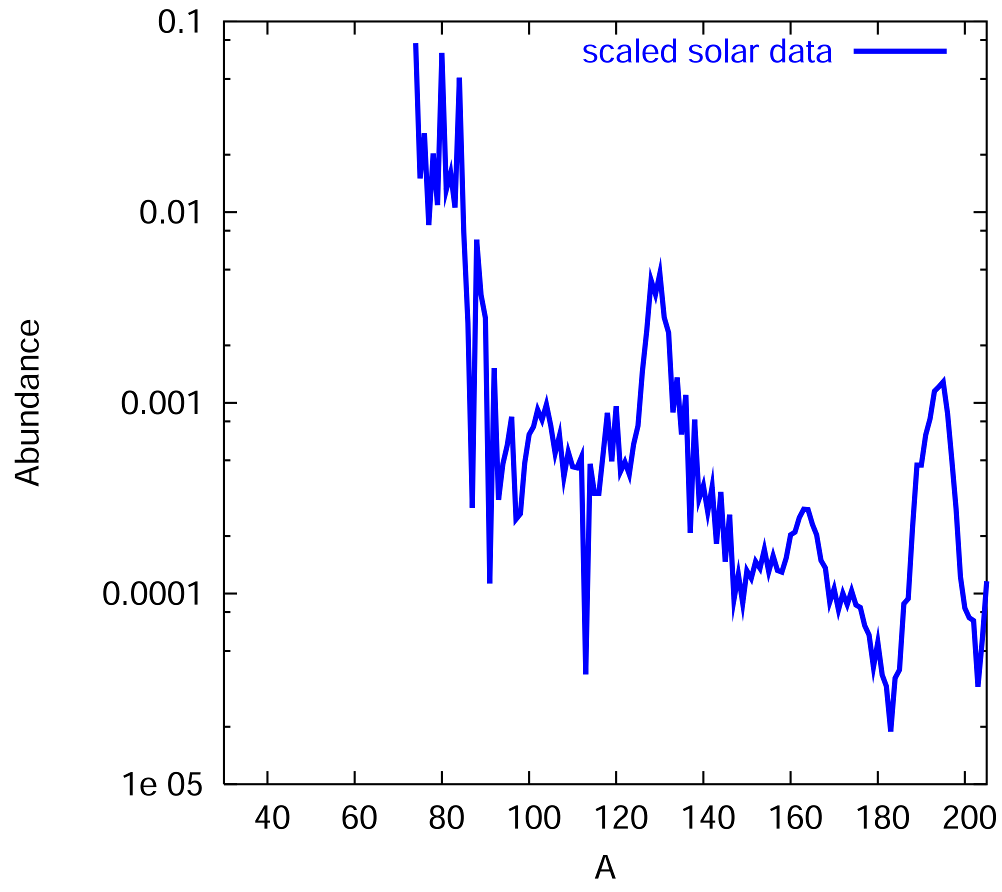


rapid neutron capture as compared with beta decay

Can the neutrino driven wind make the r-process?

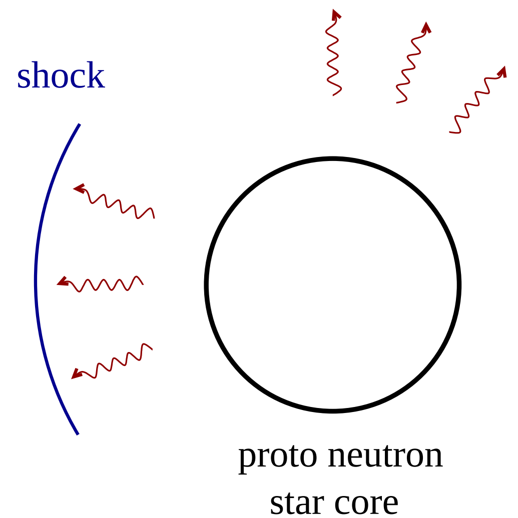
- only if there are enough neutrons in the outflow
- right neutrino/antineutrino ratios, entropy and timescale

Distribution of elements in the solar system

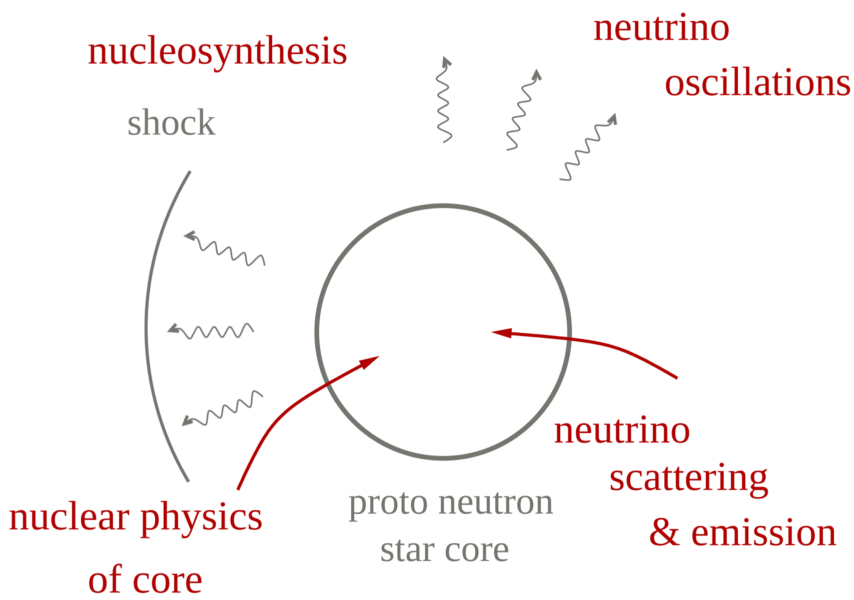


Galactic chemical evolution is an area of study where all types of nucleosynthetic events are studied for the frequency in the universe and matched to the final pattern.

Summary: cartoon of the supernova geometry



Summary: where is the nuclear physics?



Coming up next: Nucleosynthesis